Far-infrared and Sub-millimetre
Surveys of Circumstellar Discs

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Abstract

Stars of all ages and evolutionary stages are seen to be surrounded by discs of material. During the formation of a stellar system the stars are orbited by a massive protoplanetary disc composed of interstellar gas and dust, in which planet formation occurs. Between 1 and 10 Myr the protoplanetary disc disperses, leaving behind the newly formed system of planets and smaller bodies. The remaining material which has not formed into planets is referred to as a debris disc.

Even though the interstellar dust grains from the protoplanetary disc have long been removed from the system, debris discs can contain large quantities of dust due to collisions between larger bodies and cometary activity. Such dust can be detected by its thermal emission. This thesis focuses on observational studies at far-infrared and sub-millimetre wavelengths of debris discs and the late stages of protoplanetary disc evolution.

An overview of surveys for debris discs performed to date is presented, highlighting the limitations and statistical biases. The motivation, design and sample selection for two large surveys for debris discs around nearby stars, with the Herschel space observatory and the SCUBA-2 sub-millimetre camera on the James Clerk Maxwell Telescope, are described. The combination of a uniform observational strategy, longer wavelengths than previous surveys, and a large, clearly chosen sample – unbiased by stellar properties – will allow robust statistical conclusions of how the incidence and properties of debris discs depend on system parameters such as stellar mass, age, metallicity, binarity, and the presence of planets.

As a precursor to the Herschel and SCUBA-2 surveys, a volume-limited sample of 130 A type star systems was surveyed using observations at 24 and 70 μm from the Spitzer space telescope. Stellar photosphere fluxes at 24 and 70 μm, which were required to determine the presence of emission from dust, were predicted by fitting model flux distributions to optical and near-infrared photometry. Debris discs were detected around 46 systems, 12 of which – including the system with the largest dust mass – are new discoveries. This survey adds to the results of previous studies which show that debris disc incidence is not correlated with host star metallicity, despite the well known giant planet – metallicity correlation. This is in accordance with what is predicted from the core accretion theory of planet formation. The most significant result from this survey is that, contrary to results reported in a previous work, debris discs are overall less common around binary stars. Further investigation shows that systems with separations of $\sim 3-150$ AU are especially deficient of debris, whilst closer binaries and the primaries of wider binaries show debris detection rates consistent with those for single stars.

A sample of circumstellar discs around 29 young stellar systems with ages of 5–30 Myr were observed with the LABOCA sub-millimetre instrument on the APEX telescope at 870 μm, to provide disc masses or mass upper limits in support of a large Herschel programme. These targets included the η Chamaeleontis cluster and four bright Herbig Ae/Be stars which have not previously been observed at this wavelength. All but the Herbig Ae/Be stars were not detected, and $3\sigma$ dust mass upper limits of $\sim 0.1-3M_\odot$ are determined, with corresponding total disc masses of $\sim 0.03-1M_\odot$. These mass limits indicate that there is insufficient remaining material in these discs to form gas giant planets, and add to the prevailing view that protoplanetary discs typically disperse within 10 Myr and that gas giant planet formation must be completed before this time.

A search for cold dust emission from two of the Solar System’s nearest neighbours – α Centauri AB and ε Indi – was also performed with LABOCA. In both cases no debris disc emission was detected. A bright resolved feature was detected near α Centauri AB, however, follow-up observations at a second epoch, two years after the initial observations, showed that the feature is not co-moving with the stars. It is argued that the feature is most likely a pre-stellar core. The stars α Centauri A and B are detected, which is one of only very few detections of main sequence stellar photospheres at sub-millimetre wavelengths.
Declaration

I declare that this thesis is not substantially the same as any that I have submitted for a degree or diploma or other qualification at any other University. I further state that no part of my thesis has already been or is being concurrently submitted for any such degree, diploma or other qualification.

This thesis is the outcome of my own work except where specifically indicated in the text.

Neil Phillips
Edinburgh
22 November 2010
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# Table of Contents

List of Figures xi

List of Tables xiii

1 Introduction 1

1.1 The Solar System and its Debris Disc 4

1.1.1 Dust in the Solar System 6

1.1.2 Solar System Debris Evolution & the Late Heavy Bombardment 7

1.2 Classification and Evolution of Circumstellar Discs 9

1.2.1 Observational Classification and Evolution 9

1.2.2 Disc Evolution and Planet Formation Models 13

1.3 Young Stars 16

1.3.1 Pre-Main Sequence Stellar Evolution 16

1.3.2 Young Stellar Associations and Moving Groups 18

1.4 Thermal Emission 20

1.4.1 Black Body Radiation Properties 20

1.4.2 Emission from Real Objects 21

1.4.3 Luminosity and Effective Temperature 21

1.5 Dust Emission and Absorption 22

1.5.1 Grain Temperatures and Disc Radii 22

1.5.2 Disk Emission and Dust Mass 24

1.5.3 Effect of Optical Depth 25

1.5.4 Measuring Dust Masses 25

1.5.5 Fractional Luminosity 27

1.5.6 Dust Opacities 27

1.6 Dust Production, Removal and Dynamics 28

1.6.1 Radiation Pressure 28

1.6.2 Poynting-Robertson Drag 29

1.6.3 Collisions 31

1.6.4 Sublimation 32

1.6.5 Stellar Wind Pressure 33

1.6.6 Stellar Wind Drag 33

1.7 Thesis Outline 35

2 Debris Disc Surveys: SUNS and DEBRIS 37

2.1 Modelling Survey Detection limits and Selection Effects 39

2.2 Previous Surveys 41

2.2.1 IRAS 41

2.2.2 ISO 45

2.2.3 Ground Based (Sub-)Millimetre Observations 46

2.2.4 Spitzer 47

2.2.5 Science Results and the State of the Field 50

2.3 The SCUBA-2 Unbiased Nearby Stars (SUNS) survey 52
2.4 Surveys with Herschel ........................................ 54
  2.4.1 The DEBRIS survey ..................................... 55
2.5 Summary ....................................................... 61

3 Target selection for the SUNS and DEBRIS surveys ......... 63
  3.1 Introduction .................................................. 63
  3.2 Selection Criteria .......................................... 65
    3.2.1 Early Type Limit ..................................... 65
    3.2.2 Late Type Limit ....................................... 65
    3.2.3 Luminosity Class Limits ............................... 66
    3.2.4 Multiple Star Systems ................................. 67
    3.2.5 Subsample Sizes ...................................... 67
    3.2.6 DEBRIS Targets ...................................... 68
    3.2.7 SUNS Targets ......................................... 69
  3.3 Sources of Data ............................................ 70
    3.3.1 Parallaxes ............................................ 70
    3.3.2 Spectral Types ....................................... 70
    3.3.3 Photometry ........................................... 71
    3.3.4 Astrometry .......................................... 73
  3.4 Components of Multiple Systems .......................... 74
  3.5 Sample Properties ......................................... 75
    3.5.1 Completeness ......................................... 75
    3.5.2 Effective Temperatures ............................... 78
    3.5.3 Luminosities & Hertzsprung-Russel Diagram ......... 80
    3.5.4 Resolved Binary Separations ......................... 83
  3.6 Catalogue .................................................. 85
  3.7 Summary .................................................... 93

4 Properties of Nearby A Type Stars with Spitzer ............ 95
  4.1 Introduction ................................................ 95
  4.2 Data Reduction ............................................. 97
    4.2.1 MIPS-24 ............................................ 98
    4.2.2 MIPS-70 ............................................ 101
  4.3 Photosphere Photometry Prediction ......................... 106
    4.3.1 Models .............................................. 108
    4.3.2 Synthetic Photometry ................................. 111
    4.3.3 Grid of Photometric Colours ......................... 112
    4.3.4 Photometric Bands .................................. 113
    4.3.5 Zero Point Offsets .................................. 119
    4.3.6 Model Fitting ...................................... 121
    4.3.7 Predicting MIPS Photometry ......................... 122
    4.3.8 Peculiar Stars ...................................... 123
    4.3.9 Binaries .......................................... 125
  4.4 Calibration ................................................ 127
    4.4.1 MIPS-24 ............................................ 127
    4.4.2 MIPS-70 ............................................ 128
    4.4.3 PSF fitting vs. Aperture Photometry ................ 129
  4.5 Results .................................................... 131
    4.5.1 Observations Summary ................................ 131
    4.5.2 Stellar Parameters, Photometry and Excesses ........ 131
    4.5.3 Stellar Parameter Comparison ....................... 132
    4.5.4 Dust Properties .................................... 132
    4.5.5 Statistics ......................................... 147
5 LABOCA Data Reduction & Observations of Southern Circumstellar Discs 167

5.1 Introduction ................................................. 167
5.1.1 The GASPS Survey ...................................... 168
5.2 LABOCA Data Reduction ................................... 170
5.2.1 Flagging .................................................... 171
5.2.2 Calibrating the Time Series Data ...................... 172
5.2.3 Correlated Noise Removal .............................. 173
5.2.4 Frequency Domain Filtering ............................ 175
5.2.5 Baseline Subtraction ................................... 175
5.2.6 Final despiking ($\sigma$-clipping) ..................... 175
5.2.7 Map Projection ............................................. 175
5.2.8 Map Smoothing .......................................... 176
5.3 Calibration Observations .................................. 178
5.3.1 Pointing and Focus ..................................... 178
5.3.2 Sky-dips & Radiometer PWV Measurements .......... 178
5.3.3 Observations of Flux Calibrators ..................... 180
5.4 Photometric Analysis ...................................... 182
5.4.1 Measurement Types ..................................... 182
5.4.2 Measurements in this Work ............................ 182
5.4.3 Measuring Noise (RMS) in Maps ..................... 183
5.4.4 Extended Sources ....................................... 183
5.5 LABOCA Observations of Southern GASPS Targets .... 184
5.5.1 Instrument Problem During December Run .......... 184
5.5.2 $\eta$ Chamaeleontis ..................................... 188
5.5.3 Herbig Ae/Be Stars ..................................... 196
5.5.4 Other Stars ............................................. 208
5.6 Summary & Potential Follow-Up ......................... 209

6 Sub-millimetre Study of $\epsilon$ Indi and $\alpha$ Centauri with LABOCA 211

6.1 Introduction ................................................... 211
6.2 Observations and Data Reduction ......................... 212
6.3 $\epsilon$ Indi A+BC ............................................ 212
6.4 $\alpha$ Centauri AB ............................................ 215
6.4.1 Astrometry of the $\alpha$ Cen AB System ............... 217
6.4.2 Determining Pointing Offset and Bright Feature Motion 221
6.4.3 Stacking in the Frame of the Background .......... 225
6.4.4 Stacking in the Frame of the Stars ................. 227
6.5 Summary ...................................................... 231

7 Conclusions & Future of the Field 233

7.1 Summary & Conclusions .................................... 233
7.1.1 Debris Disc Surveys .................................... 233
7.1.2 Sub-mm Observations of Young Southern Stars ...... 237
7.2 Future ....................................................... 238
7.2.1 Science .................................................. 238
7.2.2 Observations .......................................... 239
7.2.3 Techniques & Ancillary data ........................ 243
7.2.4 Summary ................................................ 244

References ....................................................... 245
List of Figures

<table>
<thead>
<tr>
<th>Figure</th>
<th>Description</th>
<th>Page</th>
</tr>
</thead>
<tbody>
<tr>
<td>1.1</td>
<td>Debris discs shaped by gravitational interaction with planets</td>
<td>1</td>
</tr>
<tr>
<td>1.2</td>
<td>Debris discs imaged in scattered optical starlight</td>
<td>2</td>
</tr>
<tr>
<td>1.3</td>
<td>Sub-millimetre images of nearby debris discs</td>
<td>3</td>
</tr>
<tr>
<td>1.4</td>
<td>Locations of the Asteroid and Kuiper belts in the Solar System</td>
<td>4</td>
</tr>
<tr>
<td>1.5</td>
<td>Zodiagal light</td>
<td>6</td>
</tr>
<tr>
<td>1.6</td>
<td>Dynamical evolution of the outer Solar System in the Nice model</td>
<td>8</td>
</tr>
<tr>
<td>1.7</td>
<td>SED classification and evolutionary stages</td>
<td>11</td>
</tr>
<tr>
<td>1.8</td>
<td>Evolution of protoplanetary disc fraction of clusters</td>
<td>12</td>
</tr>
<tr>
<td>1.9</td>
<td>Schematic evolution of a typical protoplanetary disc</td>
<td>12</td>
</tr>
<tr>
<td>1.10</td>
<td>Observed giant planet – metallicity correlation</td>
<td>14</td>
</tr>
<tr>
<td>1.11</td>
<td>Evolution of PMS stars on the HR diagram</td>
<td>17</td>
</tr>
<tr>
<td>1.12</td>
<td>Nearby young stellar associations in heliocentric $UVWXYZ$ space</td>
<td>19</td>
</tr>
<tr>
<td>1.13</td>
<td>Temperatures of different grains as a function of distance from the Sun</td>
<td>23</td>
</tr>
<tr>
<td>1.14</td>
<td>Dust mass evolution of circumstellar discs from $10^5$ to $10^{10}$ yr</td>
<td>26</td>
</tr>
<tr>
<td>1.15</td>
<td>Vega imaged at 70 and 850 $\mu$m showing effects of radiation pressure</td>
<td>29</td>
</tr>
<tr>
<td>1.16</td>
<td>Poynting-Robertson drag</td>
<td>30</td>
</tr>
<tr>
<td>1.17</td>
<td>Sublimation rates of ices as a function of heliocentric distance</td>
<td>32</td>
</tr>
</tbody>
</table>

2.1 | Sky distribution of IRAS catalogue entries, and FSC magnitude distributions | 42 |
2.2 | IRAS distance limits for Kuiper and Asteroid belt analogues | 44 |
2.3 | (Sub-)Millimetre distance limits for Kuiper and Asteroid belt analogues | 48 |
2.4 | Spitzer/MIPS distance limits for Kuiper and Asteroid belt analogues | 49 |
2.5 | Photographs of the JCMT, SCUBA-2, SCUBA and UKT14 | 52 |
2.6 | Herschel | 54 |
2.7 | Dust mass sensitivity comparison for typical solar-type star | 57 |
2.8 | Herschel distance limits for Kuiper and Asteroid belt analogues | 58 |
2.9 | First results from the DEBRIS survey: three resolved discs | 59 |
2.10 | Preliminary flux density distribution for $\zeta$ Tuc using PACS photometry | 59 |
2.11 | Coverage of PACS and SPIRE scan maps used by DEBRIS | 60 |

3.1 | Evolution of spectral type, $T_{\text{eff}}$ and luminosity for low mass stars | 66 |
3.2 | Locus of luminosity classes on a HR diagram | 67 |
3.3 | B-V-colour-magnitude diagrams of primary stars with ZAMSs and isochrones | 68 |
3.4 | M type star classification method used by the PMSU survey | 72 |
3.5 | Distribution of systems in equatorial coordinates | 76 |
3.6 | Number of included systems in each subsample as a function of distance | 77 |
3.7 | Histogram of $T_{\text{eff}}$ of primaries | 79 |
3.8 | Gray et al. [2003, 2006] $T_{\text{eff}}$ vs. $(B-T - V_T)$ for primaries, with polynomial fit | 79 |
3.9 | Bolometric corrections from Bertone et al. [2004] as a function of $T_{\text{eff}}$ | 81 |
3.10 | HR diagram for UNS primary stars with polynomial fit | 81 |
3.11 | Histogram of $\log(L/L_\odot)$ of primaries | 82 |
<table>
<thead>
<tr>
<th>Section</th>
<th>Title</th>
<th>Page</th>
</tr>
</thead>
<tbody>
<tr>
<td>3.12</td>
<td>Histograms of projected physical binary separations for the subsamples</td>
<td>84</td>
</tr>
<tr>
<td>4.1</td>
<td>Summary of MIPS-24 small source observing mode</td>
<td>99</td>
</tr>
<tr>
<td>4.2</td>
<td>MIPS-24 PSF produced using DAOPHOT</td>
<td>100</td>
</tr>
<tr>
<td>4.3</td>
<td>Effect of sky value on PSF subtracted images</td>
<td>101</td>
</tr>
<tr>
<td>4.4</td>
<td>Summary of MIPS-70 coarse scale compact source observing mode</td>
<td>102</td>
</tr>
<tr>
<td>4.5</td>
<td>Effects of filtering on MIP S-70 images</td>
<td>104</td>
</tr>
<tr>
<td>4.6</td>
<td>MIPS-70 PSFs and column averages from filtered post-BCD images</td>
<td>105</td>
</tr>
<tr>
<td>4.7</td>
<td>Computing grid of photometric colours for fitting and predicting photometry</td>
<td>107</td>
</tr>
<tr>
<td>4.8</td>
<td>Effects of interpolation on Kurucz 1221 point flux distributions for λ &gt; 10 µm</td>
<td>110</td>
</tr>
<tr>
<td>4.9</td>
<td>Combining Johnson JK response functions with atmospheric transmission</td>
<td>118</td>
</tr>
<tr>
<td>4.10</td>
<td>Comparison of MIPS-24 observed PSF fit and predicted photometry</td>
<td>128</td>
</tr>
<tr>
<td>4.11</td>
<td>Comparison of MIPS-70 observed PSF fit and predicted photometry</td>
<td>129</td>
</tr>
<tr>
<td>4.12</td>
<td>Comparison of DAOPHOT PSF fit and aperture photometry measurements</td>
<td>130</td>
</tr>
<tr>
<td>4.13</td>
<td>Comparison of fitted photosphere parameters with literature values</td>
<td>133</td>
</tr>
<tr>
<td>4.14</td>
<td>Temperature vs. [24]−[70] relationship</td>
<td>134</td>
</tr>
<tr>
<td>4.15</td>
<td>Correction of orbital radius computed from [24]−[70] assuming black body grains</td>
<td>135</td>
</tr>
<tr>
<td>4.16</td>
<td>Excess significance histograms for all detected stars</td>
<td>150</td>
</tr>
<tr>
<td>4.17</td>
<td>Excess significance histograms for single and binary stars</td>
<td>151</td>
</tr>
<tr>
<td>4.18</td>
<td>Magnitude excesses as a function of physical binary separation</td>
<td>153</td>
</tr>
<tr>
<td>4.19</td>
<td>Distribution of metallicities ([M/H])</td>
<td>154</td>
</tr>
<tr>
<td>4.20</td>
<td>MIPS images of HD 37594</td>
<td>159</td>
</tr>
<tr>
<td>4.21</td>
<td>Optical colour image of HD 37594 and SF O 21</td>
<td>160</td>
</tr>
<tr>
<td>4.22</td>
<td>PSF subtracted MIPS images of HD 37594 showing resolved nature</td>
<td>161</td>
</tr>
<tr>
<td>4.23</td>
<td>MIPS images of the CCDM 23489-2808 ABC (UNS 109) system</td>
<td>162</td>
</tr>
</tbody>
</table>

| 5.1     | LABOCA scan patterns used in this work                                 | 171  |
| 5.2     | Empirical τ_(24) (870 µm) vs. PWV relationship from flux calibrators and sky-dips | 180  |
| 5.3     | LABOCA back-end DC offsets (box averaged) during December 2008          | 187  |
| 5.4     | Distance and age of the η Cha cluster in relation to other open clusters | 188  |
| 5.5     | Weight map of η Chamaeleontis with the 18 stars over-plotted           | 192  |
| 5.6     | Weight map of η Chamaeleontis with the 18 stars over-plotted           | 192  |
| 5.7     | Stacked map of all 18 members of η Chamaeleontis                      | 193  |
| 5.8     | Average disc mass of η Cha in evolutionary context                     | 194  |
| 5.9     | SED fit for RECX 15 including Herschel and LABOCA photometry           | 194  |
| 5.10    | Full LABOCA map of HD 97048 in relation to Chamaeleon I                | 201  |
| 5.11    | LABOCA maps of HD 97048, 1005-43 and 1005-46 with elliptical Gaussian fits | 202  |
| 5.12    | LABOCA map of HD 104237 with 2-source fit and subtraction              | 203  |
| 5.13    | SEDs for mee01 group I and II sources, and explanation by disc geometry | 207  |

| 6.1     | LABOCA map of ε Indi A+BC                                               | 214  |
| 6.2     | LABOCA maps of α Cen AB                                                 | 216  |
| 6.3     | Orbit of α Cen B relative to α Cen A projected on the sky               | 218  |
| 6.4     | Full sky motion as seen from Earth of α Cen AB                          | 220  |
| 6.5     | Cross-correlation masks for determining the pointing offset between epochs | 223  |
| 6.6     | Cross-correlation surfaces for the three masks                          | 224  |
| 6.7     | Fitting an elliptical Gaussian to the bright feature                    | 226  |
| 6.8     | α Cen AB relative to the Scorpius-Centaurus-Lupus-Crux complex          | 227  |
| 6.9     | MIPS-24 image of α Cen AB with PSF subtraction                          | 229  |
| 6.10    | Photometry of the stars (α Cen AB)                                      | 230  |

| 7.1     | Artist's impression of ALMA                                            | 241  |
List of Tables

3.1 Comparison of Gray and PMSU spectral types for stars within 15.7 pc ........ 73
3.2 Summary of subsample properties ........................................ 75
3.3 Reference abbreviations used in the text and tables ...................... 86
3.4 System info: primary star, distance, confusion estimate and survey membership ... 87
3.5 Component names, positions and proper motions .......................... 88
3.6 A-K primary spectral types, Tycho photometry and effective temperatures .... 89
3.7 M type primary spectral types, effective temperatures and $B_J, V_J$ photometry ... 90
3.8 Component cross identifications with common catalogues ............... 91
3.9 Notes for specific systems .................................................. 92
4.1 Properties of the grid of photometric colours ........................... 112
4.2 Adopted zero point offsets (ZPOs) for synthetic photometry .............. 120
4.3 Fitted stellar parameters and observed and predicted MIPS photometry ... 137
4.4 Disc parameters for stars with significant excess in both bands ........ 144
4.5 Disc parameter limits for stars with significant excess in only one band ... 145
4.6 Stars with MIPS-24 excess but no MIPS-70 photometry .................. 145
4.7 Orbit configurations for discs detected in multiple star systems ........ 146
4.8 Subsample excess properties .............................................. 149
4.9 The chemically peculiar stars and their excesses ......................... 157
5.1 Summary of LABOCA observations of GASPS targets ..................... 185
5.2 Flux calibrator measurements showing reduced sensitivity in Dec 2008 .... 186
5.3 $\eta$ Cha cluster member properties, $870\mu m$ photometry and mass limits ... 195
5.4 Properties of the four Herbig Ae/Be stars ................................ 196
5.5 (sub-)millimetre photometry, spectral power law slopes and disc masses for the four Herbig Ae/Be stars ........................................... 205
5.6 Summary of stars from nearby associations, with LABOCA photometry and dust mass limits ..................................................... 208
6.1 Observation summary ...................................................... 213
6.2 Positions of the centre of mass and A and B components of $\alpha$ Cen AB ........ 220
Chapter 1

Introduction

Ever since it was discovered in the sixteenth century that the planets of the Solar System orbit around the Sun, mankind has pondered the existence of planets orbiting other stars. We are now at an exciting time where we are able to observe and study systems of planets around hundreds of other stars. In addition to planets, the Solar System harbours a wide range of smaller bodies (planetesimals), from dwarf planets, asteroids and comets to grains of Zodiacal dust. The majority of the mass of these objects, which are thought to be leftovers of the planet formation process, resides in the Asteroid and Kuiper belts at distances of approximately 2.8 and 45 AU from the Sun respectively. Collisions and disintegration of planetesimals (e.g. comet trails) lead to the plane of the Solar System being littered with dusty debris, the spatial distribution of which is shaped by gravitational interactions with the planets (Fig. 1.1). Hundreds of such debris discs have now been detected around other stars, with studies showing that at least 10–30% of Sun-like and hotter stars possess such material.

Figure 1.1: Left: Artist’s impression of the ε Eridani debris disc, with dust confined to rings by gravitational interaction with planets (Image credit: NASA/JPL-Caltech/T. Pyle). Right: Present day distribution of dust generated in the Kuiper Belt, from models of Kuchner and Stark [2010].
CHAPTER 1. INTRODUCTION

Planetesimals and planets form from interstellar (primordial) dust and gas in dense circumstellar discs around newly born stars. These protoplanetary discs generally dissipate within the first 10 Myr after the birth of the star due to accretion and photoevaporation, leaving a system of planets, planetesimals and debris. Dusty circumstellar discs appear to be ubiquitous at all stellar ages and evolutionary stages – debris discs have even been discovered around white dwarfs [e.g. Su et al., 2007, Farihi et al., 2008].

There are three primary emission mechanisms by which circumstellar discs can be observed: optical and near-infrared (near-IR) starlight scattered by dust; gas emission lines; and thermal emission from dust at infrared to millimetre wavelengths. Detecting scattered starlight is challenging due to extreme contrast between the star and the disc surface brightness, however, such observations allow the spatial resolution of facilities such as the Hubble Space Telescope (HST) to be used to resolve the structure of discs such as in Fig. 1.2. Spectroscopic observations at near-IR to millimetre wavelengths are used to probe the conditions of the gas in protoplanetary discs, although interpretation of such observations requires complex models, and instrumental sensitivity typically only permits emission from a small selection of molecules to be detected (most notably carbon monoxide). To date there have been very few detections of gas in debris discs, although small quantities of gas are likely to exist in debris discs due to evaporation of ices from comets and dust grains which stray inwards towards the star from the cold outer regions.

Figure 1.2: Two debris discs around nearby A type stars imaged in scattered optical starlight using a coronograph to block the majority of the light from the stars. Both systems have direct detections of planets. Left: Fomalhaut at 7.7 pc imaged by HST [Kalas et al., 2008]. Right: β Pictoris at 19.4 pc imaged with 3.6 m ground based telescope using adaptive optics (central circle with planet detections from 8 m telescope using PSF subtraction rather than coronagraphy) [Lagrange et al., 2010].

Observing thermal emission from dust grains at infrared and longer wavelengths is by far the most prolific and generally applicable method of detecting and characterising circumstellar discs.
This is possible due to the large surface area per unit mass exhibited by dust grains, which allows dust masses of a few thousandths of an Earth mass \((M_\oplus)\) to routinely be detected. Generally the emission is not significantly spatially resolved at these wavelengths due to instrumental resolutions of \(\sim 5-20''\), although large discs have been resolved around nearby stars [e.g. Holland et al., 1998, Fig. 1.3].

![Image](image.png)

**Figure 1.3:** Sub-millimetre (850 \(\mu m\)) images of nearby debris discs, shown at the same physical scale. From left to right: \(\tau\) Ceti (G8.5 V, 3.7 pc), \(\epsilon\) Eridani (K2 V, 3.2 pc), Vega (A0 V, 7.7 pc), Fomalhaut (A4 V, 7.7 pc), \(\eta\) Corvi (F2 V, 18.3 pc).

The temperature of the majority of dust in discs is at least an order of magnitude colder than the stellar effective temperature. Consequently, the star to disc contrast decreases with increasing wavelength, allowing dust to be detected as excess flux compared to the expected stellar flux. This is the method by which essentially all known circumstellar discs have been detected, and this method is used throughout this thesis. Characterisation and modelling of discs generally relies upon photometric (or spectro-photometric) observations spanning a wide wavelength range to sample the spectral energy distribution (SED) of the dust.

This thesis primarily concerns surveys of nearby stars at wavelengths from 24 to 870 \(\mu m\) which aim to investigate statistically how debris disc incidence rates and properties vary with system properties such as stellar mass, metallicity, age, multiplicity and the presence of known planets. A study of young systems with ages encompassing protoplanetary disc dispersal and early debris disc evolution (5–30 Myr) using 870 \(\mu m\) observations is also performed. The contents of each chapter of the thesis is outlined at the end of this chapter. The remainder of this chapter presents supporting background and theory for the rest of the thesis.
1.1 The Solar System and its Debris Disc

Whilst we cannot be sure how typical it is of planetary systems in general, it is still worthwhile to consider the Solar System as a reference with which to understand and compare the features of other systems. The Solar System has the Sun at its centre: a main sequence (hydrogen burning) star with a mass of \( M_\odot \approx 2 \times 10^{30} \text{kg} \approx 3.3 \times 10^5 M_\oplus \) (\( M_\oplus \approx 6 \times 10^{24} \text{kg} \) is the mass of the Earth), an effective temperature (see §1.4.3 for definition) of \( \sim 5780 \text{K} \), a luminosity of \( \sim 4 \times 10^{26} \text{W} \), and a spectral type of G2 V. After the Sun, the next most massive objects are the four gas giant planets with orbital semi-major axes of \( \sim 5-30 \text{AU} \) and masses of \( \sim 15-300 M_\oplus \): Jupiter, Saturn, Uranus and Neptune. Then there are the four rocky terrestrial planets with orbital semi-major axes of \( \sim 0.3-1.5 \text{AU} \) and masses of \( \sim 0.05-1 M_\oplus \): Mercury, Venus, Earth and Mars, and satellites (moons) of the planets with masses up to \( \sim 0.025 M_\oplus \). In addition to the planets, which are all sufficiently massive to gravitationally dominate their orbits, there are many smaller bodies swarming throughout the Solar System. These range from the largest dwarf planets such as Eris and Pluto with masses of \( \sim 0.002 M_\oplus \) down to sub-micron sized grains of dust. The orbits of the planets and the vast majority of smaller bodies lie close to a plane known as the ecliptic, which is approximately perpendicular to the Sun’s spin axis.

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\text{Figure 1.4: Locations of the Asteroid and Kuiper belts in the Solar System. Top: the Solar System within Jupiter’s orbit (a = 5.2 AU), showing the Asteroid Belt. Bottom: The Solar System out to \( \sim 50 \text{AU} \), showing the Kuiper Belt outside of Neptune’s orbit (a = 30.1 AU). Image adapted from http://www.nasa.gov/mission_pages/spitzer/multimedia/20081027b.html (Image credit: NASA/JPL-Caltech).}
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There are three primary reservoirs of smaller bodies (planetesimals) in the Solar system. The Asteroid Belt lies between the orbits of Mars and Jupiter, with a total mass of \( \sim 6 \times \)
$10^{-4} M_\oplus$ and a mean orbit radius of 2.8 AU [Krasinsky et al., 2002]. The Asteroid Belt has been studied observationally for over two centuries (the largest object, Ceres, was discovered in 1801), facilitated by the relative proximity to the Earth, which currently allows bodies with sizes down to a few hundred metres to be detected. 

Outside the orbits of the gas giant planets, extending from approximately 30 AU to at least 50 AU, lies the Kuiper Belt\footnote{Although the same “Kuiper Belt” is most commonly used, including in this thesis, Gerard Kuiper was not the first astronomer to consider the existence of such a belt of planetesimals. Indeed Kuiper’s reasoning suggested that such a belt should no longer exist at the present epoch. The belt is also less commonly referred to as the Edgeworth-Kuiper Belt after Kenneth Edgeworth who had previously suggested such a belt and that it was a source of comets. See e.g. http://www.icq.eds.harvard.edu/ke.html for a more complete discussion.}. Observational study of the Kuiper Belt only really started in the early 1990s, with the detection of the first Kuiper Belt Object (KBO) other than Pluto in 1992 [Jewitt and Luu, 1993]. The difficulty of observing KBOs is simply due to their distance from the Earth and Sun, and our current inventory is still highly observationally biased due to the sky coverage of sufficiently deep surveys. Our knowledge of the total mass of objects in the Kuiper Belt is quite uncertain, although a value of $\sim 0.1 M_\oplus$ is most often assumed [Vitense et al., 2010, Jewitt et al., 1998]. The approximate locations of the Asteroid and Kuiper Belts in relation to the planets are shown in Fig. 1.4.

At much larger distances, of several thousand AU and greater, there is a third reservoir of bodies known as the Oort Cloud [Oort, 1950]. The distances to the Oort cloud preclude the observation of any objects in the cloud with the current instrumentation available. However, the Oort cloud is generally accepted as the source of long period comets such as the famous 1P/Halley and C/1995 O1 (Hale-Bop), which through close encounters with other objects have their orbits perturbed such that they travel into the inner Solar System. Observations of long period comets suggest that the Oort cloud is essentially spherical, and extends to perhaps $5 \times 10^4$ AU. The origin of the Oort cloud is uncertain. It is possible that a large fraction of the objects were not formed in the Solar System, but were instead captured from the Sun’s birth environment after being ejected from neighbouring systems [Levison et al., 2010].

In addition to these three substantial reservoirs of planetesimals, objects are also found in a wide variety of other locations in the Solar System. Example groups include Trojans orbiting at Lagrange points in the orbits of planets (most notably Jupiter), Near Earth Asteroids which pass inside the Asteroid Belt, Scattered Disc objects at Kuiper Belt and greater distances from the Sun, and comets which traverse both the outer and inner solar system on highly eccentric orbits. The Scattered Disc objects typically have perihelia (closest distance from the Sun) in the Kuiper Belt, however, they can have very large eccentricities which take them to aphelia (farthest distance from the Sun) of well over 100 AU. The comets are thought to originate in the Kuiper Belt (short period comets) and the Oort cloud (long period comets). Objects from
these regions contain significant fractions of ices (primarily H$_2$O, but also NH$_3$, CO, CO$_2$, CH$_4$
and others), which sublimate to form a comet tail when a comet comes close enough to the Sun
that its temperature exceeds around 100 K (see later).

1.1.1 Dust in the Solar System

The ecliptic plane is littered with small solid particles – dust grains – which can be seen in the
night sky as the zodiacal light (the zodiac refers to the constellations of stars which straddle
the ecliptic plane). The zodiacal light is a combination of scattered starlight which is visible
at optical wavelengths, and thermal emission which is most clearly visible from Earth at mid-
infrared wavelengths (e.g. Fig. 1.5). Due to observational bias, the zodiacal light observed from
Earth is primarily caused by small dust grains with sizes of $\sim$1–100 $\mu$m located less than $\sim$2 AU
from the Sun [Mann et al., 2006]. Measurements from spacecraft have, however, allowed the
properties of zodiacal dust to be explored out to 50 AU from the Sun [Humes, 1980, Grün et al.,
1993, Gurnett et al., 1997, Landgraf et al., 2002, Poppe et al., 2010]. Dust grains follow slightly
non-keplerian orbits due to radiation and solar wind forces explained later in this chapter,
resulting in their lifetime before they spiral into the Sun or are destroyed by collisions being
typically $\ll$1 Myr. This short lifetime in comparison to the 4.5 Gyr age of the Solar System
[Burns, 2010] suggests that the zodiacal dust must be continuously replenished.

![Figure 1.5: Zodiakal light, as seen in the mid-IR. Image is an all-sky map in galactic coordinates
from the Infrared Astronomical Satellite (IRAS). The zodiakal light is the diffuse blue band, with the
emission peaking between 12 and 25 $\mu$m (see e.g. Nesper et al. [2006, 2010] for IRAS zodiakal light

There are multiple sources of the dust: comet trails and spontaneous disruptions [Mukai,
1985, Liou et al., 1995, Lisse, 2002, Nesvorný et al., 2010], collisions of objects in the Asteroid
and Kuiper belts [Backman et al., 1995, Stern, 1996], interstellar dust grains captured as the
Solar System passes through the interstellar medium [Bertaux and Blamont, 1976, Grün et al.,]
1.1. THE SOLAR SYSTEM AND ITS DEBRIS DISC

1993, Krüger and Grün, 2009, Draine, 2009], and collisions of interstellar dust with Solar System objects [Yamamoto and Mukai, 1998]. The relative contributions of the different sources vary with distance from the Sun, and for different sizes of dust grains. The drag forces which act on grains cause them to spiral inward, so dust from comets and collisions can only contribute to regions inward from where it is generated. The grain size dependence of the contributions is due to the inherent grain size distributions of each dust source (e.g., interstellar dust grains typically have a maximum size of $\sim 1 \mu m$ [Kim et al., 1994]), and due to the size dependant effects of the drag forces and probabilities of destructive collisions between grains.

In the inner Solar System (approximately inward of Jupiter’s orbit at 5 AU) the dust, with a total mass of $\sim 2 \times 10^{-9} M_\odot$, is primarily generated by comets and collisions in the asteroid belt, with the former currently thought to be dominant [Nesvorný et al., 2010]. Between Jupiter and Saturn ($\sim$ 5–10 AU, i.e. outside of the asteroid belt) comets are by far the dominant source of dust [Landgraf et al., 2002]. Beyond Saturn there is very little cometary activity, and collisions in the Kuiper Belt are the dominant dust source. Models of collisional dust production in the Kuiper Belt combined with measurements from the Voyager spacecraft suggest that dust grains below $1 \text{ mm}$ in size are generated at a rate of $\sim 5 \times 10^4 \text{ kg/s}$ and the total mass of such grains in the Solar System is $\sim 10^{-6}$–$10^{-5} M_\odot$ [Gurnett et al., 1997, Jewitt and Luu, 2000, Landgraf et al., 2002, Moro-Martín and Malhotra, 2003]. By comparison, the masses of dust in detected debris discs around other stars are at least an order of magnitude larger than this, although as debris discs are currently only detected around $\sim 10\%$ of Sun-like stars [e.g. Trilling et al., 2008] it is still unclear whether the Solar System is more or less dusty than average. Determining how representative the Solar System and its debris disc are among planetary systems is one of the primary motivations of performing ever more sensitive surveys for debris discs around other stars.

1.1.2 Solar System Debris Evolution & the Late Heavy Bombardment

We, as life on Earth, are fortunate that there is not much more debris in the inner Solar System. Impacts of asteroids and comets with planets are thought to have significant implications for the evolution and survival of life [Kring, 2003], and have been attributed as the cause of mass extinction events on Earth [Schulte et al., 2010]. There is strong evidence from the cratering histories of the Moon and other bodies that the inner Solar System was a much more dangerous place in the past. In particular these cratering studies indicate a sharp decline in the rate of impacts 3.9 Gyr ago, with evidence for a peak in cratering activity just prior to this time. This peak in activity, $\sim 700 \text{ Myr}$ after the planets formed, is referred to as the Late Heavy

\footnote{Note that around more luminous stars than the Sun, radiation pressure (§1.6.1) can dominate over drag forces in the outer regions of systems causing small dust grains to move outward from where they are generated.}
Bombardment (LHB) or the Lunar Cataclysm [Tera et al., 1973, 1974, Strom et al., 2005].

\[ \Delta[70] = 4 \]

\[ \Delta[24] = 0 \]

Figure 1.6: Dynamical evolution of the outer Solar System in the Nice model. Left: epoch just prior to Jupiter–Saturn resonance crossing (879 Myr in the model). Centre: Just after the LHB has started (882 Myr). Right: 200 Myr after the LHB, when only 3% of the initial planetesimal population mass remains. Planet orbits: green = Jupiter, yellow = Saturn, cyan = Uranus, blue = Neptune. Image credit: Mark Booth.

The LHB has been replicated in dynamical models of the Solar System [the Nice model, Gomes et al., 2005]. In these models the gas giant planets were initially in more compact orbits, with Neptune’s orbit inside that of Uranus (see Fig. 1.6). Outside the orbits of the gas giants lay a belt of planetesimals with an inner edge at \( \sim 15.5 \text{ AU} \) and a mass of \( \sim 35M_\oplus \). Due to gravitational interaction between the planetesimal belt and the planets, the orbits of Jupiter and Saturn migrated. At an epoch of \( \sim 880 \text{ Myr} \) in the model, Jupiter and Saturn reached an orbital resonance \( (P_{\text{Saturn}}/P_{\text{Jupiter}} = 2) \), at which point the orbits of Neptune and Uranus became unstable and they scattered outwards into the planetesimal belt causing a massive disturbance of the planetesimals. This disruption of the planetesimal belt is the explanation of the LHB. The planetesimals scattered inwards were rapidly removed dynamically or by collisions, and the planetesimals scattered outwards which did not leave the Solar System form the Kuiper Belt and Scattered Disc seen today at >35 AU [Levison et al., 2008].

The combination of these models, the cratering records, and the present day planetesimal content of the Solar System (\( \lesssim 0.1M_\oplus \)) indicate that the planetesimal mass prior to the LHB was around three orders of magnitude greater than today. This would have made the Solar System readily detectable in the mid/far-IR by observers around stars in the solar neighbourhood.

With reference to the survey performed in chapter 4 of this thesis, the fractional excess flux of the pre-LHB Solar System debris disc at 24 and 70 \( \mu \)m would have been \( \sim 0.2 \) (\( \Delta[24] = 0.2^m \)) and \( \sim 70 \) (\( \Delta[70] = 4.6^m \)), making the Solar System among the brightest observed discs around Sun-like stars [Booth et al., 2009].
1.2 Classification and Evolution of Circumstellar Discs

Stars are formed from gas and dust in dense cores in molecular clouds. As a core collapses, angular momentum conservation dictates that any rotation of the material is amplified, causing the infalling material to be funnelled into a primordial disc. The collapsing material will form a star, or stars, at the centre. Further gravitational instabilities can lead to the formation of other stars or sub-stellar objects. Over time, the remaining primordial material surrounding the disc (the envelope) feeds into the disc or is blown away by radiation, outflows and winds from the forming protostars. The primordial disc evolves due to accretion of material onto the stars, and due to processing of the gas and dust in the disc. These discs of evolving material are the birth ground of planetesimals and planets, and as such are referred to as protoplanetary discs. Eventually, the protoplanetary disc orbiting the stars disperses, leaving bare (proto)stars, planets and planetesimals (asteroids and comets). The planetary system will continue to evolve as the orbits of the planets and planetesimals perturb one another. During this post-protoplanetary disc evolution, the destruction of planetesimals can generate significant quantities of dust, which form a debris disc.

1.2.1 Observational Classification and Evolution

Observationally, the evolution of young stellar objects (YSOs – protostars and their circumstellar material), has been classified by the characteristics of their spectral energy distributions (SEDs), as shown in Fig. 1.7. Initially, Lada and Shu [1990] proposed three classes of YSO based on the SED slope at near- and mid-infrared (∼2–50 μm) wavelengths. Lada Class I sources have spectral slopes which rise with increasing wavelength in the near/mid-IR. These objects are typically deeply embedded within molecular clouds. The emission from the protostar(s) is heavily obscured by surrounding material, and the majority of the luminosity is emitted from dust in the disc and envelope. Class II sources have near/mid-IR SED slopes which are decreasing with increasing wavelength. Whilst often still in the vicinity of molecular clouds, these are typically less embedded than class I sources. The (proto)star(s) in these sources are relatively unobscured unless the circumstellar disc is edge-on. Class III sources are essentially bare (proto)stars, with no significant near/mid-IR emission from circumstellar material.

Over the past two decades, advances in observing capabilities, particularly in the far-IR and (sub-)millimetre regimes (∼50 μm–2 mm), have required refinement and extension of the classification system of Lada and Shu [1990]. Greene et al. [1994] defined ranges of near/mid-IR slopes (νF_ν ∝ λ^α) to distinguish the three Lada classes, as well as defining ‘flat spectrum’ sources in between classes I and II (I: α > 0.3, flat: −0.3 < α < 0.3, II: −1.6 < α < −0.3, III: α < −1.6). Additionally, (sub-)millimetre observations have identified two further classes...
of YSOs which were, until recently, undetectable at mid-IR wavelengths. Class 0 sources, proposed by Andre et al. [1993], have SEDs dominated by envelope emission. The protostars are completely obscured, so the only near/mid-IR emission is due to nebulosity illuminated by the protostars, hot inner disc, or outflows. Secondly, dense cores in molecular clouds, with temperatures as low as \( \sim 10 \) K, have been detected [Ward-Thompson et al., 1994, 2007, and references therein]. These prestellar cores are the earliest evolutionary stage, although for each core it must be determined whether they are gravitationally unstable and hence will indeed collapse to form stars.

An extra class of system, transition objects (or transition discs), potentially in between class II and III systems in evolutionary terms, has been proposed ['transition' term first used in Strom et al., 1989]. These systems are very similar to class II systems, but lack circumstellar near-IR emission. The explanation for this is that the inner few AU of these systems have been largely cleared of material, most likely due to either photoevaporation caused by UV radiation from the stars [e.g. Alexander, 2008] or due to the gravitational influence of planets [e.g., Calvet et al., 2002, Megeath et al., 2005, Sicilia-Aguilar et al., 2009, Merín et al., 2010]. Although generally less massive than class-II discs, the transition objects are quite a heterogeneous group, with total disc masses of \( \ll 1 M_{\text{Jup}} \) to \( 40 M_{\text{Jup}} \) [Cieza et al., 2008b, 2010].

Schematic representations of the SEDs of these various classes of object are shown in Fig. 1.7. In this thesis, chapters two, three, four and six are primarily concerned with surveys for debris discs. Chapter five presents observations of a variety of young objects, from class I to III, including a number of transition objects. In chapter 6, it is proposed that a bright background source located approximately 50" from the nearby \( \alpha \) Centauri system is a prestellar core.

The timescales of the various phases, and indeed whether or not the SED classes strictly correspond to a series of evolutionary stages, is a matter of debate [e.g., Evans et al., 2009]. However, studies of large numbers of stars in young clusters have shown that the fraction of stars with near/mid-IR detected discs drops rapidly over the age range 3–10 Myr [Strom et al., 1989, Haisch et al., 2001, Hillenbrand, 2005]. This decline in disc fraction is shown in Fig. 1.8. The decline can be modelled as an exponential decay with an e-folding time of 2.5 Myr [Williams and Cieza, 2011]. Surveys of the same clusters at far-IR and sub-millimetre wavelengths [e.g. Andrews and Williams, 2005, 2007] typically find very few extra discs than seen at near/mid-IR wavelengths, suggesting that the dispersal process is rapid \( \lesssim 0.5 \) Myr Williams and Cieza, 2011], with the cold outer disc being removed soon after the hot inner accretion disc.

A high-level physical description of the evolution of protoplanetary discs is shown in Fig. 1.9. The sub-figures (a)–(d) approximately correspond to SED class I, II, transition object and debris disc respectively, although no envelope is shown in (a).
1.2. CLASSIFICATION AND EVOLUTION OF CIRCUMSTELLAR DISCS

Figure 1.7: SED classification and evolutionary stages. Dashed lines and grey regions: (proto)star as a reddened black body; solid lines and pink regions: complete SED. Prestellar core: cold ($T \sim 10$–20 K), dense region in a molecular cloud in the process of infall. Class 0: SED dominated by envelope, protostar has formed but is highly extincted, any near-IR emission is from nebulous. Class I: SED rising in near/mid-IR, dominated by massive flared circumstellar disc and remaining envelope, star moderately extincted. Class II: SED falling in near/mid-IR, SED dominated by circumstellar disc, little envelope, star relatively un-extincted. Transition Object (TO): A Class II object with a lack of circumstellar near-IR emission, indicating the inner disc regions have been largely cleared. Class III: Naked star on or approaching the ZAMS, potentially with planets and a debris disc.
**Figure 1.8:** Evolution of protoplanetary disc fraction of clusters, determined by the presence of a near-IR excess – i.e. Class I–II sources. This shows that the inner regions of most discs are cleared by an age of 5 Myr. Taken from Wyatt [2008].

**Figure 1.9:** Schematic evolution of a typical protoplanetary disc, seen in cross-section. Taken from Williams and Cieza [2011].
1.2.2 Disc Evolution and Planet Formation Models

The evolution of circumstellar discs from primordial discs of interstellar gas and dust to mature systems of planets and planetesimals is highly complex, and many of the processes involved are still not well understood.

It is widely accepted that rocky planets and planetesimals are formed by the process of core accretion [e.g. Pollack et al., 1996]. The core accretion theory explains planet formation as a bottom-up process, starting from the coagulation of dust grains. Whilst the core accretion theory can explain the formation of gas giant planets, it is also possible that these can be formed by a top-down process where a region of a disc collapses due to gravitational instabilities [e.g. Boss, 1997].

Core Accretion

The initial coagulation of dust grains is facilitated by the gas in the disc, which orbits at slightly sub-keplerian velocities due to a radial pressure gradient. The gas exerts a drag force on the dust grains, which causes a range of dust orbital velocities between the gas velocity and the keplerian velocity. The small relative velocities between grains allow for collisions in which the grains loosely stick together to form aggregates. Growth to kilometer size planetesimals is thought to occur rapidly in $\sim 10^4$ yr [Weidenschilling, 1988], at which point self-gravity becomes significant in holding the objects together. Growth of objects at this stage can be slow, however, once the mass becomes large enough that the escape velocity significantly exceeds the relative velocities between planetesimals, the accretion rate starts to increase with mass and a period of runaway growth ensues, creating lunar mass protoplanets in $\sim 10^5$ yr [Wetherill and Stewart, 1989]. The runaway growth stops when a protoplanet has cleared its orbit of objects to accrete. Objects formed in protoplanetary discs tend to have very low eccentricity orbits due to the drag from the gas. Gravitational interactions between protoplanets and planetesimals, however, eventually increase the eccentricities of the smaller bodies such that further collisions occur. This stage is referred to as oligarchic growth, and can proceed for as long as $\sim 10^6$ yr (long after gas dispersal from protoplanetary discs), before terrestrial mass planets are formed [Wetherill, 1990].

If a protoplanet reaches a critical mass of approximately $10M_{\oplus}$ while a significant amount of gas remains in the disc, then gas will rapidly be accreted onto the protoplanet, forming a gas giant planet [Pollack, 1984]. If gas giant planets are formed by core accretion, their presence in a system implies that critical mass ($\sim 10M_{\oplus}$) cores were formed within a few million years in order for rapid gas accretion to have been able to take place. The timescale of planetesimal accretion is proportional to the density of dust and planetesimals in the disc [e.g. Rice and Armitage, 2005], so the likelihood of a disc to produce gas giant planets increases with the mass...
of the disc and the metallicity (dust:gas mass ratio).

There are a number challenges to the core accretion model which planet formation researchers are currently trying to overcome. The greatest of these are to do with the growth of objects in the centimetre to kilometer size regime. The first issue is that as objects grow to centimetres or so in size, rather than being entrained in the gas as smaller grains are, they follow near keplerian orbits, with gas drag exerting a force that causes them to spiral in towards the star where they will be accreted. The second issue is that aggregates of order 100 m in size, which are too small for self-gravity to hold together, are especially fragile and can be easily destroyed in even very low velocity collisions [Stewart and Leinhardt, 2009], making it difficult to build kilometer and larger size bodies which are gravitationally bound. The other significant issue is that the time required to build cores massive enough to become gas giant planets is typically similar to the dissipation timescales of discs [e.g. Rafikov, 2011].

Gas Giant Planet – Metallicity Correlation

The vast majority of currently known exoplanets (planets around stars other than the Sun) have been detected by measuring the radial velocity variation of the host star due to its orbit around the system barycentre. This technique is primarily sensitive to massive planets with short orbital periods, as these cause larger radial velocity variations with shorter periods [e.g. Udry and Santos, 2007]. As a result the majority of detected planets are gas giant planets in close orbit around their star, although the timespan of surveys is now long enough that orbits out to ~5 AU have been detected.

![Figure 1.10: Observed giant planet – metallicity correlation. Taken from Fischer and Valenti [2005].](image)

The observational bias of planet detection surveys currently allows few statistical conclusions to be drawn about the properties of exoplanets (although they clearly show that there are numerous systems quite unlike the Solar System). One robust trend that has been found,
however, is that the probability of detecting a gas giant planet is strongly correlated with the host star metallicity, as shown in Fig. 1.10 [Santos et al., 2004, Fischer and Valenti, 2005].

This correlation is consistent with the formation of gas giant planets by core accretion [Ida and Lin, 2004, Wyatt et al., 2007a], as the time taken to form critical mass cores is lower in more metal rich (i.e., more dusty) discs of a given mass, so they are more likely to be formed before the gas disperses from the disc. If the giant planets were formed directly by gravitational collapse then either no correlation with metallicity [Boss, 2002] or an anti-correlation with metallicity [Cai et al., 2006] would be expected (this is due to the effect of metallicity on the cooling rate of overdense regions, which is related to their gravitational stability).

Investigating the metallicity dependence of the detection rate of debris discs is currently the best method to observationally probe the metallicity dependence of the formation of smaller, rocky bodies [e.g., chapter 4 of this thesis; Greaves et al., 2006, Saffe et al., 2008]. These studies show no metallicity dependence on the likelihood for a system to at least form planetesimals.
1.3 Young Stars

Closely related to the evolution and classification of circumstellar discs is the evolution of the host stars and the surrounding birth environment. Young stars are rarely found in isolation, and studies of the evolution of circumstellar discs are dependant on the clustering of stars born at similar epochs in order to reliably estimate ages. The following two subsections give an overview of the pre-main sequence (PMS) evolution of stars and the clustering of young stars near the Sun. The timescales of PMS evolution, and particularly how these vary with stellar mass, are important to bare in mind when studying a young group of stars. The groups of young stars near the Sun give general context to observations of nearby star formation and circumstellar disc studies. These topics are particularly relevant to chapter 5 of this thesis.

1.3.1 Pre-Main Sequence Stellar Evolution

As a protostar collapses, it eventually reaches a state of quasi-hydrostatic equilibrium, where pressure balances the force of gravity. At this point the effective temperature is $\sim 3500$ K, approximately independent of stellar mass. As the protostar contracts, the temperature remains approximately constant, so the evolution on a Hertzsprung-Russel (HR) diagram, as shown in Fig. 1.11, is vertically downward [Hayashi, 1966]. This phase is referred to as the Hayashi track, and it continues while convection occurs throughout the majority of the stellar interior. For stars more massive than approximately $0.5 \ M_\odot$, a significant (containing greater than half the stellar mass) radiative core develops during the collapse, at which point the effective temperature starts to increase, causing the evolutionary track on the HR diagram to become approximately horizontal as the protostar collapses further. Eventually, the core temperature of the protostar becomes sufficient for nuclear fusion of Hydrogen to ignite, which halts the contraction and marks the start of the main sequence life of the star. The position of stars in the HR diagram when they reach the main sequence is referred to as the zero age main sequence (ZAMS). The timescales of the contraction depend on the luminosity of the protostar. For a solar mass star, the Hayashi track phase lasts for approximately 10 Myr, nuclear burning starts at approximately 30 Myr, and the ZAMS is reached at approximately 60 Myr [Hayashi, 1966]. For stars more massive than a solar mass, the Hayashi track timescale is negligible as they rapidly develop radiative cores. For a 2.3 $M_\odot$ star (corresponding to an early A-type main sequence star) the total time to reach the ZAMS is approximately 3 Myr [Henyey et al., 1955]. Pre-main sequence (PMS) evolutionary tracks and isochrones on a HR diagram, starting from 0.1 Myr, are shown in Fig. 1.11.

With reference to Fig. 1.11, it can be seen that when young groups of stars are observed, with ages $\lesssim 10$ Myr, stars with masses below about 1 $M_\odot$ will still be on Hayashi tracks, with
1.3. YOUNG STARS

Figure 1.11: PMS evolutionary tracks for stars of mass 0.1 to 6.0\(M_\odot\) in the HR diagram, from an age of 0.1 Myr to the ZAMS. Isochrones at ages of 1, 10 and 100 Myr are shown with dotted lines. Note that, for instance, at an age of 10 Myr stars with \(M \gtrsim 2.0 M_\odot\) will be on the ZAMS, whereas lower mass stars will typically still be above and to the right of the ZAMS (more luminous and redder). Taken from Palla and Stahler [1999].

Effective temperatures of 3000–4000 K (spectral types K/M). More massive stars will already be approaching the ZAMS, with effective temperatures above 5000 K (G and earlier type). This has led to a distinction being made between YSOs with stellar masses above and below \(\sim 1-2 M_\odot\). The low mass stars are referred to as T Tauri stars [Joy, 1945], and the higher mass stars are referred to as Herbig Ae/Be\(^3\) stars [Herbig, 1962]. Both groups of stars share the same observational features which can distinguish them from main sequence stars of the same spectral types: luminosity above the ZAMS, irregular optical variability, Hydrogen Balmer \(\alpha\) (H\(\alpha\)) and other emission lines, high Lithium abundance, and ultraviolet and infrared excess continuum emission. Note that Herbig Ae/Be stars may already be on the ZAMS. The T Tauri stars are further subdivided based on the strength of H\(\alpha\) emission (the equivalent width), which is a measure of the rate at which material is being accreted onto the star, and the presence of any detectable infrared excess. Classical T Tauri stars (CTTS) have strong H\(\alpha\) emission (equivalent width \(\gtrsim 5-20\) Å), and typically significant excess at all infrared wavelengths. Weak-lined T Tauri stars (WTTS) have weak H\(\alpha\) emission (equivalent width \(\lesssim 5-20\) Å) and typically have less infrared excess than CTTS. If a WTTS has no detectable infrared excess then it is

\(^3\)The less common Fe and Ge type stars are often referred to as Herbig stars as well
classified as a Naked T Tauri star (NTTS). More evolved examples of WTTS/NTTS, which are still above the main sequence but show Lithium depletion, are referred to as post T Tauri stars (PTTS). Lithium is rapidly removed from the photospheres of convective stars once the core temperature reaches approximately $2.5 \times 10^6$ K, which typically occurs at ages of more than 10 Myr in T Tauri stars.

1.3.2 Young Stellar Associations and Moving Groups

Stars are formed in molecular clouds, which typically disperse on Myr timescales [e.g. Hartmann, 2001, and references therein] due to clearing by radiation and winds from the newly formed stars. As a consequence, stars formed in a particular region tend to have ages within at most a few Myr of one another. The stars inherit their velocities from the cloud, and initially the relative motions between members are small. Whilst dense clusters$^4$ may remain bound as open clusters for several hundred Myr [Wielen, 1971], smaller clusters and loose associations of stars disperse over timescales of several to tens of Myr. The members of such clusters and associations can retain similar space velocities for periods much longer than the dispersal timescale. The remnants of associations and clusters which are in the process of dispersing are referred to as moving groups or streams [Larson, 2002]. As the members of moving groups can usually be assumed to have formed at approximately the same epoch, ages accurate to 10–20% can be determined by averaging ages determined for individual members or by isochrone fitting to Hertzsprung-Russel or colour-magnitude diagrams.

The first moving groups identified in the solar neighbourhood (within \( \sim 100 \) pc) were the Ursa Major$^5$ and Hyades super-clusters, with ages of approximately 300 and 600 Myr respectively [Zuckerman and Song, 2004]. At larger distances of \( \sim 100–600 \) pc there are several moving groups dominated by O and B type stars – the OB associations [see e.g. de Zeeuw et al., 1999]. The three nearest OB associations – Lower Centaurus Crux, Upper Centaurus Lupus and Upper Scorpius, with mean distances of 116–148 pc and ages of 5–17 Myr [de Zeeuw et al., 1999, Preibisch and Mamajek, 2008] – are dynamically related, and together form the Sco OB2 or Sco-Cen association.

Within the last two decades several young moving groups of \( \sim 10–100 \) members, with ages of \( \lesssim 50 \) Myr, have been identified within \( \sim 100 \) pc of the Sun [see review in Torres et al., 2008]. The heliocentric positions and space motions of stars in several of these groups are shown in Fig. 1.12. As the spatial extent of these groups are often 100 pc or more, they are spread all over the sky. Member stars are identified based on signs of youth (X-ray emission, Li abundance,

$^4$A cluster is typically defined as having at least an order of magnitude higher density than the field. Associations typically have similar or lower densities than the field.

$^5$The Ursa Major group is also referred to as the Sirius group, as Sirius is its brightest member.
Figure 1.1.2: Nearby young stellar associations in heliocentric U/V/W/X/Y/Z space. Taken from Torres et al. [2008]. Filled stars: β Pic; filled circles: Tuc-Hor; open circles: Columbia; crosses: Carina; filled triangles: TW Hya; open triangles: ε Cha; open stars: Octans; open squares: Argus; filled squares: AB Dor.

Hα emission etc.) and their space motion. The ages of the groups shown in Fig.1.12 span 6–70 Myr.

The knowledge of the ages, combined with the proximity and youth of the stars in these moving groups has made them prime targets for studies of the evolution of stars and circumstellar discs at epochs after molecular cloud dispersal. Examples of such studies are discussed throughout this thesis, and all the target stars observed in chapter five are members of associations or moving groups.
1.4 Thermal Emission

1.4.1 Black Body Radiation Properties

The irradiance per unit frequency (power per unit area of emitting body, per unit solid angle, per unit frequency) of a black body with temperature $T$, at frequency $\nu$ or wavelength $\lambda$, is given by Planck’s function,

$$ B_\nu(T) = \frac{2h\nu^3}{c^2} \frac{1}{e^{h\nu/kT} - 1} = \frac{2hc}{\lambda^3} \frac{1}{e^{hc/\lambda kT} - 1} $$

(1.1)

where $h = 6.626 \times 10^{-34}$ Js is Planck’s constant, $c$ is the speed of light, and $k = 1.381 \times 10^{-23}$ m$^2$kg s$^{-2}$K$^{-1}$ is Boltzmann’s constant. This function has a single maximum, and tends to zero as $\nu$ tends to zero and infinity. The maximum of $B_\nu$ occurs at a wavelength of

$$ \lambda_{\text{max}} = \frac{5098 \mu m}{T} $$

(1.2)

It is useful to consider the cases where the exponent, $h\nu/kT$, is significantly less than and greater than one. By noting the series expansion of an exponential,

$$ e^x = \sum_{n=0}^{\infty} \frac{x^n}{n!} = 1 + x + \frac{x^2}{2} + \frac{x^3}{6} + ..., $$

(1.3)

it is seen that when the exponent is $\ll 1$, the second order and higher terms can be neglected, allowing the Planck function to be simplified to,

$$ \lim_{h\nu/kT \ll 1} B_\nu(T) = \frac{2\nu^2 kT}{c^2} \frac{1}{\nu} = \frac{2kT}{\lambda^2} $$

(1.4)

This is the Rayleigh-Jeans law, which applies when $\lambda \gg 10^4 \mu m/T$. In the limit where the exponent is $\gg 1$, the $-1$ in the denominator can be ignored. This yields Wein’s law,

$$ \lim_{h\nu/kT \gg 1} B_\nu(T) = \frac{2h\nu^3}{c^2} e^{-h\nu/kT} = \frac{2hc}{\lambda^3} e^{-hc/\lambda kT} $$

(1.5)

which applies when $\lambda \ll 10^4 \mu m/T$.

The above equations are for irradiance per unit frequency. The corresponding versions for irradiance per unit wavelength can be obtained by noting that

$$ F_\lambda = \frac{d\nu}{d\lambda} F_\nu = \frac{\lambda^2}{c} F_\nu $$

(1.6)
The wavelength which maximises $B_\lambda$ is given by Wein’s displacement law,

$$\lambda_{\text{max}} = \frac{2898 \, \mu m}{T}. \quad (1.7)$$

### 1.4.2 Emission from Real Objects

Real objects do not emit as perfect black bodies. This can be accounted for by introducing a wavelength dependant emissivity, $\epsilon_\nu$, such that the spectral irradiance is given by:

$$I_\nu(T) = \epsilon_\nu B_\nu(T). \quad (1.8)$$

For an object in thermal equilibrium, the absorptivity is equal to the emissivity (Kirchoff’s law of thermal radiation).

### 1.4.3 Luminosity and Effective Temperature

The luminosity (total radiated power) of a black body of area $A$ and temperature $T$ is

$$L = A \int_0^\infty B_\nu(T)d\nu \int d\Omega = A\sigma_{SB}T^4 \quad (1.9)$$

where $\Omega$ is integrated over the half-sphere, and $\sigma_{SB} = 5.670 \times 10^{-8} \, \text{W m}^{-2}\text{K}^{-4}$ is the Stefan-Boltzmann constant. For objects which are not a perfect black body, an effective temperature can be defined for any point on a surface such that,

$$\int_0^\infty I_\nu(T)d\nu \int d\Omega = \int_0^\infty \epsilon_\nu B_\nu(T)d\nu \int d\Omega = \sigma_{SB}T_{\text{eff}}^4. \quad (1.10)$$

The effective temperature is one of the primary physical parameters used to describe stars. Although a single $T_{\text{eff}}$ is usually ascribed to a star, in reality $T_{\text{eff}}$ varies over the surface due to star spots (patches of low $T_{\text{eff}}$), and oblateness caused by rotation.
1.5 Dust Emission and Absorption

1.5.1 Grain Temperatures and Disc Radii

Grains can be considered to be in thermal equilibrium, which means that the absorbed and emitted power are equal. In an optically thin disc, such as the majority of debris discs, the absorbed power is due entirely to the light from the star. For an isothermal grain (small, thermally conductive, or rotating) of radius $a$ and temperature $T_g$, treating the star as a spherical black body of effective temperature $T_*$ and radius $R_*$ gives,

$$
\pi a^2 \left( \frac{R_*}{r} \right)^2 \int_0^\infty \epsilon_\nu B_\nu(T_g) d\nu = 4\pi a^2 \int_0^\infty \epsilon_\nu B_\nu(T_g) d\nu,
$$

where $\epsilon_\nu$ is the grain’s wavelength dependant emissivity (and absorptivity). If a grain behaves as a perfect black body ($\epsilon_\nu = 1$) at wavelengths covering the majority of the stellar emission and the grain’s emission, then Eqn. (1.11) simplifies to,

$$
\pi a^2 \frac{L_*}{4\pi r^2} = 4\pi a^2 \sigma_{SB} T_g^4,
$$

and therefore,

$$
T_g = \left( \frac{L_*}{16\pi r^2 \sigma_{SB}} \right)^{1/4} = 278 \left( \frac{L_*}{L_\odot} \right)^{1/4} \left( \frac{r}{1\text{ AU}} \right)^{-1/2} \text{ K},
$$

In reality, $\epsilon_\nu$ is approximately constant at wavelengths shorter than the size of a grain, but at longer wavelengths $\epsilon_\nu$ decreases. The variation of $\epsilon_\nu$ at longer wavelengths is usually well approximated by a power law in wavelength, with typical exponent values between 1 and 2. The exponent is $\sim 1$ for amorphous materials, $\sim 2$ for crystalline or metallic materials, and $\sim 1.5$ for typical interstellar dust [Seki and Yamamoto, 1980, Witt, 1989, Backman and Paresce, 1993].

Defining a critical wavelength, $\lambda_0$, at which $\epsilon_\nu$ transitions from a constant to a power law, the grain temperature in two useful cases can be computed [Backman and Paresce, 1993]. If a grain is much larger than the wavelength of the peak in the stellar emission ($a \gtrsim 1\mu m$), but is smaller than the peak in the grain’s emission, then the temperature of amorphous grains is,

$$
T_g = 468 \left( \frac{L_*}{L_\odot} \right)^{1/5} \left( \frac{r}{1\text{ AU}} \right)^{-2/5} \left( \frac{\lambda_0}{1\mu m} \right)^{-1/5} \text{ K}.
$$

This case corresponds well with the typical scenario in debris discs. For ISM dust grains, which are so small that $\lambda_0$ is shortward of the majority of stellar emission, the grain temperature is,

$$
T_g = 636 \left( \frac{L_*}{L_\odot} \right)^{2/11} \left( \frac{r}{1\text{ AU}} \right)^{-4/11} \left( \frac{T_*}{T_\odot} \right)^{3/11} \text{ K}.
$$

(1.14)
Equations (1.13), (1.14) and (1.15) are plotted as a function of \( r \) for a solar-type star in Fig. 1.13. Note that Eqn. (1.14) is only valid when grains are cold enough that the majority of the grain emission is at wavelengths longer than \( \lambda_0 \) (\( \lambda_0 \approx a \)). In Fig. 1.13 the temperatures from (1.14) are only plotted when above the black body temperature – the grains act as black bodies at higher temperatures.

\[ T_g / K \]

\[ r / AU \]

**Figure 1.13:** Temperatures of different grains as a function of distance from a solar like star. The black body temperature is from Eqn. (1.13). The temperature of amorphous grains is from Eqn. (1.14), with \( \lambda_0 = a \), when it is above the black body temperature. The temperature of ISM grains is from (1.15).

From Fig. 1.13 it can be seen that the grain temperature can vary by a factor of several at a particular distance from the star, with the potential variation increasing with distance. For the majority of circumstellar discs, the only means of determining their spatial structure is by modelling their SEDs, as they are too small or have surface brightnesses too low to image. The SED fitting will typically determine the physical temperature of dust grains. The distance of the dust from the star must then be inferred from relations such as Eqns. (1.13), (1.14) and (1.15). It is rarely possible to also determine the grain properties from the SED (at least for the outer regions of discs), so there is considerable uncertainty in distances derived from temperatures. For debris discs it is common to state distances for black body grains from Eqn. (1.13), with the caveat that the real distance will be larger by up to a factor of \( \sim 3 \) [e.g., Wyatt, 2008].
1.5.2 Disc Emission and Dust Mass

For any group of thermal emitters at a temperature $T$, a cross-section, $\sigma(\lambda)$, can be defined such that the observed flux density at a given wavelength can be determined from,

$$F_\nu(\lambda, T) = \frac{\sigma(\lambda) B_\nu(\lambda, T)}{d^2},$$  \hspace{1cm} (1.16)

where $d$ is the distance from the observer. For the case of an optically thin medium, $\sigma$ can be expressed in terms of the mass of the emitters,

$$F_\nu(\lambda, T) = \frac{\kappa(\lambda) M_{\text{dust}} B_\nu(\lambda, T)}{d^2},$$  \hspace{1cm} (1.17)

where $\kappa(\lambda)$, with dimensions of area per unit mass, is called the opacity of the medium. If the emitters were all identical spherical grains, then the opacity would simply be,

$$\kappa(\lambda) = \frac{\pi a^2 \epsilon(\lambda)}{\lambda} = \frac{3\epsilon(\lambda)}{4\rho}$$  \hspace{1cm} (1.18)

As described above, the emissivity of grains is approximately constant for $a \gtrsim \lambda$, so at any wavelength, grains much larger than the wavelength ($a \gtrsim 10 \lambda$) contribute little to the opacity if smaller grains are present, due to the $1/a$ dependence of $\kappa$. The exact wavelength dependence of the opacity depends on the grain size distribution and the grain types (emissivity power law exponent, as described above). It is also highly unlikely that even in a narrow ring of dust, as is often seen in debris discs, there will be a single grain temperature, due to the different grain sizes (Fig. 1.13). Even with all these possible variables, it has been found that the SEDs of many dusty media, including the ISM, most debris discs, and the outer regions of protoplanetary discs, can be well fitted by a single characteristic temperature and an opacity of the form,

$$\kappa(\lambda) = \begin{cases} 
\kappa_0 & \lambda \leq \lambda_0 \\
\kappa_0 \left(\frac{\lambda}{\lambda_0}\right)^{-\beta} & \lambda > \lambda_0
\end{cases}$$  \hspace{1cm} (1.19)

The critical wavelength, $\lambda_0$, and the opacity power law exponent, $\beta$, can be determined by fitting the SED of a source. A reference opacity, either $\kappa_0$ or $\kappa$ at some fiducial wavelength, must however be chosen, as the dust mass is not known. The choice of reference opacity is discussed below. For each type of dusty medium there will usually be typical values of $\lambda_0$ and $\beta$. If the grain size distribution is expected to be dominated by grains smaller than the wavelengths of the majority of the grain thermal emission, e.g. for the ISM or young primordial/protoplanetary discs, then $\kappa(\lambda) \propto \lambda^{-\beta}$ is assumed to hold at all wavelengths e.g., Andrews and Williams,
2005]. Typically $\beta \sim 2$ for the ISM. For debris discs $\beta \sim 1$ and $\lambda_0 \sim 200 \mu\text{m}$ are typical, although there can be considerable variation in $\lambda_0$ [e.g., Wyatt, 2008]. The value of $\lambda_0$ can be thought of as a characteristic grain size, as the spectrum would be the same if all grains had this size and their emissivity power law exponent was equal to $\beta$.

Occasionally a continuous function is used for the opacity, in place of Eqn. (1.19) [e.g. Najita and Williams, 2005],

$$\kappa(\lambda) = \kappa_0 \left( 1 - e^{-\left(\frac{\lambda}{\lambda_0}\right)^{-\beta}} \right),$$

which can be seen from Eqn. (1.3) to give the desired behaviour for $\lambda \gg \lambda_0$ and $\lambda \ll \lambda_0$, and a smooth transition in between.

### 1.5.3 Effect of Optical Depth

Above it was assumed that the dusty medium was optically thin along the line of sight to the observer at all wavelengths. The optical depth, $\tau(\lambda)$, for a medium of uniform line of sight thickness, and total cross sectional area, $A$, is,

$$\tau(\lambda) = \frac{\kappa(\lambda) M_{\text{dust}}}{A}. \quad (1.21)$$

The observed flux density is then,

$$F_\nu(\lambda, T) = \left( 1 - e^{-\kappa(\lambda) M_{\text{dust}} A} \right) \frac{B_\nu(\lambda, T)}{d^2}, \quad (1.22)$$

which simplifies to Eqn. (1.17) for the optically thin case, $\tau \ll 1$. The effect of Eqn. (1.22) on the observed spectrum is equivalent to reducing $\beta$ at wavelengths for which $\tau$ is significant. It can be seen that this is a similar effect to increasing $\lambda_0$ defined above.

### 1.5.4 Measuring Dust Masses

The mass of dust can be calculated from Eqn. (1.22) or (1.17) if the opacity at some wavelength is assumed and $T_{\text{dust}}$, $\beta$ and $\lambda_0$ can be determined. There are several complications to this. Firstly, the opacity at any wavelength is dominated by grains with a similar size to the wavelength, if present in the system, as smaller grains have lower emissivity, and larger grains naturally contribute less to the opacity (Eqn. (1.18)). Secondly, unless there is a means for computing the optical depth to a reasonably high accuracy, it is not possible to use Eqn. (1.22), and only a lower limit of the mass can be determined. As the optical depth decreases with increasing wavelength, it is desirable to use photometry at wavelengths long enough that the optical depth is assured to be negligible.
For debris discs, which are generally optically thin at all wavelengths, it is also desirable to determine dust masses from the longest wavelengths possible. As will be described later, the smaller the size of grains, the more susceptible they are to forces which can greatly perturb their orbits or remove them from a system. As a result the fraction of the total dust mass which is contained in the small ($a \lesssim 100\,\mu m$) grains can vary greatly between discs. The small grains will also generally have a wide range of temperatures due to having a much wider range of orbits than the larger grains.

The desire to use the longest possible wavelengths to weigh discs must be balanced with the fact that the flux density drops as $\lambda^{-(2+\beta)}$ in the Rayleigh-Jeans limit. The best compromise of immunity to unknown optical depth and uncertain small grain properties, versus sensitivity to detect the dust, is typically found in the (sub-) millimetre regime. The most commonly used wavelengths are around $850\,\mu m$ and $1.3\,mm$, as these can be observed efficiently with large ($10-30\,m$) ground based telescopes.

Dust masses determined from (sub-) millimetre photometry show a clear distinction between the maximum masses of debris discs, of $\sim 10^{-1}\,M_\odot$, and typical protoplanetary disc dust masses of $\gtrsim 10^1\,M_\odot$, as shown in Fig. 1.14. During the age range of $\sim 5-30\,Myr$, the dust mass drops by at least two orders of magnitude. Dust masses in systems spanning this age range are studied in chapter 5 of this thesis.

**Figure 1.14:** Evolution of the dust mass of circumstellar discs, showing the transition from protoplanetary to debris discs between $\sim 5$ and $30\,Myr$. The masses are computed from $850\,\mu m$ photometry, all assuming $\kappa_\nu(850\,\mu m) = 0.17\,m^2kg^{-1}$. Taken from Wyatt [2008]
1.5. DUST EMISSION AND ABSORPTION

1.5.5 Fractional Luminosity

A parameter often used to characterise circumstellar discs is the ratio of the luminosity of the disc to the luminosity of the star,

\[ f = \frac{L_{\text{IR}}}{L_{\odot}}. \]  \hspace{1cm} (1.23)

This is the fraction of the starlight which is absorbed and re-radiated by dust. This doesn’t correspond to one particular quantity, as it depends not only on the total cross-sectional area of dust, but also strongly on the radial distribution. If the dust were distributed spherically around the star then \( f \) would be the optical depth of the dust to starlight \((f = 1 - e^{-\tau} \sim \tau)\).

For a disc, \( f \) is thus a lower limit on the optical depth of the disc to the starlight.

There is an empirical boundary at \( f = 10^{-2} \), which separates protoplanetary discs (above) from debris discs (below). This corresponds approximately to the point above which the optical depth of a disc to starlight starts to become significant.

1.5.6 Dust Opacities

As mentioned previously, the wavelength dependence of the opacity can be determined by fitting \( T_{\text{dust}} \), \( \beta \) and \( \lambda_0 \) to the observed flux density distribution or SED. However, to determine physical dust masses it is necessary to assume an absolute opacity at some wavelength. The uncertainty on the assumed opacity is typically at least a factor of two, so it is important that the assumed opacity is quoted alongside derived dust masses, to allow the masses to be scaled for comparison with other works. As (sub-)millimetre photometry is typically used for weighing discs, it is common to use an assumed opacity at either 850 \( \mu \text{m} \) or 1.3 mm.

For debris discs, the most commonly assumed opacity is \( \kappa(850 \mu\text{m}) = 0.17 \text{ m}^2/\text{kg} \) [Zuckerman and Becklin, 1993, Wyatt, 2008, and many other references]. This value has continued to be used, as studies since Zuckerman and Becklin [1993] have shown that \( \kappa(850 \mu\text{m}) \) could either be a factor of 3–5 less than this for spherical grains, or a factor of 3–5 higher for fractal grains [Najita and Williams, 2005, Beckwith et al., 1990, Pollack et al., 1994, Wright, 1987]. For consistency, \( \kappa(850 \mu\text{m}) = 0.17 \text{ m}^2/\text{kg} \) will be assumed for debris discs throughout this thesis.

For protoplanetary discs, an assumed opacity of \( \kappa(300 \mu\text{m}) = 1.0 \text{ m}^2/\text{kg} \) from Beckwith et al. [1990] is typically used\(^6\) [e.g., Andrews and Williams, 2005]. Assuming \( \beta = 1 \), the corresponding values at 850 \( \mu\text{m} \) and 1.3 mm are 0.35 and 0.23 \( \text{m}^2/\text{kg} \) respectively. Beckwith et al. [1990] note that this is a compromise, and the lowest and highest values in the literature are a factor of five either side of their adopted value. The value \( \kappa(850 \mu\text{m}) = 0.35 \text{ m}^2/\text{kg} \) is used to estimate the masses of protoplanetary discs around Herbig Ae/Be stars in chapter five of this thesis.

\(^6\)The actual value used in Beckwith et al. [1990] is a factor of 100 lower, but it is used with total disc masses, assuming a gas to dust ratio of 100.
1.6 Dust Production, Removal and Dynamics

A number of processes can act on dust grains in circumstellar discs in such a way as to destroy them, alter their trajectories from keplerian orbits, or remove them from the system. Here the dominant processes affecting grains in debris discs are described.

1.6.1 Radiation Pressure

Stellar radiation exerts a force on grains which opposes the gravitational force. The magnitude of both radiation and gravitational forces is inversely proportional to distance from the star, hence the ratio of these forces is independent of orbital radius for any particular grain. As the radiation force is proportional to a grain’s cross-sectional area, but the gravitational force is proportional to its volume, it can be seen that the radiation force becomes increasingly dominant as the grain size is reduced, and below some size grains can be blown out of the system.

For a simplified case of black spherical grains of radius \( a \) and density \( \rho \), at a distance \( r \) from the star, the ratio of radiation to gravitational force, \( \beta \), can be calculated from,

\[
F_{\text{rad}} = \frac{L_* a^2}{4 r^2 c} \quad \text{and} \quad F_{\text{grav}} = G \frac{4 \pi a^3 \rho M_*}{3 r^2},
\]

which gives [Backman and Paresce, 1993],

\[
\beta \equiv \frac{F_{\text{rad}}}{F_{\text{grav}}} = \left( \frac{3}{16 \pi c G} \right) \frac{L_* \, \mu m}{a \rho M_*} = 0.57 \left( \frac{1 \, \mu m}{a} \right) \left( \frac{1 \, \text{g cm}^{-3}}{\rho} \right) \left( \frac{L_*}{L_\odot} \right) \left( \frac{M_*}{M_\odot} \right).
\]  

(1.24)

Clearly, when \( \beta \) exceeds unity, grains are accelerated out of the system. More generally, the effect of radiation pressure on the orbits of grains is equivalent to reducing the mass of the star by a factor \((1 - \beta)\). For a grain released from a massive body (for which \( \beta \approx 0 \)) in a circular orbit, the eccentricity of the grain will be [Burns et al., 1979],

\[
e = \frac{\beta}{1 - \beta}
\]

(1.25)

Thus grains for which \( \beta < 0.5 \) remain in bound orbits with elevated eccentricities, and grains for which \( \beta > 0.5 \) leave the system on hyperbolic orbits in times comparable to orbital timescales (< \( 10^4 \) yr). These are both important outcomes for debris discs. Firstly, there is a grain size dependant smearing out of dust generated from massive bodies, with larger grains following orbits similar to the parent bodies, but smaller grains having increased eccentricities. Secondly, systems will be depleted of grains for which \( \beta > 0.5 \), as even if they are continuously produced they will be lost at a greater rate than larger grains for which \( \beta < 0.5 \).

In reality, computing \( \beta \) is more complicated than Eqn.(1.25) due to the wavelength depen-
dence of the effective cross-section of grains. This results in a maximum value of $\beta$ occurring for grain sizes approximately equal to the wavelength of the peak in the stellar SED. For the Sun, the maximum value of $\beta$ is $\sim 0.5$ when $a \sim 0.5 \mu m$ [Barns et al., 1979]. Thus for G-type and cooler main-sequence stars, radiation pressure does not cause grains to be blown out. For hotter stars, for which the stellar SED peaks at shorter wavelengths ($\lambda_{\text{max}} \propto 1/T_\star$), and $L_\star/M_\star$ is greater, grains below a “blowout” size are ejected. For A-type main-sequence stars, $T_\star L_\star/M_\star$ can be as much as 50 times that of the Sun, and the blowout size can exceed $10 \mu m$ [Backman and Paresce, 1993]. Discs around such stars must be devoid of dust grains smaller than the blowout size, unless they have been produced within the past $\sim 10^4$ years.

Evidence of the ejection, or eccentricity elevation, of small grains from the debris disc around an A-type star is seen in mid-IR images of Vega. These show a smooth cloud of dust extending out to $\sim 500$ AU [Su et al., 2005, Shibahara et al., 2010]. At sub-millimetre wavelengths, which are mostly sensitive to large dust grains ($a \gtrsim 100 \mu m$), a much more compact and clumpy dust distribution is seen [Holland et al., 1998].

![Image](image.png)

**Figure 1.15:** Vega imaged at 70 and 850 $\mu m$ showing effects of radiation pressure. Large grains, as seen in the 850 $\mu m$ image (right) are likely trapped in resonances with a planet. Smaller grains, as seen in the 70 $\mu m$ image (left) have highly eccentric or hyperbolic orbits which extend to several hundred AU. The images are at the same scale and are taken from Su et al. [2005] and [http://www.roe.ac.uk/ukatc/science/content/vega-disk.html](http://www.roe.ac.uk/ukatc/science/content/vega-disk.html).

### 1.6.2 Poynting-Robertson Drag

As well as exerting a radial force on dust grains, stellar radiation also exerts a tangential force which opposes their motion. This is the Poynting-Robertson (PR) effect, and causes grains to steadily spiral in towards the star. The effect is explained diagrammatically in Fig. 1.16.
Figure 1.16: Poynting-Robertson drag. (a) In rest frame of dust grain the radiation appears to come from in front of the grain (angle from radial direction: $\theta = v/c$ for $v \ll c$), so radiation pressure causes acceleration opposite to orbital velocity. (b) In the rest frame of stellar system the grain only feels radiation pressure radially, but radiation (either scattered or emitted) from the dust grain is blue shifted in the forward direction and red shifted in the backward direction, leading to a net loss of forward momentum. Image from http://en.wikipedia.org/wiki/File:Poynting-Robertson_effect.png.

It can be shown [Wyatt and Whipple, 1950, Burns et al., 1979, Backman and Paresce, 1993, Kalas, 2010] that the orbit decay time (orbital velocity divided by acceleration due to PR) for a grain of radius $a$ and absorption efficiency $Q_a$ is,

$$t_{PR} = \left( \frac{16\pi c^2}{3} \right) \frac{a \rho r^2}{Q_a L_*} \approx \frac{10^3}{Q_a 1 \mu m} \frac{a}{1 \mu m} \frac{\rho}{1 g cm^{-2}} \left( \frac{r}{1 A U} \right)^2 \frac{L_\odot}{L_*} \text{yr.}$$

For grains larger than the wavelength of the peak in the stellar SED ($a \gtrsim 1 \mu m$), $Q_a \approx 1$. The $r^2$ dependence means that PR drag seriously limits the lifetime of dust within the inner regions of systems. Thus PR drag dictates that any dust observed in the inner regions of systems must either be transient or continuously replenished. In addition, the dependence on $a$ means that the dust population in the inner regions of systems will tend to be dominated by larger grains which are depleted less rapidly.

Dust created in planetesimal belts which are exterior to the orbits of planets will spiral inward through the orbits of the planets. This dust can either be accreted onto the planets, dynamically ejected, or trapped, at least temporarily, in gravitational resonances with the planets (in analogy to the Trojan asteroids in Jupiter’s orbit). Over-densities of dust caught at such resonant points are thought to be the cause of the clumpy emission seen in sub-millimetre images of some debris discs [e.g., Wyatt, 2006].

Noting that the temperature, $T_g$, of a grain in thermal equilibrium is proportional to
(L*/r^2)^{1/4}$, the PR timescale can be re-written as,

$$t_{PR} \approx 10^3 \frac{a}{1 \mu m} \frac{\rho}{1 g \ cm^{-3}} \left( \frac{278 K}{T_g} \right)^4 \text{yr}. \quad (1.28)$$

This is useful to determine approximate PR timescales for dust belts with temperatures fitted from observed SEDs.

### 1.6.3 Collisions

Bodies of all sizes, from the smallest dust grains to large planetesimals, can be broken down by collisions. The time scale for collisions between bodies is proportional to the orbital period, $P$, and inversely proportional to the fractional surface density of the disc, $\sigma(r)$ (effectively the face-on optical depth). Specifically, it can be shown that [Backman and Paresce, 1993, Kalas, 2010],

$$t_{\text{coll}} \geq \frac{P}{8\sigma(r)} \sim \frac{1}{8\sigma(r)} \left( \frac{r}{1 \text{AU}} \right)^{3/2} \left( \frac{M_\star}{M_\odot} \right)^{-1/2} \text{yr}, \quad (1.29)$$

where $\sigma(r)$ is the fractional surface density at radius $r$.

The outcome of a collision on a body is determined by the specific kinetic energy, $Q$, which depends on the size of the body, $a$, the size of the impactor, $a_{im}$, and the collision velocity, $v_{\text{coll}}$ [Wyatt and Dent, 2002].

$$Q = \frac{1}{2} \left( \frac{a_{im}}{a} \right)^3 v^2_{\text{coll}}. \quad (1.30)$$

If $Q$ is especially low, a collision can result in an inelastic rebound, or the bodies sticking together. This, however, is considered to rarely be the case in debris discs [Wyatt and Dent, 2002]. For higher specific energies, either cratering will occur, in which less than half the mass of a body is ejected, or for still higher energies, a body will shatter entirely into pieces having less than half the original body’s mass. A typical threshold for shattering is $\sim 10^4$ J/kg, which for colliding bodies of comparable sizes equates to a collision velocity of $\sim 100$ m/s. The orbital smearing effect of radiation pressure on small grains is sufficient to create relative velocities greater than this threshold for grains of 1–10 times the blowout size, regardless of random vertical motions [Backman and Paresce, 1993]. Thus, for small grains, collisions are almost always destructive. This acts to remove dust from systems over time by producing fragments which are below the blowout size, or spiral inward by PR drag.

As $\sigma$ is generally proportional to the mass, and hence infrared luminosity of a disc, the collisional timescale within planetesimal belts of detected discs is typically much shorter than the PR timescale. In clear inner regions of discs, however, the PR timescale can dominate, and the effects of collisions can be negligible.
1.6.4 Sublimation

Grains composed largely of ices can be destroyed by the sublimation of the ices. Sublimation of ices is also the cause of cometary activity, which is an important source of dust in the inner regions of planetary systems. The rate of sublimation for each type of ice is a very strong function of temperature, and is also dependant on heating rate [e.g. Brown and Bolin, 2007] and on the surrounding substrate material [e.g. Fraser et al., 2001, Mukai, 1996]. The sublimation rate, \( \gamma \), as a function of heliocentric distance in the Solar System (a proxy for temperature - a relationship of \( T = 255r^{-0.35} \) K is assumed [Mukai, 1996]) for four important molecules are shown in Fig. 1.17.

![Graph showing sublimation rates of ices as a function of heliocentric distance.](image)

**Figure 1.17:** Sublimation rates of ices as a function of heliocentric distance. Taken from Mukai [1996].

The timescale for an ice grain of mass \( m \), surface area \( A \), density \( \rho \) and radius \( a \), to completely sublimate is [Mukai, 1996],

\[
\tau_{\text{subl}} = m/\gamma A = \rho a/3\gamma \sim 10^{-8} \left( \frac{a}{1 \mu\text{m}} \right) \left( \frac{1 \text{ kg m}^{-2}\text{s}^{-1}}{\gamma} \right) \text{ yr.} \quad (1.31)
\]

Thus with reference to Fig. 1.17, the lifetime of grains with significant fractions of water and ammonia ices are limited to years or less within approximately \( \sim 4 \) and \( \sim 10 \) AU of the Sun respectively, although there is considerable variation due to differing grain emissivities (Eqn. (1.11)). The distance at which water ice sublimates rapidly (although not clearly defined) is often referred to as the ‘frost line’ or ‘snow line’. Icy grains migrating inward due to PR drag will be destroyed as they traverse the frost line, and grains found within the frost line must be entirely rocky. For other stars these sublimation distances scale approximately as \( L^1/4 \) (Eqn. (1.13)).
1.6.5 Stellar Wind Pressure

Stellar winds can exert a force radially outward on grains in the same way as radiation pressure. For an isotropic radial wind, the force on a spherical grain of radius $a$ at distance $r$, is,

$$F_{\text{wind}} = \frac{\dot{M}_* v_{\text{wind}} a^2}{4r^2},$$

(1.32)

where $\dot{M}_*$ is the mass loss rate of the star, and $v_{\text{wind}}$ is the speed of the wind at the distance of the grain. This has the same $a$ and $r$ dependence as the force due to radiation pressure in Eqn. (1.24). The combined effect can be seen by redefining $\beta$ in Eqn. (1.25) to include stellar wind pressure,

$$\beta = \frac{F_{\text{rad}} + F_{\text{wind}}}{F_{\text{grav}}} = \left(\frac{3}{16\pi G}\right) \frac{1}{a p M_*} \left(\frac{L_*}{c} + \dot{M}_* v_{\text{wind}}\right).$$

(1.33)

For the Sun, $\dot{M}_* \sim 2 \times 10^{-14} M_\odot/\text{yr} \sim 1 \times 10^9 \text{kg/s}$, $v_{\text{wind}} \sim 10^5$–$10^6 \text{m/s}$, and $\frac{L_*}{c}/\dot{M}_* v_{\text{wind}} \sim 100$. This situation is similar for K to F-type stars, with radiation pressure dominating. For A-type stars the convection which drives stellar winds becomes negligible, and as a result radiation pressure dominates completely. However, for M-type stars, which are generally highly active, it is thought that mass loss rates could be up to 1000 times the current solar value [Plavchan et al., 2005]. This, combined with luminosities of $<0.1 L_\odot$, means that the stellar wind can dominate, and plays a similar role to radiation pressure in discs around earlier-type stars. Evidence for this is seen in the debris disc of AU Microscopii, which is a star of spectral type M1 V in the $\sim 10 \text{Myr}$ old $\beta$ Pic association. Optical and near-IR scattered light images reveal dust extending to radii of $\sim 300 \text{AU}$ [Kalas, 2010], but modeling of the surface brightness profile in these images, and the infrared and sub-millimetre SED, implies that the majority of the large grains ($a \gtrsim 1 \text{mm}$) are contained in a narrow ‘birth ring’ at approximately 43 AU [Strubbe and Chiang, 2006]. The dust outside this ring is thought to be in elliptical orbits due to stellar wind pressure, in analogy to the radiation pressure driven grains seen around the $\sim 1000$ times more luminous A-type stars $\beta$ Pictoris and Vega (Fig. 1.15).

1.6.6 Stellar Wind Drag

Stellar winds also cause a drag force analogous to the PR drag caused by starlight. It can be shown that [Plavchan et al., 2005],

$$\frac{F_{\text{wind, drag}}}{F_{\text{wind, rad}}} \approx \frac{c}{v_{\text{wind}}} \frac{F_{\text{PR}}}{F_{\text{rad}}}. $$

(1.34)

For typical wind speeds, $c/v_{\text{wind}}$ is a factor of $\sim 100$–$1000$. For the Solar System the result is that the drag force due to the solar wind is approximately a third of the PR drag force. For
more active or less luminous stars, the wind drag dominates, and orbit decay timescales can be much shorter than predicted from PR drag alone [Plavchan et al., 2005],

\[
\frac{t_{\text{wind, drag}}}{t_{\text{PR}}} \approx \frac{L_*}{M_\star c^2}
\]  

(1.35)
1.7 Thesis Outline

1. **Introduction.** Background on relevant aspects of circumstellar disc evolution, pre-main sequence stellar evolution and associations of young stars; properties of thermal emission; emission and absorption by dust, including relationships between observed flux density, dust mass and dust temperature, and dust temperature and orbital distance; and the physical processes which act on dust in discs (focused on debris discs).

2. **Debris Disc Surveys: SUNS and DEBRIS.** This thesis primarily focuses on surveys of nearby stars in search of thermal emission from debris discs (chapter five is the only exception). This chapter provides an overview of previous surveys and their limitations. The need for large unbiased surveys using far-IR and sub-millimetre wavelengths is demonstrated, and two such surveys are outlined – the DEBRIS Herschel Key Programme and the SCUBA-2 Unbiased Nearby Stars (SUNS) survey.

3. **Target Selection for the SUNS and DEBRIS Surveys.** These debris disc surveys aim to determine statistically what system properties (e.g. stellar mass, metallicity, age, binarity, presence of planets) lead to the presence and properties of debris discs. To achieve this, and to maximise legacy value, the targets needed to be drawn from a clearly defined and maximally complete sample. This chapter describes the selection criteria and the work involved in the selection process. The catalogue of target systems and distributions of properties within the sample are presented.

4. **Properties of Nearby A Type Stars with Spitzer.** This chapter presents a survey of a volume limited sample of 130 A type star systems (primary star has spectral type of A) at 24 and 70 µm using the Spitzer space telescope. Previous studies have shown that such observations detect debris discs around ∼30% of A type systems. The survey presented here is the least biased and most homogeneous (in terms of data reduction, calibration and analysis) study performed to date. Dust temperatures, orbital distances and masses are computed for all detected discs. Debris detection rates are compared as a function of stellar effective temperature, metallicity, binarity and binary separation. These results are compared to other works, and two particularly interesting discs discovered in this work are discussed.

5. **LABOCA Data Reduction & Observations of Southern Circumstellar Discs.** This chapter presents the results, data reduction and analysis from a programme to measure the masses of discs around stars with ages of 5–30 Myr using 870 µm observations. This age range encompasses the final stages of protoplanetary disc evolution and the initial
stages of debris disc evolution. The targets are all part of a large \textit{Herschel} Key Programme studying gas in circumstellar discs. The mass measurements and upper limits are discussed in the context of rocky planet formation and gas giant core formation timescales. The photometry obtained for four Herbig Ae/Be stars is combined with literature millimetre photometry to examine the disc geometries and optical depths.

6. \textbf{Sub-millimetre Study of \textit{\varepsilon} Indi and \textit{\alpha} Centauri with LABOCA.} Two of the Sun’s nearest G/K spectral type neighbours were observed at 870 \textmu m in search of cold dust not previously detected at $\leq 70$ \textmu m. This work was motivated by sub-millimetre images of debris discs around similar stars such as \textit{\varepsilon} Eridani and \textit{\tau} Ceti. Observations of \textit{\alpha} Centauri at two epochs separated by two years are used to allow background emission to be distinguished from emission associated with the system, utilising the system’s large proper motion. Dust mass upper limits are determined for the components in the \textit{\varepsilon} Indi system, and a bright feature detected in the \textit{\alpha} Centauri system is discussed.

7. \textbf{Conclusions & Future of the Field.} The thesis concludes with a summary of the results and conclusions from each chapter. The science drivers for future circumstellar disc surveys and the potential of future observing facilities in this field are described, and required development of techniques and ancillary supporting data are highlighted.
Chapter 2

Debris Disc Surveys: SUNS and DEBRIS

This chapter presents the current state of observational surveys, both past and present, for thermal emission from debris discs. The most fruitful searches for debris discs have so far been performed in the mid/far-IR, with the majority of detected discs resembling massive analogues of the Kuiper belt in the Solar System. A brief history of surveys performed with IRAS, ISO, Spitzer and ground-based sub-millimetre telescopes will be given, including pertinent scientific conclusions and limitations. It will be seen that from the results of previous surveys it is difficult at best to draw statistical conclusions on what system properties lead to the presence and properties of debris discs. This is due to biases in the selection of targets (many surveys have, understandably, targeted stars thought to have a high probability of harbouring debris discs) and biases in the detectable range of dust temperatures due to differing wavelength coverages.

Determining the fraction of systems with debris discs, and the properties of discs, has important implications for understanding planet formation and the evolution of planetary systems. The detection of debris in a system implies that planet formation processes had at least proceeded as far as producing planetesimals, the collisions between which generate the observed dust. Although the present day levels of dust in the Kuiper Belt are below the detection limits of surveys of other stars, it is thought that in its early stages the Solar System contained a much larger mass of planetesimals, with the mass of dust being especially elevated during the Late Heavy Bombardment phase (LHB). It is likely that some detected debris discs are undergoing such a LHB phase, and determining how common these scenarios are has important implications for placing the evolution of the Solar System in context. The ‘dustiness’ of systems also
has important implications for the development of life in planetary systems.

New sub-millimetre instruments offering vastly improved sensitivity over their predecessors, and the recently successfully launched and commissioned *Herschel* Space Observatory, are allowing large surveys of stars for debris discs in the wavelength range 70–850 μm. The author has been involved in the planning and preparatory work for two such surveys, which are discussed at the end of this chapter. These surveys are designed to provide a census of the debris disc incidence and properties around nearby A–M spectral type main sequence field stars. By utilising observations from optical to sub-millimetre wavelengths for a clearly chosen, minimally biased, sample of stars, it will be possible to provide robust statistical constraints on the properties and causes of debris discs. The target selection for these surveys is the subject of chapter 3, and a precursor survey of A-type stars with *Spitzer* is presented in chapter 4.
2.1 Modelling Survey Detection limits and Selection Effects

Interpreting the statistical results from debris disc surveys requires selection effects to be understood. The key results from surveys are typically how the detection rates of debris discs (i.e. how likely it is to detect a debris disc) vary with host star parameters. One of the trends typically seen is a variation of debris detection rate with stellar spectral type (effective temperature and luminosity): the detection rate increases from M through A spectral types. These results will be described in more detail in the following sections. To understand whether these trends are a selection effect, and to explore the relative abilities of different instruments to detect discs around stars of different types, a model is presented here to predict the number of stars of different types around which dust belts of given radius and mass can be detected. This is achieved by determining the maximum distance at which a disc would be detectable as a function of stellar effective temperature, and comparing these with the space densities of stars of different spectral types.

The minimum dust flux detectable for an instrument is modelled as a multiple, $n$, of the quadrature sum of the RMS noise, $\sigma_{\text{noise}}$, and a fraction, $f$, of the stellar photosphere flux, $F_*$, to represent flux calibration uncertainty:

$$F_{\text{dust, min}} = n \sqrt{\sigma_{\text{noise}}^2 + (fF_*)^2}$$  \hspace{1cm} (2.1)

The calibration uncertainty should include the photosphere flux prediction uncertainty, of typically $\sim 2-5\%$ (see e.g. chapter 4 for detailed discussion), when significant in comparison to the instrumental calibration uncertainty. Generally the instrumental calibration uncertainty dominates, with typical 1σ values of 5-15% for most mid-IR to millimetre wavelength instruments. *Spitzer/MIPS-24* is a notable exception, with instrumental calibration uncertainty of only 1σ $\sim 1\%$.

The maximum detectable distance for a mass $M_{\text{dust}}$ of dust, with temperature $T_{\text{dust}}$ and opacity $\kappa(\lambda)$, is found by substituting for $F_{\text{dust, min}}$ the flux from eqn. (1.17). The stellar flux can be approximated to sufficient accuracy for these purposes (within a factor of two at mid-IR and longer wavelengths) by assuming the star is a spherical black body with temperature $T_{\text{eff}}$ and luminosity $L_*$. Substituting for $F_{\text{disc, min}}$ and $F_*$ in eqn. (2.1) and rearranging then yields,

$$d_{\text{max}} = \frac{1}{n\sigma_{\text{noise}}} \left[ \left( \kappa(\lambda)M_{\text{dust}}B_\nu(\lambda,T_{\text{dust}}) \right)^{2} - \left( n f B_\nu(\lambda,T_{\text{eff}})\frac{L_*}{4\pi\sigma_{SB}T_{\text{eff}}^4} \right)^2 \right]^{1/4}$$  \hspace{1cm} (2.2)

The situation where the term in square brackets becomes negative corresponds to the fractional excess flux due to the dust being smaller than $n$ times the calibration uncertainty – i.e. the
disc would be undetectable even with infinite signal to noise ratio.

The opacity is determined using eqn. (1.19), with a reference opacity of \( \kappa(850\,\mu m) = 0.17\, m^2/kg \), exponent of \( \beta = 1.0 \), and critical wavelength of \( \lambda_0 = 200\, \mu m \) assumed, i.e.,

\[
\kappa(\lambda) = 0.17\, m^2/kg \times \begin{cases} 
\left( \frac{200\, \mu m}{850\, \mu m} \right)^{-1.0} & \lambda \leq 200\, \mu m \\
\left( \frac{\lambda}{850\, \mu m} \right)^{-1.0} & \lambda > 200\, \mu m 
\end{cases}
\] (2.3)

These values are typical for debris discs [e.g. Wyatt, 2008]. The dust temperature is determined from the orbital radius and stellar luminosity using eqn. (1.14), assuming amorphous grains of size \( a = 100\, \mu m \). Where this temperature would be below the temperature for black body grains (eqn. (1.13)), the black body grain temperature is used. The stellar luminosity is determined as a function of stellar effective temperature using an empirical relationship for main sequence stars determined in chapter 3 (§3.5.3, eqn. (3.8)).

In the following sections plots of \( d_{\text{max}} \) as a function of stellar temperature are used to show the abilities of photometric bands of various instruments to detect massive Asteroid and Kuiper belt analogues (Figs. 2.2, 2.3, 2.4, 2.8 for IRAS, ground based (sub-)mm facilities, Spitzer/MIPS, and Herschel respectively). These belts have radii of 2.8 and 45 AU respectively, with dust masses selected to yield meaningful detection distances (\( \gtrsim 10\, \text{pc} \) for at least some \( T_{\text{eff}} \) range, in at least one photometric band). To give an idea of the number of stars around which these discs could be detected as a function of \( T_{\text{eff}} \), stars from five maximally complete volume-limited samples of \( \sim 125\, \text{M}, \text{K, G, F and A type main sequence stars, compiled in chapter 3, are overplotted. The model debris discs would be detectable around stars which fall below the } d_{\text{max}} \text{ curve for an instrumental band. The detection significance, } n, \text{ is set to 3 in all cases} \) i.e. \( d_{\text{max}} \) is the maximum distance at which a 3\( \sigma \) detection of the dust would be obtained.
2.2 Previous Surveys

2.2.1 IRAS

The Infrared Astronomical Satellite [IRAS; Clegg, 1980], with a 60 cm liquid helium cooled primary mirror, performed an all sky survey in four bands centred at 12, 25, 60 and 100 μm during a ten month mission in 1983. In total 96% of the sky was covered by scanning the sky from a near polar orbit with a 103 minute period. The optical resolution was diffraction limited for the 25–100 μm bands. IRAS was primarily intended for producing catalogues of sources rather than high fidelity imaging – the detector beams are elongated in the cross-scan direction (mostly ~ 4–5′) and differ between detectors, and the spatial sampling varies over the sky. Products from the IRAS mission include catalogues of point sources and extended sources, low resolution (1.5′/pix, 4′ FWHM) images\(^1\), and tools for analysing the scan data\(^2\) and reconstructing higher resolution images\(^3\). The catalogues of point sources have been used extensively in studies of debris discs, and they are used in chapter four of this thesis.

Catalogues of Point Sources

The IRAS Point Source Catalogue [PSC; Beichman et al., 1988] contains approximately 246,000 high reliability detections from the scan data which are unresolved, covering all of the sky which was scanned at least twice (96%). Positional uncertainties are specified by an error ellipse, with median 1σ major and minor axis values of 18″ and 4″.

The IRAS Faint Source Catalogue [FSC; Moshir et al., 1990] was released later and was produced by point-source filtering and co-adding the scan data, resulting in a typical increase in sensitivity of one magnitude (factor of 2.5) over the PSC. Median 1σ major and minor axis positional uncertainties are 20″ and 4′′ (14″ and 2′′ for sources also in the PSC). Median 1σ photometric uncertainties for detections are 11%, 9%, 15% and 18% for 12 to 100 μm respectively. As the analysis used to produce the FSC offered little benefit in highly confused areas, the galactic plane (|b| < 10°) and a few other confused areas are omitted from the FSC. The FSC contains only approximately 173,000 sources as a result. The spatial distribution of PSC and FSC entries are shown in Fig. 2.1 (top). The Vega magnitude distributions of the detections in the four bands in the FSC are shown in Fig. 2.1 (bottom). The magnitudes at which incompleteness starts to become noticeable are approximately 5.6\(^m\) (0.23 Jy), 3.8\(^m\) (0.28 Jy), 1.7\(^m\), (0.32 Jy) and −1.1\(^m\) (1.27 Jy) at 12 to 100 μm respectively. The limiting magnitudes are approximately 1\(^m\) fainter than these.

\(^1\)most recently, IRIS: Miville-Deschênes and Lagache [2005].
\(^2\)e.g., SCANPI: http://irsa.ipac.caltech.edu/IRASdocs/scanpi/
\(^3\)e.g., HIRES: http://irsa.ipac.caltech.edu/IRASdocs/hires_over.html
Figure 2.1: Top: sky distribution of IRAS PSC and FSC entries. Bottom: IRAS FSC magnitude distributions (0.1 μm bins). Black lines: values with quality two (low quality detections) or three (high quality detections). Red lines: values with quality three only (there are none at 100 μm). Approximate magnitudes at which the gradient turns over (indicating the onset of significant incompleteness), shown by green vertical lines, are: 5.6 (0.23 Jy), 3.8 (0.28 Jy), 1.7 (0.32 Jy) and −1.1 (1.27 Jy).
Debris Disc Surveys with *IRAS*

Since IRAS made the discovery of the first debris disc, around Vega, during early calibration observations [Aumann et al., 1984], many authors have mined the *IRAS* catalogues and scan data in search of other stars exhibiting excess IR emission over their photospheres [Aumann, 1985, Odenwald, 1986, Sadakane and Nishida, 1986, Cote, 1987, Aumann, 1988, Walker and Wolstencroft, 1988, Aumann and Probst, 1991, Patten and Willson, 1991, Stencel and Backman, 1991, Oudmaijer et al., 1992, Cheng et al., 1992, Zuckerman et al., 1995, Mannings and Barlow, 1998, Plets and Vynckier, 1999, Fajardo-Acosta et al., 2000, Moór et al., 2006, Rhee et al., 2007]. These works typically joined catalogues of nearby or bright stars with the *IRAS* catalogues using astrometry, often requiring the matched *IRAS* entries to have detections in at least the 12 and 25 μm bands. The *IRAS* colours ([12] − [25], [12] − [60] or [25] − [60]), which would nominally be zero for stellar photospheres, would be examined for a significantly positive (red) value. Due to the uncertainties in the *IRAS* photometry, of typically 1σ ∼ 10–15% (0.1–0.15 m), colour thresholds between 0.3 m and 1.6 m were typically used. The 100 μm *IRAS* band has typically not been used for the detection of excess due to its low sensitivity compared to the 60 μm band, and the significant complications due to confusion from galactic cirrus at 100 μm (even at high galactic latitudes).

Useful reviews of works up to 1999, and statistical analysis of the incidence rates of debris discs, are given in Mannings and Barlow [1998] and Plets and Vynckier [1999]. These showed that a total of approximately 180 debris discs around main-sequence stars were detectable by *IRAS* over the whole sky (88 in approximately half the sky in Mannings and Barlow [1998]), that the debris disc hosts were primarily of late B to early G spectral type, and that the incidence rate of debris for stars of these spectral types was approximately 15%.

More than 20 years on from the *IRAS* mission, the *IRAS* catalogues are still yielding new results on debris discs. Moór et al. [2006] performed a systematic study of IRAS detected discs of high fractional luminosity ($f > 10^{-4}$), and showed that almost all stars with $f > 5 \times 10^{-4}$ are in stellar moving groups with ages less than 100 Myr. Rhee et al. [2007] performed yet another debris disc search, joining the *Hipparcos* catalogue [Perryman et al., 1997] with the *IRAS* FSC and PSC, taking several steps to avoid false detections. This study yielded a total of 146 main sequence stars with high reliability 60 μm excess emission within 120 pc of the Sun, including 33 previously unidentified systems. The spectral type distribution of debris disc hosts found by Rhee et al. [2007] is in agreement with that of Mannings and Barlow [1998] – the detections are dominated by stars of late B to early G spectral type.

The low, and variable, spatial resolution of IRAS (∼1'), leads to the significant possibility of false excess detections due to confusion of sources. Authors have typically taken steps to
minimise this by, for instance, ignoring IRAS detections which matched multiple stars in the
catalogues they were matching with. However, spurious excess detections due to confusion with
uncatalogued objects [e.g., Lisse et al., 2002], or non-circumstellar dust [e.g., Kalas et al., 2002,
Gaspár et al., 2008], remain a significant problem which can only be remedied by follow-up
observations at higher spatial resolution.

Detection Limits and Selection Effects

The dust masses of debris discs detected by IRAS span $10^{-3}$ to $1M_\oplus$ [Rhee et al., 2007], which
is to be compared with the present Kuiper Belt dust mass of $\sim 10^{-6} - 10^{-5}M_\oplus$ [e.g., Moro-
Martín and Malhotra, 2003]. The spectral type distribution of IRAS detected debris disc hosts
is dominated by stars of late B to early G spectral type, however, it can be shown that this
is a selection effect. Figure 2.2 shows the maximum distance at which a massive Kuiper Belt
analogue ($r = 45$ AU) of $M_{\text{dust}} = 10^{-2}M_\oplus$, and a massive Asteroid Belt analogue ($r = 2.8$ AU)
of $M_{\text{dust}} = 10^{-4}M_\oplus$, can be reliably detected by the four IRAS bands as a function of stellar
effective temperature. These plots were produced using the model described in the previous
section, and both the flux sensitivity of each band (values from Fig. 2.1) and a $3\sigma = 0.6^{\mu}$
calibration uncertainty are accounted for. For comparison, the five volume limited spectral
type subsamples of Phillips et al. [2010] (chapter 3) are over-plotted.

![Kuiper Belt analogue (r=45AU, M=10^{-2}M_\oplus)](image1)

![Asteroid Belt analogue (r=2.8AU, M=10^{-4}M_\oplus)](image2)

**Figure 2.2:** IRAS distance limits for Kuiper and Asteroid belt analogues with dust masses of $10^{-2}$ and
$10^{-4}M_\oplus$ respectively, as a function of stellar spectral type (effective temperature). Points represent the
closest $\sim 125 M$ (red), K (orange), G (green), F (cyan) and A (blue) spectral type main sequence stars
to the Sun (not including secondary stars in multiple systems). The model belts would be detectable
for stars below the $d_{\text{max}}$ curves. This shows that, for instance, IRAS could detect such belts around
$\sim 100 A$ type stars, whilst not being able to detect any around $M$ type stars.
2.2. PREVIOUS SURVEYS

It can readily be seen that the spectral type distribution of stars below the distance limits closely matches the distributions of debris disc hosts in Rhee et al. [2007] and Mannings and Barlow [1998]. This indicates that there is no reason to infer a strong trend of debris disc incidence rate with spectral type based on the IRAS detections.

2.2.2 ISO

The European Space Agency’s Infrared Space Observatory [ISO; Kessler et al., 1996], like IRAS, had a 60 cm liquid helium cooled primary mirror. ISO operated from the end of 1995 to April 1998, performing pointed observations. Four instruments were available: a short wavelength (2.5–17 μm) camera [ISOCAM; Cesarsky et al., 1996], an imaging photopolarimeter with bands between 3 and 200 μm [ISOPHOT; Lemke et al., 1996], and two spectrometers [SWS and LWS; de Graauw et al., 1998, Swinyard et al., 1998]. The key advantages over IRAS for photometry and imaging were pixel sizes matched to the optical resolution (diffraction limited down to ~5 μm), and the ability to detect far fainter sources due to more sensitive detectors and the longer integration times permitted by the pointed nature of the mission. Unfortunately, a combination of unexpected detector transient response, and confusion, precluded the use of efficient chopped point-source photometry with ISOPHOT. Debris disc surveys with ISOPHOT had to switch to more time consuming raster observations, with a reduction in the number of targets or the number of bands [e.g., Habing et al., 2001].

A small number of surveys of a significant number of stars (9–150 each), searching for debris discs, were performed with ISOPHOT [Abraham et al., 1998, Habing et al., 1999, 2001, Decin et al., 2000, Spangler et al., 2001, Laureijs et al., 2002]. Reviews are presented in Decin et al. [2003] and Ábrahám et al. [2003]. In general these surveys were driven by photometry at 60 μm, with 44'' square detector pixels and a typical 1σ noise of 21 mJy (approximately a factor of five deeper than IRAS). Relatively few new unambiguous statistical results were achieved from the ISOPHOT surveys. The typical debris detection rate of ~15% [Decin et al., 2003] is the same as found by IRAS. As with IRAS, few warm debris discs were detected by using shorter wavelength (primarily 25 μm) photometry [Laureijs et al., 2002]. It was shown that debris discs around young stars (≤400 Myr) are typically more common and more massive than around older stars [Habing et al., 1999, Spangler et al., 2001, Decin et al., 2003], however, a clear relationship of detection rate or mass as a function of age could not be determined due to incompleteness, small sample sizes and statistical biases [Ábrahám et al., 2003, Decin et al., 2003].

The legacy of the ISOPHOT debris disc surveys is primarily the provision of high quality photometry (deeper, better calibrated and less confused than IRAS) at 60–170 μm for bright discs, and the ‘cleaning’ of the IRAS debris disc samples by providing higher resolution 60 μm
images than IRAS, which can show confused detections [e.g., Moór et al., 2006, Rhee et al., 2007]. The outcomes of the ISO debris disc surveys highlighted the need for large (>200 stars) surveys, to an even deeper depth, and with efficient observations at wavelengths longer than 60 μm, to make significant advances in understanding the factors affecting the incidence and properties of debris discs.

2.2.3 Ground Based (Sub-)Millimetre Observations

The low spatial resolution, and correspondingly poor confusion limited sensitivity, of IRAS and ISO at their longer wavelengths (100–200 μm) made them most productive at detecting dust emission at 60 μm (Fig. 2.2). The temperature of dust which has its peak flux density at 60 μm is ~80 K. For lower dust temperatures the dust mass sensitivity of IRAS and ISO drops rapidly. Dust in the Kuiper Belt in the Solar System has a temperature of ~50 K, and resolved scattered light images of debris discs have shown dust at much greater distances from their star than the Kuiper Belt [e.g., Smith and Terrile, 1984]. There is thus strong motivation for observations at longer wavelengths than 60 μm, with better spatial resolution than IRAS and ISO, to detect cold dust. In the late 1990s, the advent of sensitive bolometer arrays operating in the sub-millimetre (sub-mm) and millimetre (mm) atmospheric windows (450 μm, 850 μm and 1.3 mm) [e.g., Kreysa et al., 1998, Holland et al., 1999] on large ground-based telescopes (10–30 m), allowed the possibility of reliably detecting such cold dust for the first time.

Observing dust at these long wavelengths is challenging, as the flux density distribution of debris discs is typically dropping as \( \lambda^{-3} (\beta = 1) \), whilst the atmospheric transmission gets progressively worse for the shorter wavelength windows (450 μm and shorter windows are only accessible from the very driest sites for a small fraction of time). The sensitivity at these wavelengths is ultimately limited by confusion from extragalactic point sources (dusty high-z galaxies), although the integration times required to reach this limit are often impractically long.

The first instrument to provide sufficient sensitivity was the Sub-millimetre Common User Bolometer Array [SCUBA; Holland et al., 1999], operating on the 15 m James Clerk Maxwell Telescope (JCMT\(^4\)) on Mauna Kea, Hawaii. SCUBA was able to perform extragalactic confusion limited imaging at 850 μm (1σ = 0.7 mJy/beam – a 3σ limiting Vega magnitude of 1.0\(^\text{m}\)), although this required several hours of integration time. Observations to a similar depth at 450 μm were possible, but typically required 20 hours of the very best weather at Mauna Kea. Initial studies concentrated on the most conspicuous discs detected by IRAS [e.g., Vega, β Pic, Fomalhaut and ε Eri; Holland et al., 1998, Greaves et al., 1998]. As these targets are close to the

\(^4\)http://www.jach.hawaii.edu/JCMT/
Sun, their discs were spatially resolved in the 15" FWHM JCMT 850 μm beam. Modelling of these images [e.g., Wyatt and Dent, 2002, Wyatt, 2006] has shown the wealth of information, on both debris and planets with which debris interacts, which can be extracted from resolved sub-mm images. Several other discs detected in the mid/far-IR have been followed up by (sub-)mm imaging [e.g., Greaves et al., 2004b, Marsh et al., 2006, Corder et al., 2009].

Several surveys of 7–32 stars have been performed [Holmes et al., 2003, Wyatt et al., 2003, Greaves et al., 2004a, Liu et al., 2004, Najita and Williams, 2005, Lestrade et al., 2006, 2009, Nilsson et al., 2009, 2010]. These have necessarily had to focus on specific groups of stars: previously detected discs [Holmes et al., 2003, Nilsson et al., 2009, 2010], stars with known planets [Greaves et al., 2004a], solar analogues [Najita and Williams, 2005], young stars [Wyatt et al., 2003, Liu et al., 2004, Lestrade et al., 2006, Nilsson et al., 2009] and late-type (K/M) stars [Lestrade et al., 2006, 2009].

The ability of (sub-)mm observations to detect cold dust has been especially beneficial for detecting dust around low luminosity stars (K/M-type dwarfs). Fig. 2.3 shows distance limits for deep (sub-)mm observations (3σ detections, with σ = 5, 1 and 0.3 mJy at 450 μm, 850 μm and 1.3 mm) for the same star and disc models as in Fig. 2.2. Whilst IRAS was insensitive to dust at Kuiper Belt like distances around mid-G and later stars, the number of stars within the detection distance limits of (sub-)mm observations are approximately equal for all spectral types to late-M. Such detected discs around late-type stars are AU Mic, GJ 182 and GJ 842.2 [Liu et al., 2004, Lestrade et al., 2006].

Whilst ground-based (sub-)mm observations have the potential to offer significant advances in the detection of discs around late-type stars, the detection of dust at large distances from earlier type stars, and to constrain the dust properties (masses, temperatures and grain sizes) of discs detected at other wavelengths; the integration times of several hours per star have made it unfeasible to perform large surveys from which statistical conclusions may be drawn. The primary impact of (sub-)mm observations on the field to date has been the photometry and resolved images of bright mid/far-IR detected discs, and the detection of a small number of discs around young late-type stars.

2.2.4 Spitzer

The Spitzer space telescope [Werner et al., 2004] can be seen as an evolution of the ISO concept. The primary mirror was helium cooled, with a diameter of 85 cm. Three instruments were available: a 4-band near-IR camera [IRAC; 3.6, 4.5, 5.8 and 8 μm; Fazio et al., 2004], a 3-band mid/far-IR camera [MIPS; 24, 70 and 160 μm; Rieke et al., 2004b,a], and a mid-IR spectrometer [IRS; 5–38 μm; Houck et al., 2004]. Spitzer performed routine cryogenic operations between
Figure 2.3: (Sub-)Millimetre distance limits for Kuiper and Asteroid belt analogues with dust masses of $10^{-2}$ and $10^{-4} M_\odot$ respectively (as in Fig. 2.2 for *IRAS*), as a function of stellar spectral type (effective temperature). These assume typical 3σ detections in deep ground based observations ($\sigma = 5, 1$ and 0.3 mJy at 450 µm, 850 µm and 1.3 mm respectively). This shows that, by coincidence, the number of stars for which such belts are detectable within the distance limits are essentially independent of spectral type. This lack of stellar bias is one of the key advantages of (sub-)mm debris disc surveys.

January 2004 and April 2009, and is currently executing a warm programme using only the two shortest wavelength IRAC bands. Debris disc surveys have primarily used the 24 and 70 µm MIPS bands and IRS. The 160 µm MIPS band has only been used to observe a small number of discs due to poor sensitivity compared to the other bands, a near-IR light leak (especially problematic for nearby stars), and confusion.

The key advantages of *Spitzer* /MIPS over *ISO* were improved sensitivity and calibration accuracy, and the ability to produce fully sampled diffraction-limited images (nominal image pixel scales of 2.45 and 4.0′′/pix at 24 and 70 µm respectively). For the shortest typical observations (integration time ~100 s), 1σ point-source noise levels of approximately 0.1 and 5 mJy (12.1″ and 5.5″), and 1σ random flux calibration uncertainty of approximately 1% and 4%, were achieved at 24 and 70 µm respectively. The sensitivity and calibration stability yielded unprecedented dust mass detection limits. Fig. 2.4 (top) shows the distance detection limits for the same models as Figs. 2.2 and 2.3, and for dust masses a factor of 50 lower (bottom). *Spitzer* provided almost two orders of magnitude greater dust mass sensitivity than *IRAS*, and the ability to detect significant numbers of debris discs around early type stars at distances of over 100 pc.

Many debris disc surveys have been performed with *Spitzer*, observing over 1000 stars in total. Papers to date have focused on particular groups of stars: Sun-like (F–K type) stars
2.2. PREVIOUS SURVEYS

Figure 2.4: Spitzer/MIPS distance limits for Kuiper and Asteroid belt analogues as a function of stellar spectral type (effective temperature). Top: dust masses of $10^{-2}$ and $10^{-4}M_\odot$ as in Figs. 2.2 and 2.3, showing that Spitzer offers greatly improved sensitivity (larger distance limits) to discs around mid-K and earlier type stars, although sensitivity for M type stars is poor relative to (sub-)mm observations. Bottom: dust masses of $2 \times 10^{-4}$ and $2 \times 10^{-6}M_\odot$, a factor of 50 lower than in Figs. 2.2 and 2.3, showing that such low masses can be detected around significant numbers of A–G type stars. As for IRAS, the spectral type sensitivity bias shown here appears to be borne out in the Spitzer detection statistics for A, F–K and M type stars.

[Beichman et al., 2005, 2006a,b, Bryden et al., 2006, Hillenbrand et al., 2008, Trilling et al., 2008, Lawler et al., 2009, including the Formation and Evolution of Planetary Systems (FEPs) survey], B/A type stars [Rieke et al., 2005, Su et al., 2006, Morales et al., 2009], M type stars [Gautier et al., 2007, Plavchan et al., 2005, 2009], multiple systems [Trilling et al., 2007], planet hosts [Bryden et al., 2009], and clusters [Stauffer et al., 2005, Gorlova et al., 2006, Siegler et al., 2007, Cieza et al., 2008a, Currie et al., 2008a,b, Gáspár et al., 2009, Sierchio et al., 2010].
There have also been many studies of individual systems (for instance, using special observing modes to give increased spatial sampling), and surveys filling in volume limited samples of stars [Koerner et al., 2010, and chapter 4 of this thesis].

Key results to date include: determining approximate debris incidence rates of 33%, 16% and ∼4% for A, F to early K, and late K and M-type stars respectively (although these include significant biases, both in sample selection and intrinsic sensitivity as shown in Fig. 2.4); showing that debris incidence rates for multiple systems are the same or even higher than for single star systems; and revealing numerous warm discs due to the excellent quality of MIPS 24 μm photometry.

Over the next few years there will be a period of consolidation, which will likely result in numerous statistical studies of debris discs. The vast number of stars observed by Spitzer, combined with the impressive sensitivity and photometric accuracy, will provide a significant legacy, including a many-fold increase in the number of known debris discs. The variety of observing modes, wavelength coverage and integration times used for different sources, however, will complicate the analysis and ultimately limit the conclusions which can be drawn. As Fig. 2.4 shows, the lack of sensitivity at wavelengths longer than 70 μm means that potential discs around late type stars still remain largely unexplored.

### 2.2.5 Science Results and the State of the Field

#### Spectral Type and Related Stellar Parameters

All of the surveys of statistically significant samples of stars performed to date have observed at wavelengths of ≤70 μm. These have consistently found that debris discs are most commonly detected around B/A type stars, with detection rates of 30% or higher with Spitzer [Su et al., 2006]. For ‘Sun like’ F–K type stars the detection rates have typically been a factor of two lower than for B/A type stars [Hillenbrand et al., 2008, Trilling et al., 2008]. For M type stars very few debris discs have been detected at all [e.g. Gautier et al., 2007], and those which have been detected are typically around very young stars in moving groups [Plavchan et al., 2009].

As has been shown above in Figs. 2.2 and 2.4, this result can be explained as an observational bias due to the use of only wavelengths of 70 μm and less. For instance, discs which would be easily detectable at 60/70 μm around an A type star at a distance of 40 pc from the Sun could be undetectable around any M type stars. This bias is primarily due to the dependence of the dust temperature, and hence the peak wavelength of dust thermal emission, on stellar luminosity. Removing, or at least lessening, this bias is one of the primary motivations for longer wavelength surveys as described in the following sections.
2.2. PREVIOUS SURVEYS

Debris Disc Evolution

One of the primary goals of many of the surveys performed with Spitzer has been to investigate the evolution of debris disc properties. One such survey was the Formation and Evolution of Planetary Systems [FEPS, Meyer et al., 2006] survey, which observed ~328 Sun like stars (defined in this case as approximately F8–K0 spectral types) spread over the age range 3 Myr to 3 Gyr. The evolution of debris discs around A type stars has also been investigated [Su et al., 2006]. Evolutionary studies are generally complicated by the difficulty of determining the ages of stars, so there are often significant selection effects in order to observe stars with known ages, and simply due to the lack of stars younger than ~100 Myr within ~100 pc of the Sun.

The overall result from these studies has been that the maximum observed 24 and 70 μm flux excess, and maximum fractional luminosity, decrease with age as \( \sim t^{-1} \) [Su et al., 2006, Hillenbrand et al., 2008, Wyatt et al., 2007b, Wyatt, 2008]. The detection rate of debris discs, at least at 70 μm, however, appears to be largely independent of age.

Host Star Metallicity

Studies have been performed to look at how the host star metallicity distributions differ between debris disc host stars and stars lacking detected discs. Greaves et al. [2006] performed this analysis for F–K type stars which are in radial velocity planet detection surveys, and have accurate metallicities determined by Fischer and Valenti [2005], as used in establishing the giant planet – metallicity correlation (§1.2.2). Saflle et al. [2008] measured metallicities for 113 B–K type debris disc host stars found from the literature. Both of these studies showed no significant difference between the metallicity distributions of debris hosting and non-debris hosting stars, in contrast to a clear difference between giant planet hosting stars and other stars. The effect of metallicity on debris detection rates is investigated within a volume limited sample of 130 A type stars in chapter 4 of this thesis.

Binary Stars

It is clear that the presence of multiple stars in a system could affect the presence and properties of debris discs due to gravitational interactions. Trilling et al. [2007] performed a survey of binary systems with F and late-A spectral type primaries. They concluded that debris incidence rates were overall higher than for single stars, and that there was a deficit of debris discs in systems with separation of order 10 AU. The comparison between debris detection rates for binary and single stars in that work, however, is not robust, as the detection criteria used for the two samples are quite different. This is due to using detections for single stars from a different work [Su et al., 2006], which applied much more stringent detection criteria. To test
the true effect of binarity and binary separation on debris disc incidence (for A type stars at least) is one of the goals of chapter 4 of this thesis.

2.3 The SCUBA-2 Unbiased Nearby Stars (SUNS) survey

The Sub-millimetre Common User Bolometer Array 2 instrument [SCUBA-2; Holland et al., 2003, 2006], operating simultaneously at 450 and 850 μm, is currently undergoing commissioning on the JCMT. SCUBA-2 differs in two primary ways from previous and current astronomical bolometer arrays. Firstly, the detectors employ superconducting transition edge sensors (TES) instead of traditional semiconducting thermistors. This provides improved sensitivity – the per bolometer sensitivity of SCUBA-2 is expected to be several times greater than its predecessor, SCUBA. Secondly, the number of bolometers in each focal plane (∼5000) is over an order of magnitude greater than any current or previous instrument, which increases mapping speed by the same factor. The result is that SCUBA-2 will be able to produce confusion limited 850 μm maps (1σ noise < 1 mJy/beam) of the 7′ × 7′ field of view in as little as one hour. The combination of the short integration times for compact fields, and wide area mapping speed, make SCUBA-2 ideally suited to large surveys. Seven such programmes, each allocated between 300 and 1100 hours, have been accepted and together make up the JCMT Legacy Survey\(^5\).

![Figure 2.5: Left: Photograph of SCUBA-2 (left) and its predecessors, SCUBA (right) and UKT14 (center). The large size of the SCUBA-2 cryostat is due to the size of re-imaging lenses which are necessary to match the compact detector arrays to the full JCMT field of view. Right: the James Clerk Maxwell Telescope (JCMT).](image)

The SCUBA-2 Unbiased Nearby Stars survey [SUNS; Matthews et al., 2007] is allocated a total of 300 hours to observe 500 nearby A–M type main-sequence star systems in search of debris discs. The observations will reach a 1σ RMS noise of approximately 1 mJy/beam at 850 μm. This will be an order of magnitude more systems than have previously been observed.

\(^5\)http://www.jach.hawaii.edu/JCMT/surveys/JLSDescriptors.html
to a comparable depth at sub-millimetre wavelengths. The primary goals of this survey are,

- To determine unbiased statistics on the incidence of discs around nearby stars
- To constrain disc masses, temperatures, dust grain properties and spatial distribution for discs previously detected in the mid- and far-IR
- To discover discs too cold to be detected in the mid/far-IR
- To be the basis of source lists for future observing campaigns using instruments such as ALMA
- To provide limits on the presence of dust around nearby stars that are vital to future missions such as Darwin and TPF.

The statistical nature of the survey is of particular importance. Dust mass detection limits for observations at $850 \, \mu m$ are relatively unbiased by dust temperature. This differs from shorter wavelengths where dust mass sensitivity decreases rapidly when shortward of the thermal emission peak ($T \lesssim 5100/\lambda [\mu m]$). The adoption of a homogeneous observation depth, in contrast to the majority of other surveys since IRAS, further simplifies statistical analysis.

The targets comprise five spectral type sub-samples, each containing the 100 closest stellar systems to the Sun in the declination range $-40^\circ < \delta < +80^\circ$ with primaries of spectral types A, F, G, K and M. The choice of overall sample size was driven by the need to compare detection rates between various sub-samples, such as systems with and without planets or the spectral type classes, with sizes of $\sim 10$–$20\%$ of the overall sample and detection rates of order $10\%$. The selection and general properties of these targets are the subject of the following chapter. The declination limits are chosen such that all targets can be observed at an elevation of at least $30^\circ$ (airmass, $\sec z = 2.0$). There is no bias based on previous detections of dust or planets, or any system parameters other than distance and spectral type. The statistical results of the survey will apply in general to field stars.

The key advantage of the JCMT over other ground based telescopes such as APEX, and space based far-IR/sub-millimetre observatories such as Spitzer and Herschel, is its large size and correspondingly small beam sizes and low confusion limits. The $850 \, \mu m$ FWHM beam size of $15''$ will allow larger discs to be spatially resolved. Discs detected at $850 \, \mu m$ which show signs of significant structure will be followed up with deep $450 \, \mu m$ observations (FWHM beam size of $8''$). Such structure in sub-millimetre images of discs is important as the spatial distribution of large grains can be used to infer the presence and locations of planets. Although both bands are observed simultaneously by SCUBA-2, the integration time required to reach the same dust mass surface brightness sensitivity at $450 \, \mu m$ is much longer due to the higher angular resolution.
and poorer atmospheric transmission. The time for these follow-up observations (several hours each) will have to be applied for.

There are ancillary uses of the SUNS data. The observations are randomly distributed over the sky and will cover at least the SCUBA-2 instantaneous field of view of $7' \times 7'$ (i.e., a total area of at least $6.8$ square degrees for the 500 maps) to an almost confusion-limited depth. This will yield statistics on sub-mm galaxies [e., Smail et al., 1997], and any objects detected in the fields will lend themselves to follow-up with Adaptive Optics (AO) instruments, as the SUNS target stars will make excellent nearby AO guide stars.

### 2.4 Surveys with Herschel

The European Space Agency’s *Herschel* Space Observatory [Pilbratt, 2008] has a 3.5 m diameter passively cooled primary mirror (four times the diameter of *Spitzer’s*), and three instruments operating between 55 and 672 μm. *Herschel* bridges the gap between *Spitzer* and ground-based sub-millimetre facilities, and spans the peak of thermal emission from cold dust ($\sim 10–90$ K). Two instruments are particularly suited to debris disc surveys. The Photodetector Array Camera and Spectrometer [PACS; Poglitsch et al., 2010] provides imaging at 70, 100 and 160 μm, and the Spectral and Photometric Imaging Receiver [SPIRE; Griffin et al., 2010] provides imaging and point-source photometry at 250, 350 and 500 μm. The third instrument, the Heterodyne Instrument for the Far-Infrared [HIFI; de Graauw et al., 2010] is a heterodyne spectrometer. PACS has two detector arrays, one for 70/100 μm and one for 160 μm. These share the same field of view, allowing simultaneous observation at 70 or 100 μm and 160 μm. Each of the three bands of SPIRE have a separate detector array. These observe the same field of view, allowing simultaneous observation in all three bands.

![Herschel spacecraft and instruments](image)

**Figure 2.6:** *Herschel*. Left: spacecraft. Centre: instruments (HIFI at the front, PACS to the left, and SPIRE to the rear right). Right: orbit.

The size of *Herschel’s* mirror gives greatly improved spatial resolution compared to previous
observatories at the PACS wavelengths (factors of four and twelve over Spitzer and IRAS respectively). The PACS FWHM beam sizes in the relevant observing mode are approximately 5.6", 6.7" and 11" at 70, 100 and 160 μm respectively [Poglitsch et al., 2010]. This results in vastly improved mitigation of confusion, both extragalactic and due to galactic cirrus. The predicted extragalactic confusion limits (1σ) at 100 and 160 μm are 0.52 and 1.7 mJy/beam respectively. At the SPIRE wavelengths the beam sizes are larger [18", 25" and 37"; Griffin et al., 2010] which, combined with the typical SED of background galaxies, causes extragalactic confusion to be a significant limitation. In all three SPIRE bands the extragalactic confusion limits (1σ) are 6–7 mJy/beam [Nguyen et al., 2010].

It is anticipated that Herschel will have a routine science operations period of of at least three years, ultimately limited by the supply of onboard cryogen [Pilbratt, 2008]. The majority (57%) of routine observing time is allocated to 42 Key Programmes, half of which are exclusive to Herschel instrument teams (Guaranteed Time Key Programmes, GTPs), and half which were proposed by the general astronomical community (Open Time Key Programmes, OTKPs). Among these are three programmes (two OTKPs of 140 hours, and a GTPK of 61 hours) dedicated to debris discs around nearby stars. The Disc Emission via a Bias-free Reconnaissance in the Infrared/Sub-millimetre OTKP [DEBRIS; Matthews et al., 2010] is statistical survey, with a large minimally biased sample of 446 targets. The DEBRIS survey is discussed in more detail below. The DUST disks around NEarby Stars OTKP [DUNES; Eiroa et al., 2010] is a more targeted study of primarily Sun-like (F–K spectral type) nearby stars which are already known to harbour debris discs or exoplanets. The Stellar Disc Evolution GTPK is a detailed study of six of the most famous and massive nearby debris disc systems (Vega, Fomalhaut, β Pic, ε Eri, τ Cet, AU Mic), using imaging in all PACS and SPIRE bands as well as spectroscopy of gas emission lines.

Herschel was launched on 14 May 2009, entering orbit around the Sun-Earth L2 point (Fig. 2.6, right) on 13 July 2009. A period of commissioning and Performance Verification (PV) was completed by mid October 2009. This was followed by a Science Demonstration Phase (SDP), in which test observations from the Key Programmes were performed. These SDP observations have been crucial in optimising the observing modes used by the Key Programmes, and also produced early science results. Herschel is now performing routine observations.

2.4.1 The DEBRIS survey

The DEBRIS survey is designed in a similar manner to the SUNS survey with SCUBA-2, and the two surveys are highly complimentary and share the same goals. DEBRIS will observe a total of 446 nearby A–M spectral type stellar systems at 100 and 160 μm using PACS, with
a target 100 $\mu$m 1$\sigma$ noise of 1.2 mJy/beam. Approximately 100 systems will be followed up with SPIRE. The target selection is covered in the following chapter. The selection criteria are essentially as for the SUNS survey, with the declination cut replaced by a cut in predicted cirrus confusion noise$^6$. At the time of writing, approximately 20% of DEBRIS observations have been performed.

The PACS wavelengths are ideally suited to such a statistical survey, as they offer the greatest ever dust mass sensitivity for Kuiper Belt analogues. Fig. 2.7 compares the detectable dust mass as a function of dust temperature for 3$\sigma$ excess detections with various instruments for a solar analogue at 17 pc (the median distance for G-type stars in the DEBRIS and SUNS surveys). The figure includes the effect of calibration uncertainty, which sets a lower limit on the detectable excess ratio when the stellar photosphere is significantly detected. PACS offers the greatest ever sensitivity to dust in the 20–100 K temperature range, and thus DEBRIS will be the deepest ever large survey for cold, Kuiper Belt like, debris discs in terms of dust mass.

The angular resolution at the PACS wavelengths also offers the opportunity to spatially resolve many discs for the first time.

For detected discs, photometry at the SPIRE wavelengths is important for modelling the dust properties as, when combined with the PACS photometry, the wavelength region where the dust opacity transitions from constant to a power law is well covered.

**SDP Observations**

SDP observations with PACS were performed using two candidate observing modes – a chopped mode and a scan map mode. Eight targets were observed with PACS during the SDP, which were chosen to have known (at least tentatively) excess and a wide range of predicted flux densities and angular extents. Preliminary results for three of the best resolved discs in the SDP data ($\beta$ Leo, $\beta$ UMa, and $\eta$ Crv) have been published in Matthews et al. [2010] (Fig. 2.9).

A further result from SDP was the finding of an especially cold (∼30 K) disc around $\zeta$ Tuc, the 5th closest F-type star to the Sun at a distance of 8.6 pc. This disc had previously only been detected at 70 $\mu$m [Trilling et al., 2008], and so the temperature was unknown. The author is obtaining ground-based 870 $\mu$m imaging with the LABOCA bolometer array (see chapter 5) for this object (ESO proposal 386.C-0707, PI: Phillips). The flux density distribution of $\zeta$ Tuc, including the preliminary PACS 100 and 160 $\mu$m photometry, is shown in Fig. 2.10.

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$^6$Emission from galactic cirrus is typically characterized by a modified black body with $T \sim 20–30$ K and $\beta \approx 2$. This results in the lowest contrast for debris discs, which have $\beta \approx 1$, around the peak of the cirrus emission at $\sim 100–200$ $\mu$m
Figure 2.7: Dust mass sensitivity (3σ detection limit, including calibration uncertainty) as a function of dust temperature, and corresponding orbit radius for black body grains, for a typical solar-type star in the DEBRIS and SUNS surveys (Teff = 5800 K, L = L⊙, d = 17 pc). The pale colored regions highlight the temperature / radius ranges in which Herschel (green), SCUBA-2 (red) and Spitzer (blue) are the most sensitive facility. Assumed 3σ flux detection limits and 3σ calibration uncertainties: SCUBA-2: 3.0 mJy, 30% (predicted for SUNS observations); PACS 160 μm: 12.8 mJy, 15% (typical measured DEBRIS map RMS, target calibration uncertainty); PACS 100 μm: 5.1 mJy, 15% (as 160 μm); MIPS 70 μm: 20 mJy, 10%; IRAS 60 μm: 600 mJy, 60%; MIPS 24 μm: 0.1 mJy, 6% (entirely calibration dominated).

Routine Observations

Based on analysis of the SDP observations, all routine observations are being performed using a compact scan map mode as shown in Fig. 2.11. Cross scanning is used to minimise the effects of 1/f noise. Such noise is evident as linear features in the left and right maps in Fig. 2.9, as these observations did not use cross scanning (the centre images were performed in the chopped mode). With Herschel, cross scanning requires a separate observation for each scan orientation, although these are linked (performed consecutively) to reduce overheads. As of 9th September 2010, 107 targets have been observed. The mean 1σ noise for these observations, using the current reduction pipelines, are 1.73 and 4.25 mJy/beam at 100 and 160 μm respectively. Future improvements to the data reduction, including optimisation of filtering, will likely reduce these values (cf. the target of 1.2 mJy/beam at 100 μm).
Figure 2.8: Herschel distance limits for Kuiper and Asteroid belt analogues with dust masses of $2 \times 10^{-4}$ and $2 \times 10^{-6} M_\oplus$ respectively (as in the bottom row of Fig. 2.4 for Spitzer/MIPS), as a function of stellar spectral type (effective temperature). These show that Herschel’s 100 and 160 μm bands offer the possibility to detect such low mass discs around late-F to M type stars, which would not have been detectable by Spitzer.

Pointings and Multiple Systems

The author has been responsible for generating the pointings for the DEBRIS observations using the database developed in the following chapter. For multiple star systems, the pointing is centred between components within 20″ of the primary star. This applies to 50 of the 446 DEBRIS systems. The limit of 20″ is a compromise between including secondary stars and maintaining even coverage of potentially resolved discs around the primary stars (refer to the region of maximal coverage in Fig. 2.11). Components separated further than 20″ from the primary will receive less coverage or not be observed at all. This applies to 49 systems. There are a small number of exceptions due to overlap between targets of the DUNES survey (and potentially other surveys) which directly target secondary stars in some DEBRIS systems.
24. SURVEYS WITH HERSCHEL

Figure 2.9: First results from the DEBRIS survey: three resolved discs observed during the Science Demonstration Phase. Taken from Matthews et al. [2010].

Figure 2.10: Preliminary flux density distribution for ζ Tuc using PACS photometry. The region to be probed by LABOCA is highlighted.
Figure 2.11: Coverage of PACS and SPIRE scan maps used by DEBRIS. The region of maximal coverage in the PACS maps is approximately a circle with a diameter of 50°.
2.5 Summary

There have been many surveys for debris discs in the 26 years since the detection by IRAS of Vega’s disc. The catalogues and archived data from IRAS have been searched by many authors and detections have been followed up by many other facilities. The detection of $\sim$200 debris discs by IRAS compares very favourably to all instruments and surveys since, highlighting the benefit of large unbiased surveys. The IRAS data products are still regularly used today (e.g., chapter 4 of this thesis). The minimum dust mass detectable by IRAS for Kuiper Belt analogues around the nearest A type stars is approximately $10^{-3} M_{\oplus}$, however, the minimum mass increases rapidly for later spectral types (Fig. 2.2). It is thus unsurprising that the majority of IRAS detected discs are around A and F type stars.

ISO provided improved photometry for IRAS detected discs, detected a small number of discs which were below the IRAS detection limits, and allowed IRAS detections to be checked for source confusion.

Spitzer made a great leap forward in sensitivity, imaging resolution and field of view. Many surveys have targeted well over 1000 stars in total, searching for debris. Dust mass detection limits are typically at least an order of magnitude lower than IRAS or ISO. Despite the impressive sensitivity of Spitzer at wavelengths of 70 $\mu$m and less, the lack of longer wavelength capability means that as for IRAS, the Spitzer detection statistics are inherently biased toward dust around early type stars (Fig. 2.4).

(Sub-)millimetre surveys have so far had relatively little impact on detection statistics. However, (sub-)mm observations have made a huge contribution to the interpretation of bright discs by spatially resolving emission from large dust grains. It has been shown (Fig. 2.3) that the number of stars within the distance limit for detecting a given mass of dust is approximately independent of spectral type, offering the possibility for (sub-)mm surveys to discover discs around late type stars which were not detectable at shorter wavelengths.

New generations of (sub-)mm instruments will for the first time make it feasible to perform deep surveys of hundreds of stars. The SUNS survey will use the new SCUBA-2 bolometer array instrument on the 15 m JCMT to survey 500 stellar systems, comprising the nearest 100 systems accessible from JCMT with A, F, G, K and M spectral type primaries.

Herschel fills the gap between the previous infrared space missions and ground-based sub-mm facilities. PACS performs sensitive high resolution (FWHM $\sim$ 6–12$''$) imaging at 70–160 $\mu$m, reaching the lowest ever dust mass limits and spatially resolving many discs. The DEBRIS Key Programme is observing 446 stellar systems selected from the same sample as the SUNS survey. The following chapter describes the selection and properties of the targets of these surveys.
Chapter 3

Target selection for the SUNS and DEBRIS surveys

3.1 Introduction

The solar neighbourhood is an ideal testing ground for the study of debris discs and planetary systems. Close proximity maximises dust mass sensitivity and can allow systems to be spatially resolved. Systems near the Sun span a wide range of stellar parameters, e.g., mass, age, metallicity and multiplicity. Whilst determining these parameters may not be easy, the diversity included in volume limited samples makes them ideal for legacy surveys where one may wish to investigate trends as a function of many system parameters.

This chapter presents a sample of 629 systems of nearby stars, composed of five all-sky volume limited subsamples of systems with main-sequence primaries of spectral type A, F, G, K and M. The systems have each been assigned an Unbiased Nearby Stars (UNS) identifier, composed of the spectral type letter and a number increasing with distance within each subsample. These form the basis of the target lists of the SUNS and DEBRIS surveys which were described in the previous chapter. The volume limits were set indirectly by the SUNS criterion that each subsample should contain 100 systems in the declination range $-40^\circ < \delta < +80^\circ$. It is hoped that this sample will be of use in other analyses of nearby stars, in a similar way to the Catalogue of Nearby Stars [CNS3; Gliese and Jahreiß, 1991], which contains stars known to be within 25 pc of the Sun in 1991.

To facilitate the target selection, a database was created from several catalogues relevant to nearby stars. This was necessary as no single catalogue covered all the required distance and spectral type ranges with sufficient completeness, or provided up to date measurements.
of all necessary parameters. Entries from catalogues were matched by a combination of cross
identifications, astrometry (positions and proper motions), and heuristics. The result was a ta-
ble containing cross-identifications, astrometry, distances and spectral types for approximately
$6 \times 10^5$ stars. This table was used to select the targets and determine memberships of multiple
systems. Further checks and searches for components of multiple systems have been performed.

This work was primarily performed during 2006 and 2007, and the target lists were essen-
tially frozen at the time of the DEBRIS survey proposal submission in October 2007. The
author has been responsible for all the work presented here, with the exception of the choice of
selection criteria for the SUNS survey. The author was heavily involved in the development of
the DEBRIS survey, including the selection criteria. This work has previously been published
in Phillips et al. [2010].

64
3.2 Selection Criteria

The systems all have primaries (defined here as the component with the brightest visible magnitude) which are believed to be main-sequence (i.e., hydrogen burning) stars\textsuperscript{1}. The sample is split into 5 volume limited subsamples based on spectral type: A, F, G, K and M. In the rest of this work the term “X type system” is used to mean “system with X type primary”. Using separate subsamples is necessary due to the steep nature of the stellar mass function [e.g., Kroupa et al., 1993] which, for example, means that a single volume limited sample would contain over 100 times as many M type systems as A type systems. The choice to use spectral types to split the sample, rather than stellar mass, was purely practical as, with the exception of certain binary systems, stellar masses cannot be directly determined observationally. Using spectral types does, however, have the effect that the subsamples cover quite different ranges in logarithmic mass space (approximate $\Delta \log M$ of 0.70, 0.20, 0.14, 0.21, 0.22 for M-A respectively).

3.2.1 Early Type Limit

The early type, upper mass, spectral type limit of A0 ($\sim 2.5M_\odot$) was chosen as stars of earlier type are too rare in the solar neighbourhood to build a suitably large sample. The short lifetime of O and B type stars (\(\lesssim 300\) Myr) means that they are generally located in regions of recent star formation, open clusters or moving groups. The distance required to obtain a sufficient number of B type systems would be approximately 97 pc, which is significantly greater than the galactic disc scale height for such stars [\(\sim 45\) pc; Reed, 2000], creating significant bias of the positions of systems toward the galactic plane. The combination of potentially significant interstellar extinction at optical wavelengths, confusion from galactic cirrus at far-IR wavelengths, and the inherent biases of their youth and clustering, mean that O and B type systems are not included in what is generally a sample of field stars distributed uniformly on the sky with negligible interstellar extinction and ages of several hundred Myr or greater.

3.2.2 Late Type Limit

A conservative late type limit of M7.0 was chosen to avoid the inclusion of any brown dwarfs, and also to improve the completeness of the M type subsample. M type stars span the largest log $M$ range of any of the spectral classes, so making a cut at M7.0 does not restrict the statistical usefulness of the sample. The cut at M7.0 effectively excludes brown dwarfs older

\textsuperscript{1}Pre-main sequence (PMS) stars have not been specifically excluded, although very few PMS primaries are expected within the subsample distance limits.
than a few hundred Myr, and places a lower mass limit of $\sim 0.1M_\odot$ for stars older than $\sim 1$ Gyr (Fig.3.1).

\textbf{Figure 3.1:} Left: Spectral type as a function of mass for 3 different ages; 10Gyr (solid line), 1Gyr (dashed line) and 100Myr (dash-dotted line). The dotted vertical line is the hydrogen burning minimum mass ($M_{\text{HBBM}}$). Taken from Baraffe and Chabrier [1996]. Right: Evolutionary tracks for objects between 0.001$M_\odot$ and 1$M_\odot$ on a HR diagram. The dotted lines are isochrones of 1Myr, 10Myr, 100Myr and 5Gyr from right to left. Note that for objects below about 0.1$M_\odot$, $T_{\text{eff}}$ alone cannot distinguish masses. Taken from Chabrier and Baraffe [2000].

\section*{3.2.3 Luminosity Class Limits}

Initially it had been proposed to only include systems with primaries of spectroscopic luminosity classes V and IV-V. This criterion was retained for G, K and M classes, but was relaxed for A and F type stars, where there is little distinction between luminosity classes V–III (Fig.3.2), and there is not a simple relationship between luminosity class and evolutionary stage [e.g., Gray et al., 2001b,a]. Candidates for the A and F samples (and other candidates without accurately known luminosity classes) were evaluated using their position on a Johnson B.V absolute colour-magnitude diagram. Fig. 3.3 shows such diagrams for the final sample overlaid with solar composition isochrones and zero age main sequences (ZAMS) for metallicities from +0.5 to −2.0. A certain amount of leeway had to be allowed for unknown metallicity, and uncertainties in photometry (e.g. unresolved secondaries in close binaries) and parallax.
3.2.4 Multiple Star Systems

There is no discrimination of multiple star systems, and they are included naturally within the volume limits. Common proper motion stars (with compatible parallax where available) are considered as members of the same system, with no specific limit on the binary separation. Stars with common space motion but large ($\gg 1^\circ$) angular separation are not, however, considered as systems. This definition of system membership was primarily chosen for convenience of target selection, but fits well with the statistical goals of the surveys. With the exception of stars in moving groups, each system can be considered to represent a different point in age and composition. The fact that several interesting objects (e.g., with known IR excess or planets) are considered here as secondaries does not affect the statistical usefulness of the sample, although it has the disadvantage that such objects may not be observed by the surveys (see below).

3.2.5 Subsample Sizes

The number of systems in each subsample was determined by the selection criteria for SUNS, which required 100 systems in each subsample in the declination range $-40^\circ < \delta < +80^\circ$. The SUNS sample sizes were chosen to allow detection rates for various subsets, e.g. planet hosts, to be distinguished [Matthews et al., 2007, and chapter 2 of this thesis]. The size of the all-sky
Figure 3.3: Johnson B,V absolute colour-magnitude diagrams for system primaries. Left: Overlaid with zero age main-sequences (ZAMS) for stars from $0.2 M_\odot$ upwards with [Fe/H] = +0.5, 0.0, −0.5, −1.0, −2.0 (from top to bottom). The ZAMS curves are produced from [$\alpha$/Fe] = 0.0, $Y = 0.245 + 1.6Z$ evolutionary tracks from the Dartmouth Stellar Evolution Database [Dotter et al., 2008], with values taken at 2% of the total lifetime of the stars. Right: Overlaid with [Fe/H] = 0.0, [$\alpha$/Fe] = 0.0 isochrones from the Dartmouth Stellar Evolution Database [Dotter et al., 2008] with ages of 0.25, 0.5, 1, 2, 4, 8 Gyr (with turn-offs going from left to right). The photometry is mostly converted from Tycho photometry (Tycho-2 or TDSC) using transformations for unreddened main-sequence stars. For most M type targets Johnson B,V photometry from various sources was used (see text). Note that primaries in some close binaries are not individually resolved in this photometry.

For most M type targets Johnson B,V photometry from various sources was used (see text). Note that primaries in some close binaries are not individually resolved in this photometry.

subsamples, assuming the systems are isotropically distributed, is expected to be,

$$N_{\text{all-sky}} = 100 \times \frac{\int_{-90}^{90} \int_{-40}^{40} \cos \delta d\delta}{\sin 80^\circ} = 122.9.$$  

(3.1)

### 3.2.6 DEBRIS Targets

The DEBRIS target list comprises the nearest systems presented here (all-sky), subject to a cut in the predicted 100 $\mu$m cirrus confusion level towards each system. The confusion prediction was taken from the Herschel Confusion Noise Estimator (HCNE), which is part of the Herschel Observation Planning Tool (HSPOT). Systems with total predicted confusion for point-source detections greater than 1.2 mJy/beam, corresponding to twice the predicted extragalactic confusion limit, were rejected. To even out the number of systems of each spectral type after the confusion cut, the three furthest F type systems were removed from the DEBRIS target list.

To maximise the number of systems observed, DEBRIS will not image secondary components in systems where they will not fit in the field of view ($\sim 50''$ diameter of maximal coverage, see chapter 2) with the primary. This will affect approximately 50 systems.
3.2.7 SUNS Targets

The SUNS target list is simply the nearest 100 systems in each subsample here which have $-40^\circ < \delta < +80^\circ$ (with this sample it does not make any difference whether the cut is made in B1950 or J2000/ICRS equinox declination, but J2000/ICRS should be assumed). The large ($\sim 10' \times 10'$) field of view of SCUBA-2 means that a maximum of 13 systems will have components not observed with the primary star.
3.3 Sources of Data

3.3.1 Parallaxes

Hipparcos-based parallaxes were taken from Hipparcos, the New Reduction of the Raw Data [HIPnr; F. van Leeuwen, 2007] and several papers which applied special analysis to multiple systems: the General Notes issued with the original Hipparcos catalog [HIPgn; Perryman et al., 1997], Falin and Mignard [1999], Söderhjelm [1999], and Fabricius and Makarov [2000]. Parallaxes from HIPnr were used unless one of the other resources had a lower uncertainty. In cases where more than one of the other resources provided a parallax for the same Hipparcos system, the parallax has been taken from the first resource in the order: Fabricius and Makarov [2000], Söderhjelm [1999], Falin and Mignard [1999]; HIPgn. Hipparcos parallaxes from multiple resources for the same Hipparcos system were not averaged in any way to avoid underestimating the uncertainty in the averaged values, as they have all been reduced from the same data.

The other large parallax resource used was the 4th edition of the Yale General Catalog of Trigonometric Parallaxes [GCTP or YPC; van Altena et al., 1995], which contains approximately 2300 systems not measured by Hipparcos due to the magnitude limit of $V \sim 12$ and the targeted nature of the Hipparcos astrometry mission.

In addition, for many M dwarfs, parallaxes from several smaller papers were used [e.g., Henry et al., 2006, Jao et al., 2005, Costa et al., 2005, Hershey and Taff, 1998, Benedict et al., 1999, Weis et al., 1999, Ducourant et al., 1998], as well as some unpublished values from the RECONS2 consortium (Henry, private communication).

Where reliable parallaxes from multiple independent sources, or separate parallaxes for individual components in a system, were available, an uncertainty weighted average was taken:

$$\pi_{\text{adopted}} = \frac{\sum_i \pi_i / \sigma_i^2}{\sum_i 1 / \sigma_i^2} \quad \text{and} \quad \sigma_{\text{adopted}} = \sqrt{\left(\frac{1}{\sum_i 1 / \sigma_i^2}\right)}$$ (3.2)

Two or more parallaxes were used for 81% of systems, and three or more were used for 7% of systems. These cases are mostly due to overlap with Hipparcos- and ground-based (e.g. YPC) parallaxes.

3.3.2 Spectral Types

For A-K type stars, spectral types from Gray et al. [2003, 2006] have been used where they were available. Gray et al. have been obtaining spectra and determining spectral types and stellar parameters ($T_{\text{eff}}$, $[M/H]$, log $g$) for stars considered to be within 25 pc and of spectral

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type earlier than M0 or with no spectral type in the *Hipparcos* catalogue [HIP; Perryman et al., 1997]. For stars without published Gray et al. types, types from the Michigan Catalogue of HD stars [Houk and Cowley, 1975, Houk, 1978, 1982, Houk and Smith-Moore, 1988, Houk and Swift, 1999], which includes all HD stars south of $\delta_{1900} = +05^\circ$, have been used. If types from neither Gray et al. or Houk et al. were available, types in compilations such as the 5th revised edition of the Bright Star Catalogue [BSC5; Hoffleit and Warren, 1995], HIP, or the 2nd edition of the Catalog of Components of Double & Multiple stars [CCDM; Dommanget and Nys, 2002], have been used. These fall-back types are not considered to be accurate, and were largely ignored in the selection process in favour of photometry.

For ~K5 and later stars, spectral types from the Palomar/MSU Nearby-Star Spectroscopic Survey [PMSU; Reid et al., 1995, Hawley et al., 1996] have generally been used. This provides spectral types for almost all late type stars in the 3rd Catalogue of Nearby stars [CNS3; Gliese and Jahreiß, 1991]. A large number of nearby M dwarfs also have measured spectral types in the system of Kirkpatrick et al. [1991], however, it was chosen to use PMSU types wherever possible for homogeneity. The difference between PMSU and Kirkpatrick et al. types is rarely more than one subtype, even though the spectral classification method differs significantly (Fig. 3.4).

For newly discovered nearby M dwarfs not included in the PMSU, types in the Kirkpatrick et al. system [e.g. from Henry et al., 2006] have been adopted.

Table 3.1 shows a comparison between types from the PMSU and Gray et al. [2003, 2006], with stars ordered by increasing distance. Due to the selection criteria of these two surveys, the overlapping stars are mostly near the K/M type divide. There are no cases within the distance limit used here for M type systems where the PMSU and Gray et al. types conflict as to whether a star is K or M type. There are, however, six stars within the larger K type distance limit where these sources conflict (highlighted in bold in Table 3.1). With the exception of GJ 103, the PMSU classifies these stars as late K type whilst Gray et al. classify them as early M type. For consistency across the K/M type divide, the PMSU types are used to determine the K type membership, so the five stars in Table 3.1 which are classified in the PMSU as K type are included in the K type subsample, however GJ 103, which the PMSU classified as M0 type, is not.

### 3.3.3 Photometry

Whilst distance and spectral type were the primary selection parameters, it was necessary to use photometry both for determining luminosity and when determining spectral class where only low accuracy spectral types were available. As distinguishing between dwarfs and giants for K/M type stars of known distance is very simple, and because accurate spectral types were
**Figure 3.4:** M type star classification method used by the PMSU survey. Taken from Reid et al. [1995]. Left: The full depth of the TiO absorption band, defined as $T(7126 : 7135)/T(7042 : 7046)$, was used as the primary measure of spectral type (marked TiO5 here, on the spectrum of an M4 type star). Right: TiO5 index for stars from Henry et al. [1994] (squares) and very low mass stars from Kirkpatrick et al. [1995] (circles), which were used to calibrate the TiO5 to spectral type relation. Note that negative spectral types mean K types (K7 $=-1$, K5 $=-2$). The plot shows that the TiO5 index works very well as a spectral type indicator from K5 to M6. For earlier types the TiO feature disappears, and for types later than M6 the TiO feature saturates.

Available for almost all candidates later than K5 (see above), photometry was only needed for the selection of systems on the G/K boundary and earlier. All of these candidates are bright enough to have sufficiently accurate photometry in the Tycho-2 catalogue [Høg et al., 2000], the Tycho Double Star Catalogue [TDSC; Fabricius et al., 2002], or the Tycho catalogue [Høg et al., 1997]. Where there has been a need to convert between Tycho and Johnson photometry, relationships from Høg et al. [2000] for unreddened main sequence stars have been used.

\[
V_J = V_T - 0.990(B_T - V_T) \tag{3.3}
\]

\[
(B_J - V_J) = 0.850(B_T - V_T). \tag{3.4}
\]

Although not used for the selection, Johnson $BV$ photometry has been compiled for the M type primary stars [Weis, 1996, 1991, Bessel, 1990, Rodgers and Eggen, 1974, Eggen, 1974, 1980, Leggett, 1992, Henry et al., 2006, and other catalogues already mentioned]. This has been used to produce colour magnitude diagrams (e.g., Fig. 3.3) and to compute luminosities from absolute $V$ magnitudes (§3.5.3).
Table 3.1: Comparison of Gray and PMSU spectral types for stars within the UNS K type volume \((d < 15.7\) pc\), ordered by ascending distance. Six instances where the two sources conflict as to whether a star is K or M type are highlighted in bold. All of these occur outside the UNS M type volume. Spectral types from the Hipparcos catalogue are also shown. Note that none of the stars are classified as M type in the Hipparcos catalogue due to Gray & Corbally's selection criteria.

<table>
<thead>
<tr>
<th>UNS</th>
<th>GJD</th>
<th>PMSU</th>
<th>Gray</th>
<th>HIP</th>
<th>UNS</th>
<th>GJD</th>
<th>PMSU</th>
<th>Gray</th>
<th>HIP</th>
</tr>
</thead>
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<td>K7 V</td>
<td>K7V</td>
<td>F012</td>
<td>1075</td>
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<td>M4 V</td>
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<td>M3 V ke</td>
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<td>79</td>
<td>K7</td>
<td>K9 V k</td>
<td>K5/M0V</td>
<td>K120</td>
<td>200</td>
<td>K7</td>
<td>K3 V</td>
<td>K3V</td>
</tr>
<tr>
<td>K051</td>
<td>208</td>
<td>K7</td>
<td>M0.5 V ke</td>
<td>K7</td>
<td>K121</td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>103</td>
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<td>K8 V</td>
<td>K7V...</td>
<td></td>
<td>K125</td>
<td>750</td>
<td>K7.0</td>
<td>M0 V (k)K9V...</td>
<td></td>
</tr>
<tr>
<td>1097</td>
<td>M3</td>
<td>M3 V</td>
<td></td>
<td></td>
<td>K127</td>
<td>40</td>
<td>K5</td>
<td>K7-V kK5V</td>
<td></td>
</tr>
</tbody>
</table>

3.3.4 Astrometry

Accurate positions and proper motions were necessary both for matching entries in the various catalogues used, and for finding common proper motion companions. Where possible, astrometry from Salim and Gould [2003], Gould and Chanamé [2004], Deacon et al. [2005], Subasavage et al. [2005a,b], Finch et al. [2007], Henry et al. [2006], Jao et al. [2005] have been used. For stars not included or not resolved in these, astrometry from the TDCS; Tycho-2; the Tycho Reference Catalogue [TRC; Høg et al., 1998]; Tycho; Bakos et al. [2002]; the Positions and Proper Motions catalogue [PPM; Röser and Bastian, 1991, 1993, Röser et al., 1994]; or the CCDM (in order of decreasing preference) has been used.
3.4 Components of Multiple Systems

Several steps have been undertaken to maximise the accuracy of the selection of components in multiple systems.

Using the database which was constructed for the purposes of the target selection, searches were performed for stars with common proper motion to candidate targets. This not only yielded secondary stars which had not been previously identified during the selection process, but also showed some candidates to be secondaries of other stars. In cases where common proper motion companions have independent parallax measurements, these have been checked for compatibility. Other common proper motion companions have been identified from literature [e.g., Gould and Chamache, 2004, Makarov et al., 2008].

A complete check of all components listed in the CCDM as being in the CCDM systems of the targets was performed. In many cases components listed in the CCDM are not physically associated (e.g. do not have common proper motion) with the target system. Many CCDM components have cross-identifications with other catalogues, so determining whether they have common proper motion is straightforward. For those without cross-identifications, or without accurate astrometry in other catalogues, only the astrometry in the CCDM could be used.

The process for determining system membership of CCDM components consisted of an automated search for components using the 2MASS Point Source Catalogue [PSC; Cutri et al., 2003], and the Tycho/Tycho-2 catalogues, followed by manual inspection of 2MASS and Schmidt survey images, as well as comparison with the Washington Double Star catalogue [WDS; Mason et al., 2010] in many cases. CCDM components found not to be comoving with the target systems, or not identified at all, are not included in the sample presented here, but are listed in comments in Table 3.9 and Table 8 of Phillips et al. [2010].
3.5 Sample Properties

Overall properties of the subsamples are presented in Table 3.2, including the numbers of systems containing stars with detected planets and debris discs. Fig. 3.5 shows the distribution of systems on the sky for the whole sample and for the DEBRIS survey targets, which are mostly at high galactic latitudes due to the cirrus confusion cut.

Table 3.2: Summary of spectral type subsample properties. $d_{\text{max}}$ and $N_{\text{tot}}$ are the maximum distance and number of stars in each subsample. $\rho$ is the volume number density of systems, $\rho = N_{\text{tot}}/d_{\text{max}}^3$. $\text{Med}(T_{\text{eff}})$ is the median $T_{\text{eff}}$, and $\sigma_{T_{\text{eff}}}$ is the standard deviation of $T_{\text{eff}}$ within each subsample. $N_{\text{planet}}$ is the number of systems where one or more stars are listed as planet hosts in the exoplanet.eu database (27 July 2009). $N_{\text{debris}}$ is the number of systems containing a currently detected debris disc (or other indistinguishable IR excess) as indicated by any of Rhee et al. [2007], Beichman et al. [2006a], Su et al. [2006], Trilling et al. [2008]. $N_{\text{SUNS}}$ and $N_{\text{DEBRIS}}$ are the numbers of systems included in the SUNS and DEBRIS surveys respectively.

<table>
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<tr>
<th>SpT</th>
<th>$d_{\text{max}}$ (pc)</th>
<th>$N_{\text{tot}}$</th>
<th>$\rho$ (pc$^{-3}$)</th>
<th>Med($T_{\text{eff}}$) (K)</th>
<th>$\sigma_{T_{\text{eff}}}$ (K)</th>
<th>$N_{\text{planet}}$</th>
<th>$N_{\text{debris}}$</th>
<th>$N_{\text{SUNS}}$</th>
<th>$N_{\text{DEBRIS}}$</th>
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<td>A</td>
<td>45.5</td>
<td>130</td>
<td>0.0014 $\pm$ 0.0001</td>
<td>8133</td>
<td>748</td>
<td>2</td>
<td>24</td>
<td>100</td>
<td>83</td>
</tr>
<tr>
<td>F</td>
<td>24.1</td>
<td>130</td>
<td>0.0093 $\pm$ 0.0008</td>
<td>6300</td>
<td>343</td>
<td>6</td>
<td>21</td>
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<td>94</td>
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<td>G</td>
<td>21.3</td>
<td>125</td>
<td>0.0129 $\pm$ 0.0012</td>
<td>5628</td>
<td>249</td>
<td>13</td>
<td>10</td>
<td>100</td>
<td>89</td>
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<tr>
<td>K</td>
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<td>499</td>
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<td>5</td>
<td>100</td>
<td>91</td>
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<td>117</td>
<td>0.1855 $\pm$ 0.0171</td>
<td>3175</td>
<td>288</td>
<td>5</td>
<td>1</td>
<td>100</td>
<td>89</td>
</tr>
<tr>
<td>Total</td>
<td>629</td>
<td>31</td>
<td>0.1585 $\pm$ 0.0171</td>
<td>3175</td>
<td>288</td>
<td>5</td>
<td>1</td>
<td>100</td>
<td>89</td>
</tr>
</tbody>
</table>

3.5.1 Completeness

In Fig. 3.6, the number of systems as a function of distance for each of the subsamples is shown. The F, G and K subsamples very closely follow a cubic law, indicating that it is justified to assume they are isotropically and homogeneously distributed in the relevant volumes and that there are no selection effects as a function of distance. For the M subsample there is almost certainly incompleteness at distances beyond $\sim 6$ pc [see, e.g., Henry et al., 1994] which will mostly affect the latest type stars. The deviation of the A subsample from the cubic law is likely a combination of a slight lack of systems towards the Galactic poles at the largest distances, and correlation between system positions due to the young age of A stars.
Figure 3.5: Distribution of systems in equatorial coordinates (ICRS, epoch J2000). Spectral types are coloured as in Fig. 3.3. Top: all systems, with the SUNS declination limits of +80° and −40°, and the Galactic plane shown. Middle: DEBRIS target systems with the Galactic plane shown. Bottom: systems excluded from DEBRIS due to predicted confusion cut (targets excluded for other reasons not shown).
3.5. SAMPLE PROPERTIES

Figure 3.6: Number of included systems in each subsample as a function of distance ($d_{\text{max}} = 8.58, 15.6, 21.3, 24.1, 45.5$ pc for M−A respectively). For comparison the line $N = N(d_{\text{max}}) \left( \frac{d}{d_{\text{max}}} \right)^3$ is shown. Note that the F,G,K subsamples fit well indicating no completeness trend with distance. The M subsample is likely incomplete beyond $\sim 6$ pc.
3.5.2 Effective Temperatures

As the sample was split into subsamples based on spectral class, it was expected to have a good coverage of effective temperature of primary stars from approximately 2500 to 10000 K (M7-A0 types). Fig. 3.7 shows the distribution of $T_{\text{eff}}$ for primary stars in the sample in 500 K bins. The colours in the plot indicate the contributions from the five A-M subsamples. For A-K stars, $T_{\text{eff}}$ was computed from $(B_T - V_T)$ using a fit to $T_{\text{eff}}$ for stars in the sample from Gray et al. [2003, 2006]. $(B_T - V_T)$ was chosen as opposed to other photometric colours such as $(B_1 - V_1)$ or $(V_T - K_s)$, as accurate homogeneous $B_T$ and $V_T$ photometry that is resolved down to separations of $< 0.5''$ is available for almost all of the A-K primaries from the Tycho-2 and Tycho Double Star (TDSC) catalogues. The fit of $(B_T - V_T)$ to Gray et al.’s $T_{\text{eff}}$ values is shown in Fig. 3.8. A fourth order least-squares polynomial fit was obtained:

$$
T_{\text{eff}}/K = (9646.15 \pm 37.6) - (10018.4 \pm 354.4)(B_T - V_T) + (9056.19 \pm 963.2)(B_T - V_T)^2
- (4424.10 \pm 950.5)(B_T - V_T)^3 + (807.378 \pm 302.8)(B_T - V_T)^4.
$$

(3.5)

This agrees well with a relationship from Ramírez and Meléndez [2005] with [Fe/H] = 0.0, within their range of validity of $0.344 < (B_T - V_T) < 1.715$. The RMS of residuals is 150.7 K for 302 stars, which is higher than that of Ramírez and Meléndez [2005] (104 K for 378 stars), as a larger temperature range is covered. [Fe/H] has not been used as a fit parameter, and interstellar reddening has not been accounted for (although this should be almost negligible for this nearby star sample).

For M type stars, $T_{\text{eff}}$ was determined simply from the adopted spectral type using values from Reid and Hawley [2005]. The above photometric fit for A-K stars included a point representative of a typical M0 type star at $(B_T - V_T) = 1.70$, $T_{\text{eff}} = 3800$ K, to make the fit consistent with the M star temperatures at the K/M boundary.

The peak in the $T_{\text{eff}}$ distribution at approximately 5700 K is due to the G and F spectral types covering a narrow range in $T_{\text{eff}}$. Indeed, in retrospect, there would be justification for treating F and G types as a single spectral type sample.

The effective temperatures of the primary stars are given in Tables 3.6 and 3.7, and Tables 5 and 6 in Phillips et al. [2010].
Figure 3.7: Histogram of number of primaries in 500 K $T_{\text{eff}}$ bins. Contributions from each spectral type subsample are shown in colour. For A-K stars $T_{\text{eff}}$ was derived from $(B_T - V_T)$ (or $(B_J - V_J$ in a few cases where Tycho photometry was not available) using a polynomial fit against $T_{\text{eff}}$ values from Gray et al. [2003, 2006] (see Fig. 3.8). $(B_T - V_T)$ was used in preference to the more accurate temperature indicator $(V - K_s)$, as components are resolved at very small separations in Tycho-2/TDSC photometry. For M type stars $T_{\text{eff}}$ was derived from the adopted spectral type using $T_{\text{eff}}$ values from Reid and Hawley [2005].

Figure 3.8: Gray et al. [2003, 2006] $T_{\text{eff}}$ vs. $(B_T - V_T)$ for primary stars in the sample, with 4th order polynomial fit. This fit was used to generate $T_{\text{eff}}$ values for all A-K primaries for Fig. 3.7. A point for a typical M0 type star at (1.70, 3800) was added to the fit to make it tie in with $T_{\text{eff}}$ values for M type stars derived from spectral types using relationships in Reid and Hawley [2005].
3.5.3 Luminosities & Hertzsprung-Russell Diagram

Luminosities for the primary stars were determined from absolute Johnson V band magnitudes \( M_V \) using bolometric corrections from Bertone et al. [2004], interpolated by a polynomial fit (Fig. 3.9). For each star, \( M_V \) was determined using the adopted system distances and either \( V_J \) magnitudes (M type stars) or Tycho photometry (A-K type stars):

\[
M_V = \begin{cases} 
V_J \\
V_T - 0.090(B_T - V_T) 
\end{cases} - 5 \log_{10}(d/\text{pc}) + 5. 
\tag{3.6}
\]

The log luminosity was then given by,

\[
\log(L/L_\odot) = -0.4(M_V + BC_V - 4.83), \tag{3.7}
\]

where \( M_V(\odot) \) is taken to be 4.83\(^3\). The resulting plot of \( \log(L/L_\odot) \) vs. \( T_{\text{eff}} \) is shown in Fig. 3.10. A least-squares fit of a polynomial was obtained to the points in Fig. 3.10,

\[
\log(L/L_\odot) = - (46.1621 \pm 5.261) + (411.996 \pm 411.996) \left( \frac{T_{\text{eff}}}{10^4 \text{K}} \right)
- (3114.78 \pm 638.9) \left( \frac{T_{\text{eff}}}{10^4 \text{K}} \right)^2
+ (3114.78 \pm 638.9) \left( \frac{T_{\text{eff}}}{10^4 \text{K}} \right)^3
- (3379.31 \pm 813.8) \left( \frac{T_{\text{eff}}}{10^4 \text{K}} \right)^4
+ (1891.10 \pm 535.6) \left( \frac{T_{\text{eff}}}{10^4 \text{K}} \right)^5
- (427.443 \pm 142.7) \left( \frac{T_{\text{eff}}}{10^4 \text{K}} \right)^6. \tag{3.8}
\]

The RMS of residuals is 0.19 in \( \log L \), or 53% in \( L \). This fit approximates the average main sequence for the solar neighbourhood in the spectral type range M7 to A0.

The distribution of primary star luminosities is shown in Fig. 3.11. As for the \( T_{\text{eff}} \) distribution (Fig. 3.7), there is a peak around G/F type.

\(^3\text{http://nssdc.gsfc.nasa.gov/planetary/factsheet/sunfact.html}\)
3.5. SAMPLE PROPERTIES

Figure 3.9: Bolometric corrections from Bertone et al. [2004] as a function of $T_{\text{eff}}$ for $T_{\text{eff}} < 12000$ K. The line shows a 5th order polynomial fit to these points, with RMS residual of 0.014 m.

Figure 3.10: HR diagram for UNS primary stars with polynomial fit.
Figure 3.11: Histogram of number of primaries in 0.5 dex $\log(L/L_\odot)$ bins. Contributions from each spectral type subsample are shown in colour. The value of $\log(L/L_\odot)$ for each star was computed from Equs. (3.6) and (3.7), using bolometric corrections from Bertone et al. [2004].
3.5.4 Resolved Binary Separations

For systems with resolved astrometry for multiple components, the distributions of the logarithm of the projected physical separations (in AU) for the spectral type subsamples are shown in Fig. 3.12. Where there are multiple resolved components in a system, the separations of all components from the primary are included. The minimum separation at which resolved astrometry (either positions in e.g., Tycho-2 or the TDSC, or a separation vector in the CCDM) is generally available and of reasonable reliability is approximately 0.5–1″. The projected physical separations corresponding to 1″ at the maximum distance of systems in the subsamples are shown by vertical lines in Fig. 3.12. Incompleteness will be significant below these separations.

It is interesting to note that the separation distributions are very similar for A, F, G and K subsamples, except perhaps for a slight lack of G type systems with separations of order 100 AU. For the M type subsample, however, there is generally a distinct lack of resolved multiple systems. As the secondaries are necessarily of later spectral type than the primaries, the current incomplete census of late M type stars and brown dwarfs may partly explain the lack of common proper motion companions to M type stars.
Figure 3.12: Histograms of the logarithm of projected physical binary separations for the spectral type subsamples. The vertical lines correspond to $\rho = 1.0''$ at the maximum distance in each subsample. To the left of these lines there is likely to be significant incompleteness due to a lack of resolved astrometry in the catalogues which have been used. The numbers of resolved secondaries in the five subsamples are: 36, 46, 36, 59, 45 (M–A).
3.6 Catalogue

Table 3.3 lists the references used throughout this chapter and in the other tables. Tables 3.4, 3.5, 3.6 and 3.7 define the sample, and give information used in the selection process. Each system is given an identifier of the form XNNN where X is the spectral class (subsample) and NNN is a zero-padded running number increasing with distance in each subsample. These identifiers are referred to by the acronym UNS, standing for Unbiased Nearby Stars, as in the SUNS survey name. In this thesis only sample tables containing the first few entries for each subsample are given. Full tables are available in the online version of Phillips et al. [2010].

The choice of name for components is generally in order of preference: HD, HIP, GJ, LHS, NLTT, TYC, PPM, CCDM, other catalogue name, 2MASS. For stars in multiple systems, the first identifier in that order which uniquely identifies each component is used. Where components are not resolved in any catalogues which have been used, a single entry is given.

Table 3.4 lists system properties, including the name of the primary star, the adopted distance and uncertainty, and whether the system is included in the SUNS and DEBRIS surveys.

Table 3.5 lists the components of systems which are resolved in at least one of the catalogues which have been used. Positions and proper motions are given, as well as the approximate separation from the primary where this is larger than 1". Where two references are listed for a component, the proper motion has been copied from another component in the system, and in several cases the position is computed using a relative position from the CCDM combined with the position of another component.

Tables 3.6 and 3.7 list the properties of primary stars in systems, which were used for selection in spectral type and luminosity (spectral type, photometry), and/or in the plots in this chapter (photometry, effective temperatures). Table 3.6 contains the A-K type primaries with Tycho photometry, and effective temperatures from Gray et al. [2003, 2006] and computed from $(B_T - V_T)$. For the few very bright stars where Tycho photometry is saturated, values converted from Johnson $B, V$ photometry are given. Table 3.7 contains the M type primaries with spectral types, Johnson $B, V$ photometry, and effective temperatures computed from the spectral type.

Table 3.8 gives cross identifications for system components in several common catalogues, and Table 3.9 gives comments on various specific systems. Table 3.9 includes notes for systems where there are unresolved components, or there are components listed in catalogues which have been determined to not be physically associated with the system.
Table 3.3: Reference abbreviations used in the text and tables. CDS is Centre de Données astronomiques de Strasbourg. For HIPpar the data on the CDROM published with the book was used, as it had not been added to the CDS at the time.

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<th>Abbr.</th>
<th>CDS catalogue(s)</th>
<th>Reference</th>
</tr>
</thead>
<tbody>
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<td>2MASS Point Source Catalogue [Curi et al., 2003]</td>
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<td>Bright Star Catalogue, 5th Revised Ed. [Hoffleit and Warren, 1995]</td>
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<td>CCEM</td>
<td>I/274</td>
<td>Catalogue of Components of Double &amp; Multiple stars [Domenge and Ny, 2002]</td>
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<td>I/239</td>
<td>Hipparcos Main Catalogue [Perryman et al., 1997]</td>
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<td>I/239</td>
<td>Hipparcos General Notes [Perryman et al., 1997]</td>
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<td>I/311*</td>
<td>Hipparcos, the New Reduction of the Raw Data [F. van Leeuwen, 2007]</td>
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### Table 3.4: System information

System ID, primary star name, adopted distance and uncertainty ($d = 1/\pi \pm \alpha_d/\pi^2$), number of parallax measures used, parallax references (see Table 3.3), predicted total confusion noise for point source observed with Herschel’s PACS instrument at 100 µm, which surveys system is included in (S: SUNS, D: DEBRIS). The distance for UNS G001 ($\alpha + \text{Proxima Centauri}$) does not include any contribution from Proxima, as the parallax difference from the primary is significant. This example table contains the first six systems in each subsample.

<table>
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<th>Primary</th>
<th>$d$ (pc)</th>
<th>$N_m$</th>
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<td>2.543 ± 0.004</td>
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<td>0.52</td>
<td>S D</td>
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<tr>
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<td>2.676 ± 0.019</td>
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<td>0.52</td>
<td>S D</td>
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<td>0.53</td>
<td>S</td>
</tr>
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<td>0.70</td>
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<td>HIPn, YPC</td>
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<tr>
<td>G003</td>
<td>HD 185444</td>
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<td>HIPn, YPC</td>
<td>1.53</td>
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<td>G005</td>
<td>HD 26794</td>
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<td>2</td>
<td>HIPn, YPC</td>
<td>0.52</td>
<td>D</td>
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<tr>
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<td>6.708 ± 0.021</td>
<td>2</td>
<td>HIPn, YPC</td>
<td>0.53</td>
<td>S D</td>
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<tr>
<td>F001</td>
<td>HD 61421</td>
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<td>YPC, soe69</td>
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<tr>
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<td>YPC, HIPn</td>
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<td>HD 98231</td>
<td>8.368 ± 0.055</td>
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<td>HD 36393</td>
<td>8.926 ± 0.014</td>
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<td>S D</td>
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<tr>
<td>A001</td>
<td>HD 48915</td>
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<td>HIPn, YPC</td>
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<td>HD 187642</td>
<td>5.125 ± 0.014</td>
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<tr>
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<td>HD 102647</td>
<td>11.011 ± 0.003</td>
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<td>HIPn, YPC</td>
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<td>HIPn, YPC</td>
<td>0.54</td>
<td>S D</td>
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87
**Table 3.5:** Component names, positions and proper motions: ‘Pri’ column contains ‘P’ for primary component; ‘Refs’ column gives the reference for position and proper motion; ρ column gives separation of component from primary if larger than 1.0′. ρ should be considered approximate, and time variable for smaller separations (of order 100 AU or less). It is advised to check orbital solutions to find relative positions for a particular epoch. This example table contains the first five systems in each subsample.

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<th>µ_δ (mas/yr)</th>
<th>Refs</th>
<th>ρ (arcsec)</th>
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88

**CHAPTER 3. TARGET SELECTION FOR THE SUNS AND DEBRIS SURVEYS**
### Table 3.6: A-K primary spectral types, Tycho photometry and effective temperatures: spectral type and reference; Tycho $B_T,V_T$ magnitudes with standard errors and reference; $T_{\text{eff}}$ from gray03 or gray06; $T_{\text{eff}}$ computed from Tycho photometry (see §3.5.2). Where TYC2 and TDSC give the same $B_T,V_T$ and uncertainties, TYC2 is given as the reference here. In six cases Tycho photometry is not available, so values converted from Johnson $B,V$ magnitudes using Eqs. (3.3) and (3.4) are given. This example table contains the first eight systems in each subsample.

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<th>UNS ID</th>
<th>Primary name</th>
<th>SpT</th>
<th>Ref.</th>
<th>$V_T$ (mag)</th>
<th>$B_T$ (mag)</th>
<th>$T_{\text{eff},G}$ (K)</th>
<th>$T_{\text{eff},T}$ (K)</th>
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<td>HD 22049</td>
<td>K2 V (k)</td>
<td>gray06</td>
<td>3.814 ± 0.009</td>
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<td>5005</td>
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<td>HD 201091</td>
<td>K5 V</td>
<td>gray06</td>
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<td>6.711 ± 0.014</td>
<td>TYC2 4401</td>
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</tr>
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<td>HD 209100</td>
<td>K4 V (k)</td>
<td>gray06</td>
<td>4.826 ± 0.009</td>
<td>6.048 ± 0.014</td>
<td>TYC2 4654</td>
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</tr>
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<td>K7.0</td>
<td>haw95</td>
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<td>8.476 ± 0.017</td>
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<td></td>
</tr>
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<td>HD 88230</td>
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89
### Table 3.7: M type primary spectral types, effective temperatures and Johnson $B,V$ photometry:
spectral type and reference, $T_{\text{eff}}$ determined from spectral type, Johnson $V$ magnitude and references, 
Johnson $(B-V)$ colour and references. Where multiple references are given for the photometry, the value given here is the mean of the referenced values. This example table contains the first 30 systems.

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Table 3.8: Component cross identifications with common catalogues: system ID, CCDM ID and component, Henry Draper (HD/HDE) ID, Gliese & Jahrreiss (CNS3) ID, Luyten Half Second ID, New Luyten Two Tenths ID (record number in original NLTT), Harvard Revised (BSC5) ID, Positions and Proper Motions ID, Hipparcos ID, Tycho ID, Tycho Double Star Catalogue ID and component, Bonner Durchmusterung ID, Cordoba Durchmusterung ID, Cape Photographic Durchmusterung ID, Yale Parallax Catalogue (PLX) ID, 2MASS Point Source Catalogue ID (these are determined by simple cone search and may not be reliable in some cases). This example table contains the first three systems in each subsample.

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3.6 CATALOGUE
Table 3.9: Notes for specific systems

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<th>Note</th>
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<tr>
<td>M009</td>
<td>Triple system. A C components are very close binary, B component orbits AC (Delfosse et al., 1999)</td>
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<td>M011</td>
<td>CCDM lists a third component (CCDM 00184+4:401 C), but this is not associated, CCDM 00184+4:401 C is in TYC 2794-1388-1</td>
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<td>M018</td>
<td>CCDM lists seven other components (CCDM 22281+5:741 C, D, E, F, G, H, I), but these are not associated, CCDM 22281+5:741 C is clearly visible in 2MASS image (22.38 10.42 + 57 42 44.0), but is not in the PSC, CCDM 22281+5:741 E is 2MASS 22281+5:74418, CCDM 22281+5:741 I is 2MASS 22280+5:74226, identification of CCDM 22281+5:741 G is uncertain, CCDM 22281+5:741 I is HD 233209</td>
</tr>
<tr>
<td>M024</td>
<td>CCDM lists a secondary (CCDM 17365+6:622 A, HD 160661), but this is not associated</td>
</tr>
<tr>
<td>K001</td>
<td>+ é Erick is not included in DEBRIS, as it is being observed by a Guaranteed Time Project</td>
</tr>
<tr>
<td>K002</td>
<td>PPM 80047 - FK5 703 is not a component but is the system photocentre, CCDM lists four other components (CCDM 21069+3:844 C, D, E, F), but these are not associated, CCDM 21069+3:844 C is BD +38 4345, CCDM 21069+3:844 D is BD +38 4342, CCDM 21069+3:844 E is BD +38 4349, CCDM 21069+3:844 F is TYC 3126-3076-1</td>
</tr>
<tr>
<td>K003</td>
<td>+ Indi B is a brown dwarf binary</td>
</tr>
<tr>
<td>K005</td>
<td>CCDM lists two other components (CCDM 10114+4:492 B, C), but these are not associated, CCDM 10114+4:492 B is HD 233714, CCDM 10114+4:492 C is HD 233713</td>
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<tr>
<td>K006</td>
<td>CCDM lists two other components (CCDM 04153+0:790 D, E), but these are not associated, CCDM 04153+0:790 D is 2MASS 04153226+0:790274</td>
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<tr>
<td>G001</td>
<td>Parallax distance: 1.301 ± 0.001 pc (YPC, HIP, Jan09)</td>
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<tr>
<td>G002</td>
<td>+ Ceti is not included in DEBRIS, as it is being observed by a Guaranteed Time Project</td>
</tr>
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<td>G003</td>
<td>CCDM lists a secondary (CCDM 01441+1:557 B), but this is not associated, CCDM 01441+1:557 B is 2MASS 01440770+1558304</td>
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<td>CCDM lists a secondary (CCDM 10322+6:641 B), but this is not associated, CCDM 10322+6:641 B is TYC 4446-2115-1</td>
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<tr>
<td>G005</td>
<td>CCDM lists six other components (CCDM 00091+5:740 C, D, E, F, G), but these are not associated, CCDM 00091+5:740 C is 2MASS 00083553+5741355, CCDM 00091+5:740 D is 2MASS 00005344+5751559, CCDM 00091+5:740 E is BD +57 155, CCDM 00091+5:740 F is TYC 3063-1486-1, CCDM 00091+5:740 G is BD +66 129, CCDM 00091+5:740 H is HD 256533</td>
</tr>
<tr>
<td>G006</td>
<td>CCDM lists two other components (CCDM 14513+1:906 C, D), but these are not associated, CCDM 14513+1:906 C is 2MASS 14512179+1:907087, CCDM 14513+1:906 D is 2MASS 14511264+1:906463, TYC proper motion is likely inaccurate</td>
</tr>
<tr>
<td>F001</td>
<td>Proper glue has DA white dwarf secondary (G J 280 B), CCDM 07303+0:514 C, D, E are not associated, CCDM 07303+0:514 C is 2MASS 07302181+0516077, CCDM 07303+0:514 D is not in 2MASS PSC, but has three entries in 2MASS Survey Point Source Reject Table, CCDM 07303+0:514 E is TYC 187-804-1</td>
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<td>F002</td>
<td>Spectroscopic binary (BSCD 1038), CCDM lists two wide binaries (CCDM 18211+7:245 B, C), but these are not associated, CCDM 18211+7:245 B is TYC 4437-465-1, CCDM 18211+7:245 C is 2MASS 18210008+7436092</td>
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<td>CCDM lists a secondary (CCDM 04499+0:657 B), but this is not associated, CCDM 04499+0:657 B is TYC 06-137-1</td>
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<td>F006</td>
<td>CCDM lists a tertiary component (CCDM 04445-2226 C) but this is not associated, CCDM 04445-2226 C is CPD -22 883, 2MASS 04442700-223272, ρ, θ in CCDM are suspect</td>
</tr>
<tr>
<td>A001</td>
<td>CCDM and WDS list a third component (CCDM 00451-1643 C) orbiting Sirius B, but this is not well confirmed, and is not included here, CCDM lists a wide secondary (CCDM 00451-1643 D), but this is not associated, CCDM 00451-1643 D is visible in 2MASS image (06 45 11.72 + 66 41 48.7), but is not in the PSC</td>
</tr>
<tr>
<td>A002</td>
<td>Altair is included in DEBRIS, despite just missing confusion cut (1.24 vs. 1.20 mJy/beam), CCDM lists two other components (CCDM 15098+0:852 B, C), but these are not associated, CCDM 15098+0:852 B is 2MASS J5093473+0853019, CCDM 15098+0:852 C is 2MASS J5093533+0851129 (ρ in CCDM is slightly too large)</td>
</tr>
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<td>A003</td>
<td>Vega is not included in DEBRIS, as it is being observed by a Guaranteed Time Project, CCDM lists four other components (CCDM 18309+3:847 B, C, D, E), but these are not associated, CCDM 18309+3:847 B is PPM 81557 and is visible in 2MASS images, but is flagged as a persistence artifact, CCDM 18309+3:847 C is clearly visible in 2MASS images (38 36 50.24 +38 46 44.6), but is not in the PSC, CCDM 18309+3:847 D is clearly visible in 2MASS images (38 36 51.52 +38 47 10.7), but is not in the PSC, CCDM 18309+3:847 E is 2MASS 18370125+3848126</td>
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92
3.7 Summary

This chapter has presented a sample of 629 nearby star systems, composed of five volume limited subsamples of 118–130 systems with main sequence primaries of spectral type A, F, G, K and M. These systems have each been assigned an identifier constructed from the spectral type (A, F, G, K, M) followed by a number which increases with distance within each subsample. For instance, the α + Proxima Centauri system is G001 – the closest system to the Sun with primary of spectral type G.

The target lists of the SUNS and DEBRIS surveys for debris discs were derived from the sample presented here by applying simple cuts in declination and predicted cirrus confusion noise respectively. The volume limits of the subsamples were determined from the SUNS selection criteria, which required 100 systems of each spectral type in the declination range $-40^\circ < \delta < +80^\circ$. DEBRIS has a total of 446 targets.

Many resources were combined to facilitate the selection of the subsamples, with the aim of making the selection based on the most up to date, complete and reliable information available. Searches were performed, both in the database constructed for the sample selection and in the literature, for components of multiple systems. Although the completeness of the subsamples is as high as possible, there appears to still be significant incompleteness in the M type subsample at distances beyond $\sim 6$ pc. This is a known problem [Henry et al., 1994], and nearby M type stars are still regularly being discovered [e.g. Henry et al., 2006].

Optical photometry and spectral types have been used to determine effective temperatures and luminosities for the primary stars in the sample. The distributions of $T_{\text{eff}}$ and $\log(L/L_{\odot})$ (Figs. 3.7 and 3.11) both exhibit peaks corresponding to G/F type stars, although there is significant coverage of $2500 \text{K} < T_{\text{eff}} < 10000 \text{K}$ and $-3 < \log(L/L_{\odot}) < 2$. A polynomial relationship was fitted to $\log(L/L_{\odot})$ vs. $T_{\text{eff}}$, which was used for determining the sensitivity limits of debris disc surveys in the previous chapter.

The distributions of projected binary separations were examined. The distributions for the A–K type subsamples are similar, however, a general lack of resolved M type binary systems was found.

The high completeness, combined with relatively even and wide coverage of stellar temperature, luminosity and mass, will hopefully lend the sample presented here to other studies of the solar neighbourhood.
Chapter 4

Properties of Nearby A Type Stars with Spitzer

4.1 Introduction

This chapter presents a survey of a volume limited sample of 130 A type stellar systems at 24 and 70 μm using the Multiband Imaging Photometer and Spectrometer [MIPS; Rieke et al., 2004b,a] instrument on the Spitzer space telescope. The targets of this survey are the A type sample presented in Phillips et al. [2010] and Chapter 3 of this thesis.

This survey was motivated by two main factors: to provide complementary data for the SUNS and DEBRIS surveys described in Chapter 2 of this thesis, and to perform a detailed statistical survey of debris discs around an unbiased sample of A type stars (based on the UNS target list presented previously). A Spitzer Cycle-5 proposal was submitted in November 2007 [Phillips et al., 2008] to observe all A-K type UNS systems which did not already have MIPS photometry observations listed in the Reserved Observations Catalogue1 (v9.2). The list of targets comprised 1 G, 1 F and 49 A type systems. Very few F-K targets were not already observed or scheduled for observation, as a Cycle-3 survey [Koerner et al., 2006, 2010] was already observing all stars of K or earlier type within 25 pc.

The choice to focus on A type stars in this work was many-fold. Firstly, F-K type stars (considered to be ‘Sun-like’) have already been the subject of several surveys using MIPS, culminating in the study of Trilling et al. [2008], which includes ~300 stars, and the Koerner et al. survey mentioned above [first results presented in Koerner et al., 2010]. M type stars have not been studied in great numbers with Spitzer; however, this is with good reason, as the low

1http://ssc.spitzer.caltech.edu/warmission/prophit/roc/
stellar luminosities of $10^{-2}$ to $10^{-3}L_\odot$ mean that any discs will be too cold for MIPS to detect if their radii are much over 1 AU (e.g. Fig. 2.4). Gautier et al. [2007] detected no significant excesses from samples of 62, 41 and 20 nearby M-type stars observed at 24, 70 and 160 $\mu$m respectively. The observing time which would have been required to survey a large fraction of the UNS M-type sample at 70 $\mu$m to a depth likely to yield any new detections would have been unreasonable, and poorly justified given that Herschel’s longer wavelengths are more suitable for detecting colder discs.

The 24 and 70 $\mu$m MIPS bands are ideally suited to detecting and characterising dust around A type stars\(^2\). Temperatures of black bodies with peak frequency flux density, $F_\nu$, at 24 and 70 $\mu$m are approximately 200 and 70 K respectively (Eqn. (1.2)), corresponding to radii of 12 to 100 AU for black body grains around A type stars (Eqn. (1.13), with $L = 50 L_\odot$). This makes MIPS ideal for detecting and characterising the temperatures of dust at Kuiper Belt like distances around A type stars. For realistic dust grains the orbital distances probed can be significantly greater (e.g. Fig. 1.13 and Bonor and Wyatt, 2010), allowing MIPS to detect dust at physical distances of several hundred AU around A type stars.

The success of MIPS for detecting debris discs around A type stars has previously been demonstrated in Su et al. [2006], where a total of $\sim 160$ A type stars were examined, with typical debris detection rates of over 30% at both 24 and 70 $\mu$m. The sample of stars studied by Su et al. [2006] contains many potential biases. As the goal of Su et al. [2006] was to study evolution of debris, their targets were mostly chosen from stellar clusters and associations of known age. Su et al. [2006] also incorporated some field stars which were largely observed by other surveys which specifically targeted likely or known debris disc hosts. The work presented here aimed to test the generality of the debris detection rates identified by Su et al. [2006] by using a volume limited sample of stars.

In this work the rates of debris disc incidence were explored as a function of binarity and binary separation, metallicity and effective temperature. The effect of binary separation in particular was found to greatly influence the likelihood of detecting a debris disc with MIPS. This work has also discovered 12 previously unknown debris discs, including the most massive disc within the volume, and a second debris disc in a triple star system. For all detected discs the dust temperatures, masses and orbital radii have been determined from the MIPS photometry (limits on these parameters were determined where dust was only detected in one band), providing an unbiased and homogeneously determined set of disc parameters for A type stars within $\sim 45$ pc from the Sun.

\(^2\)The 160 $\mu$m MIPS band suffers from a serious near-IR light leak that introduces significant photometric uncertainty for bright stars, which combined with limited sensitivity and requiring the highest cryogen consumption of any MIPS detectors, makes the 160 $\mu$m band unsuitable for debris disc surveys.
4.2 Data Reduction

Due to the large number of images involved in this survey (over 100 in each band), and the desire for homogeneity, all observations were fully pipeline reduced, with no user intervention for individual observations. It was found that pipeline processed images available from the Spitzer Science Center (SSC) were sufficient, although an extra automated filtering stage was applied to 70 μm images as described below. The SSC pipeline produces two sets of products: level-1 *Basic Calibrated Data* (BCD) images of individual detector integrations; and level-2 post-BCD images, produced by mosaicking all the BCD images from an observation. The level-2 post-BCD mosaic images have been used here.

Photometry was obtained from the images by PSF fitting or aperture photometry using *daophot* [Stetson]. This was carried out manually, image by image, as it was necessary to inspect for companions or contaminating objects, confusion from galactic cirrus emission, and the quality of the PSF fit. Mostly PSF fitting was used, except for systems with significantly resolved dust emission where the PSF fit was seen to be poor. The choice to use PSF fitting photometry where possible was to maximise S/N (primarily beneficial at 70 μm), and to allow photometry to be obtained where there are multiple objects with separations comparable to aperture sizes.

To minimise calibration errors, the aperture radii used were the same as used for deriving the MIPS instrumental flux calibrations in Engelbracht et al. [2007] and Gordon et al. [2007]. *daophot* calibrates its PSF fit photometry to its aperture photometry when constructing the PSF model, so the PSF fit and aperture magnitudes are in the same system and required the same aperture corrections to be applied. Absolute calibration was achieved using the surface brightness calibration in the post-BCD images and published Vega zero point fluxes. Confirming the validity of the absolute calibration of the photometry produced here has been challenging, and is discussed in §4.4.

In the rest of this chapter the 24 μm and 70 μm MIPS bands are referred to as MIPS-24 and MIPS-70 respectively. Magnitudes obtained from observations in these bands are denoted [24] and [70] respectively. The choice to refer to band names and use magnitudes, rather than the more common practice in the far-IR of stating an ‘effective’ wavelength and flux densities (either instrumental, or ‘colour corrected’), is to highlight the fact that these photometric bands have considerable width, and interpreting the photometric measurements in terms of monochromatic flux densities is misleading.
CHAPTER 4. PROPERTIES OF NEARBY A TYPE STARS WITH SPITZER

4.2.1 MIPS-24

Instrument and Observing Modes

The 24 μm band of MIPS (henceforth referred to as MIPS-24) comprises a Boeing 128×128 Si:As blocked impurity band (BIB) photodetector array, combined with a single optical configuration which provides a pixel scale of ∼2.5″/pix and FoV of ∼5 × 5′. Routine observations were performed in either a stepped scan mode, or in one of two chopped modes. The two chopped modes provide even coverage of fields of approximately 3 × 3′ and 5 × 5′, and are referred to as small and large source modes respectively. Pointings in the small source mode all keep the target on the detector, and widely vary the position of the target on the detector (see Fig. 4.1). The large source mode places the target near the array centre in all on-source pointings, to allow the whole detector FoV to be observed to an even depth, and repeats the same pattern at an off-source position to obtain a sky frame. Individual exposures have durations of either 3, 10 or 30 s.

The vast majority of observations used in this survey were taken in the small source mode with 3 s exposures. The observations obtained specifically for this survey all consist of 2 repetitions of this observing mode, resulting in a total of 28 3 s integrations which are mosaicked into the post-BCD images. The limiting point-source flux and corresponding Vega magnitude for these observations are approximately 0.1 mJy and 12′′. The faintest primary stars in this survey have $F_{24} \approx 50$ mJy, [24] ≈ 5′′, so the photometric accuracy is dominated by calibration and repeatability uncertainties.

The MIPS-24 post-BCD images have a pixel scale of 2.45″/pix, and the pixel values are in units of MJy/sr. The conversion factor used to convert the instrumental units to MJy/sr, given in the FLUXCONV FITS header value, is the same value given by Engelbracht et al. [2007] for SSC pipeline reduced images.

Photometric Methods

Photometry was mostly obtained by PSF fitting, as this allows photometry to be obtained where there are multiple objects in close proximity. In addition to binary stars, the unbiased nature of this survey means that several targets have low galactic latitude and are hence in crowded fields. Very few of the target stars have extended emission at 24 μm, so PSF fitting was able to be used almost exclusively.

PSF fit photometry was performed on post-BCD images using DAOPHOT. This required a

3 Engelbracht et al. [2007] give a value of 4.54 × 10−2 MJy sr−1/(DN s−1) for images they have reduced with their own procedures, and state that instrumental values in SSC pipeline reduced images are higher by a factor of 1.6. The FLUXCONV value given in the post-BCD image FITS headers is 4.47 × 10−2 MJy sr−1/(DN s−1), which agrees with this.
4.2. **DATA REDUCTION**

![Figure 4.1: Summary of MIPS-24 small source observing mode, showing the target positions in the central $64 \times 64$ pixels of the array (diagram from the MIPS Instrument Handbook). Offsets are relative to the centre of the array. An additional frame of shorter exposure length is taken at pointings 1 and 8 for each entire observation, although these are not used by the SSC pipeline.](image)

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model PSF to be generated from observations. Observations of 12 stars with a similar range of fluxes to the survey as a whole were used. The images were tiled into a single image to allow *daophot* to work on them. An initial PSF model was used to subtract the PSF stars from their images to check for any deviation from being a point source, and to subtract other point sources within the PSF radius of the PSF stars. The resulting PSF, with radius of 35 pixels, is show in Fig. 4.2 (this is at the image resolution, although the actual PSF model which *daophot* uses is oversampled by a factor of 2).

Aperture photometry was also obtained using *daophot* as part of the photometry process. A single aperture size of $35''$ with $40-50''$ sky annulus was used, which is the same as used in Engelbracht et al. [2007] for the MIPS-24 flux calibration. Engelbracht et al. [2007] state an aperture correction factor of 1.082 (0.084") for this aperture configuration. When constructing the PSF model *daophot* calibrates the model to the aperture photometry of the PSF stars, so the PSF photometry is in the same system as the aperture photometry and hence requires the same aperture correction. This aperture photometry is used in place of PSF fit photometry for sources with resolved emission not attributable to point sources (e.g. resolved debris discs).

The conversion of the instrumental magnitudes produced by *daophot* to calibrated Vega magnitudes is covered in §4.4.

**Sky background correction**

During both PSF generation and fitting, *daophot* assumes that the sky background level is equal to the value determined from the sky annulus used for the aperture photometry. This value is higher than the true sky background by a factor proportional to the source flux, due
to a fraction of the energy in the the PSF falling in the sky aperture. This means that the model PSF has an incorrect zero level, and the level remaining in images after PSF fitting and subtraction is higher than the actual background level. By using large apertures this effect is minimised, and for isolated stars the effect is not a major concern as the photometry is unaffected. However, where multiple objects are to be fitted, it is necessary for the background levels to be realistic.

It was straightforward to correct the sky value from the aperture value to the real value by subtracting a fraction of the stellar flux measured from aperture photometry. The fraction was determined empirically by adjusting the value whilst repeating the PSF generation, fitting and subtraction, until an optimum value was found. The resulting correction used was,

\[
S_{\text{sky,corr}} = S_{\text{sky,uncorr}} - 2.75 \times 10^{-5} F_* ,
\]

where

\[
F_* = 10^{m_* - 25},
\]

and \( m_* \) is the uncorrected aperture magnitude produced by daophot using 35\(^\prime\) source radius and 40-50\(^\prime\) sky annulus. The surface brightnesses \( S_{\text{sky,corr}} \) and \( S_{\text{sky,uncorr}} \) are in units of MJy/sr, and \( F_* \) is in units of MJy 2.45\(^\prime\)^2/sr.

This correction to sky values was applied during PSF model construction, so no correction to PSF fitted magnitudes was necessary to bring them back in line with the aperture photometry (the PSF fit magnitudes were still calibrated against the same aperture magnitudes). The sky value correction was implemented as a small program which read the aperture photometry file.
produced by DAOPHOT and produced a modified version, so only a single extra command had to be run for each image during the photometry process. Fig. 4.3 shows the effect of the correction process.

![Image](https://example.com/image)

**Figure 4.3:** Effect of sky value on PSF subtracted images. Left: original image of 1.9" star (HD 40183); Centre: PSF subtracted image without sky correction for either PSF construction or fitting; Right: with sky correction for both PSF construction and fitting. The correction applied to the sky value for this image was 0.228 MJy sr\(^{-1}\).

### 4.2.2 MIPS-70

**Instrument and Observing Modes**

The 70 \(\mu\)m band of MIPS comprises a 32 \(\times\) 32 array of Ge:Ga photoconductor detectors and three optical configurations: coarse scale imaging with a nominal pixel scale of 9.8"/pix, fine scale imaging with a nominal pixel scale of 5.3"/pix, and a low resolution spectroscopy (SED) mode. Due to a cabling problem half of the array is not used, so the detector FoV is roughly 2.6' \(\times\) 5.2' and 1.4' \(\times\) 2.8' in coarse and fine scale modes respectively. The fine scale mode is intended to provide good sampling of the PSF at the expense of area coverage, sensitivity and on-source integration time. The coarse scale mode has generally been adopted for photometric observations, as it offers a factor of four better sensitivity, however, the coarse scale mode undersamples the 18" FWHM PSF. In this work photometry has only been obtained from observations performed in the coarse scale mode. Only five targets (Vega, ζ Lep, β Ari, β UMa and τ3 Eri) have only fine scale observations, hence developing reduction procedures for the fine mode was not considered worthwhile. For the targets observed only in fine-scale mode – all but one of which have a substantial excess above the photosphere – previously published photometry was taken from Su et al. [2006].

As for MIPS-24, there are two chopped observing modes – one with small offsets for compact
sources (< 1') which keeps the target on the array at all times, and one with larger offsets for covering targets up to 2' in size – as well as scan map modes. The vast majority of observations used for this project were taken in the compact source chopped mode, which is shown in Fig. 4.4. A small number of scan map observations have also been used (for Fomalhaut, α CrB and γ Oph). In all cases an observation consists of a series of integrations at stationary pointings, interspersed with integrations of flashes of an internal calibration source called a stimulator. The stimulator flashes are used to monitor the detector responsivity, which varies on a wide range of timescales. The SSC level-1 pipeline uses the stimulator flashes and other calibration information to produce flux calibrated images of the individual pointings referred to as Basic Calibrated Data images (BCDs). In addition to the standard BCDs, BCDs are provided which have had a temporal median filter applied between frames to remove latent images of the stimulator flashes and residual responsivity fluctuations. The SSC level-2 pipeline mosaics both sets of BCDs to produce filtered and unfiltered post-BCD images with a standard pixel scale of 4.0''/pix. The effects of this filtering on the post-BCD images are discussed below.

### Table 4.4

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<td>-7.5</td>
<td>-3.7</td>
</tr>
<tr>
<td>5</td>
<td>-9.0</td>
<td>7.4</td>
<td>11</td>
<td>-7.5</td>
<td>7.4</td>
</tr>
<tr>
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<td>-9.0</td>
<td>0.0</td>
<td>12</td>
<td>-7.5</td>
<td>0.0</td>
</tr>
</tbody>
</table>

**Figure 4.4:** Summary of MIPS-70 coarse scale compact source observing mode (diagram from the MIPS Instrument Handbook). Offsets are relative to the centre of the full array. Pointings 1 and 7 are only taken in the first repetition of an observation (typically 1-20 repetitions are used). Stimulator flashes occur before pointings 2, 6 and 8.

### Filtering Effects and Compensation

The MIPS-70 detectors have a very complicated temporal response to incident radiation, and the pixel sensitivity varies with time as well. The SSC level-1 pipeline calibrates out these effects as far as possible by using a combination of the stimulator flashes (effectively flatfields) and instrumental data. Unfortunately, signal due to residual detector response effects, partly introduced by the stimulator flashes themselves, still remains in the BCD frames and hence make it into the post-BCD mosaics produced from these. This results in an uneven background.
and elevated noise.

To reduce these effects for observations which are uninterested in extended emission on scales of the chop throw (~ 1′), the SSC pipeline also provides BCD frames, and mosaicked images produced from them, which have a temporal filtering applied. The filter acts per detector pixel, subtracting off a temporal median computed from the current frame and several prior and succeeding frames. This filtering is effective at removing the remaining slow response artifacts (latents), and also reduces sky structure from cirrus etc.

The downside of the filtering (ignoring observations where extended structure is important) is that the median which is subtracted from a pixel will be skewed when a bright astronomical source falls on the pixel in one of the chop positions which is used to compute the median. For the majority of images in this survey, the target is the brightest or only source visible (a quick calculation\(^4\) based on 70 μm galaxy number counts from Béthermin et al. [2010] shows that the number density of extragalactic sources with flux > 10 mJy is ~ 0.1 arcmin\(^{-2}\), so the over-subtraction from the filtering is seen as a dark band through the target with a length extending to twice the maximum chop offset. This is shown in the central image of Fig. 4.5 and in the top PSF shown in Fig. 4.6. The strength of the effect is a complicated function (non-linear due to the median filtering) of the source flux and background brightness, so it will translate to systematic uncertainties in photometry.

A workaround for this problem, described in Gordon et al. [2007], is to mask pixels in BCD frames which contain a significant fraction of the target’s flux before applying the median filtering. Implementing this here would, however, have required all the MIPS-70 observations to be re-reduced from the raw data, as the median filtering is implemented in the first level of the SSC pipeline. There is also the potential problem that ignoring all samples containing the target in the filtering could introduce other biases.

As a more convenient alternative to largely mitigate the negative effects of the temporal filtering, a simple algorithm for correcting the post-BCD images was developed. The algorithm operates on each column in a post-BCD image, subtracting off the median value of pixels in the column in bands symmetrically above and below the target position as shown in Fig. 4.5. The bands are chosen to cover the range of chop throws and lie outside the first airy ring of the target’s PSF. The bands used are rows 52–62 and 84–94. Within the range of rows 52–94 the median value is subtracted, and outwith this the median scaled by a Gaussian of \(\sigma = 10\) pixels centred at 52 or 94 pixels is subtracted. The maximum extent of these bands and the \(\sigma\) of the Gaussian were empirically chosen to give the best correction of the dark band artifact, and

\(^4\)Béthermin et al. [2010] shows \(dN/dS_{2,5} \approx 2000\ \text{gal}\ Jy^{-1}\ \text{str}^{-1}\) for \(10 \text{ mJy} < S_{70} < 1000\ \text{mJy}\). The number density of sources brighter than 10 mJy is thus \(f_{10 \text{ mJy}} \int_0^\infty \frac{dN}{dS} dS = 2000 \int_0^{10 \text{ mJy}} S^{-2,5} dS = \frac{2000}{7,5}(10 \text{ mJy})^{-1.5} = 1.3 \times 10^5 \text{str}^{-1} = 0.11 \text{ arcmin}^{-1}\).
don’t exactly correspond to values which would be chosen based on the maximum chop throw and PSF FWHM alone. The effect of this corrective filtering on the model PSF is shown in Fig. 4.6.

**Figure 4.5**: Effects of filtering on MIPS-70 images. Left: post-BCD image without temporal median filtering of BCDs, showing background structure due to slow response effects. Centre: post-BCD image with temporal median filtering of BCDs, showing vertical dark band artifact due to flux from the bright object skewing the median subtracted from pixels on which the source falls in individual BCD frames. Right: filtered BCD image after subtracting median horizontal profile computed from rows 52-62 and 84-94 (region marked by white lines), showing the dark feature is corrected for.

**Photometric Methods**

As for MIPS-24, photometry was performed by PSF fitting using DAOPHOT, except for sources with resolved emission for which aperture photometry was used. PSF fitting was especially beneficial for MIPS-70 due to the very low S/N of many of the observations. Aperture photometry was generally performed with an aperture of radius 35'' and a sky annulus of 30–65'', as used for the MIPS-70 flux calibration in Gordon et al. [2007]. For investigating the flux calibration, smaller apertures of 16'' with 18–39'' sky annulus (also listed in Gordon et al. [2007]) were also used as they yielded greater S/N photometry than the larger apertures.

Due to the lower dynamic range of MIPS-70, the S/N of individual usable observations was significantly lower than for MIPS-24. To allow a PSF model of good S/N to be constructed, a large number of observations of stars were used. The majority of observations used to construct
Figure 4.6: Left: MIPS-70 PSFs produced with DAOPHOT from filtered post-BCD images, without (top) and with (bottom) column medians subtracted from the images. Display cuts are -2.5% and +17.5% of the peak, with linear scale. Right: Average of the column median profiles subtracted from the images of HD 163588 and HD 180711 used for constructing the MIPS-70 PSF. The vertical lines show the typical $x$ centroid ($x = 57.2$ with pixel centres as integers) and 2.5 detector pixels (6.125 pixels here) either side, corresponding to the nod offset in the compact source observing mode (also closely coinciding with the first airy ring).

The PSF model were of two stars, HD 163588 and HD 180711, which were observed regularly as repeatability calibrators for MIPS-70 photometry. In total 268 observations of these stars – 144 of HD 163588 and 124 of HD 180711 – have been used. The only observations of these stars which were excluded were taken with a different dither pattern early in the Spitzer mission. This different observing mode results in different sized post-BCD images which caused complications for tiling the images to use with DAOPHOT. Other stars were also included in the model PSF construction, including other MIPS calibrators and stars from this survey which were single, uncontaminated, showed no extended emission, and had no significant flux excess. In total 327 post-BCD images were used, which were tiled to produce a single image which was used by DAOPHOT to construct a model PSF of radius 19 pixels as shown in Fig. 4.6. Due to the far lower S/N in the MIPS-70 images compared to MIPS-24, no correction to the sky values used in the PSF construction and fitting was found necessary.

The conversion of the instrumental magnitudes produced by DAOPHOT to calibrated Vega magnitudes is covered in §4.4.
4.3 Photosphere Photometry Prediction

To determine whether a star posses an excess at far-IR wavelengths, it is necessary to determine what the observed flux would be for the stellar photosphere alone. To do this requires extrapolation from photometry obtained at shorter wavelengths where excess due to dust can be considered negligible. The extrapolation requires a model for the photosphere flux distribution (absolutely calibrated spectrum with flux units).

As the spectra of stars become increasingly complicated and dependant on atmosphere parameters as wavelength decreases from the near-IR into the optical, extrapolation from the optical is very sensitive to the choice of model. The longer the wavelength of photometry used, the less the model dependence of the extrapolation, but conversely the greater the risk of contamination by dust excess.

The method used to predict far-IR photosphere photometry in this work was to perform a fit ($\chi^2$ minimisation) of observed optical and near-IR photometric colours to a grid of synthetic colours computed from model photosphere flux distributions. The best fit model was then used with observed optical and near-IR photometry to estimate MIPS-24 and MIPS-70 photometry.

The following subsections describe the model flux distributions which have been used; the production of the grid of synthetic colours; the photometric bands which have been used; calibration of the synthetic photometry (zero point offsets); the fitting process; the prediction of MIPS photometry; and complications due to chemical peculiarities, rapid rotation and binary systems. The production of the grid of synthetic colours is shown as a flowchart in Fig.4.7.

The effects of interstellar extinction have not explicitly been accounted for in this work, as the proximity of all of the targets to the Sun (maximum distance of 45 pc) places them within the Local Bubble [Vergely et al., 2010, and references therein], in which extinction is negligible. The effects of mild extinction are largely degenerate with the stellar parameters (primarily $T_{\text{eff}}$), so there is little merit in attempting to fit the extinction here.
4.3. PHOTOSPHERE PHOTOMETRY PREDICTION

Figure 4.7: Computing the grid of photometric colours for fitting and predicting photometry from model flux distribution grids.
4.3.1 Models

There are two widely used suites of stellar atmosphere modelling codes: ATLAS9 by Robert Kurucz and collaborators, and PHOENIX by Peter Hauschildt and collaborators. Grids of model photosphere flux distributions computed using both modelling codes are available. Munari et al. [2005] gives a useful summary of many of the available grids, showing the wavelength range, spectral resolution and parameter space covered ($T_{\text{eff}}$, log $g$, $[M/H]$ and others). A significant limitation of the PHOENIX based grids (e.g. Brott and Hauschildt [2005]) is the maximum effective temperature of 10000 K, which is cooler than some A type stars. Over the temperature range of A type stars there is good agreement between the PHOENIX and ATLAS9 based grids, for example Munari et al. [2005] compare several broadband photometric colours and show that all models give similar results for stars earlier than ~G type.

For this work Kurucz ATLAS9 based grids have been used to achieve homogeneous coverage of the required $T_{\text{eff}}$ range. Two grids have been combined to give the necessary wavelength coverage from optical to far-IR, and to give high enough spectral resolution to accurately fit medium-band optical photometry. A grid from Castelli and Kurucz [2003] provides spectral coverage from 90 nm to 160 $\mu$m with resolution of $\sim 0.1\lambda$ per sample shortward of 10 $\mu$m. A grid from Munari et al. [2005] provides coverage from 250 to 1050 nm with resolution of 1 $\lambda$ per sample ($\sim 10^{-4}\lambda$ per sample). Munari et al. [2005] computed their flux distributions using the atmosphere models of Castelli and Kurucz [2003] as input. The thresholds between the two sets of flux distributions, at 250 and 1050 nm, are conveniently placed with respect to the photometric bands which have been used. All the optical bands are longward of 250 nm, and 1050 nm falls between $I$ and $J$ bands. Pertinent details of these grids and their use are now given.

**Castelli and Kurucz [2003] 1221 point ODFNEW Grid**

The complete grid of Castelli and Kurucz [2003] flux distributions (1221 points each, covering 90 nm to 160 $\mu$m) is distributed as a set of files\(^5\), each containing a grid in $T_{\text{eff}}$ (3500–50000 K with 250 K spacing\(^6\)) and log $g$ (0–5 dex with 0.5 dex spacing). Each file is for a different combination of metallicity ($[M/H]$; $-2.5$–$+0.5$ with 0.5 dex spacing, with the addition of $[M/H] = +0.2$), a element enhancement, Helium enhancement ($\Delta Y$), microturbulent velocity ($v_{\text{turb}}$), and mixing length parameter ($\ell/H$). The fluxes have units of ergs/cm$^2$/s/Hz/sr $\equiv 10^3$ W/m$^2$/Hz/sr $\equiv 10^{23}$ Jy/sr.

To reduce the parameter space only models with scaled solar abundances (no $\alpha$-element

\(^5\)http://wwwuser.oat.ts.astro.it/castelli/grids.html
\(^6\)500 K spacing for $T_{\text{eff}} > 10000$ K

108
4.3. PHOTOSPHERE PHOTOMETRY PREDICTION

or Helium enhancement), with \( v_{\text{turb}} = 2 \text{ km/s} \), and with the default mixing length parameter \( l/H = 1.25 \), have been used. Excluding \( \alpha \) enhancement is justified as field A type stars are all young thin disc stars (maximum age \( \sim 1 \text{ Gyr} \)), which generally have little \( \alpha \) enhancement [Pritzl et al., 2005]. Excluding Helium enhancement is justified as He spectral features only start to become apparent for temperatures above about 20000 K. Although Helium abundance affects the physical parameters of stars \( (L,R,T_{\text{eff}}) \) of a given mass and age, it does not require extra photosphere models for \( T_{\text{eff}} \lesssim 20000 \text{ K} \). The fixed \( v_{\text{turb}} \) and \( l/H \) are due to only flux distributions with these values being available. A type stars straddle the \( T_{\text{eff}} \) divide between radiative and convective outer layers, with early A type stars being radiative and late A type stars having significant convection. The typical values of \( v_{\text{turb}} \) and \( l/H \) are seen to vary across the \( T_{\text{eff}} \) range of A type stars (see review on convection in A type stars: Smalley [2004]). \( v_{\text{turb}} = 2 \text{ km/s} \) is, however, reasonably close to the average value. On the other hand, \( l/H = 1.25 \), is significantly high compared to typical values for all but the latest A type stars [Smalley, 2004].

The 1221 point flux distributions are very sparsely sampled longward of 10 \( \mu \text{m} \), so to avoid interpolation artifacts or integration errors, interpolation by a power law was used to add points at 1 \( \mu \text{m} \) when parsing the Castelli and Kurucz [2003] files. The interpolation was linear in log-log space:

\[
F(\lambda) = e^{a \ln \lambda + c}, \quad a = \frac{\ln F(\lambda_{i+1}) - \ln F(\lambda_i)}{\ln \lambda_{i+1} - \ln \lambda_i}, \quad c = \ln F(\lambda_i) - a \ln \lambda_i. \tag{4.3}
\]

The effect of this interpolation is shown in Fig. 4.8. Without this interpolation stage, the results of generic interpolations used in the synthetic photometry routines are highly dependant on the chosen interpolation algorithm, and systematically incorrect photometry is produced.

Munari et al. [2005] Grid

The grid of Munari et al. [2005] available online\(^7\) has a resolution of 1 \( \AA \)/sample. The grid covers the same parameter values as the Castelli and Kurucz [2003] grid, except for the omission of \( [M/H] = +0.2 \) models. Fluxes are provided in wavelength flux density units of ergs/cm\(^2\)/s/\( \AA \)/sr, as opposed to the frequency flux density units used in the Castelli and Kurucz [2003] distributions. The values from the Munari et al. [2005] distributions have been scaled by a factor of \( 10^{10} \lambda [\text{m}^2]/c \) whilst parsing the files. A further factor of 1/4 was needed to bring the fluxes in line with the Castelli and Kurucz [2003] flux distributions.

The Munari et al. [2005] grid has an added dimension of stellar rotational velocity, \( v_{\text{rot}} \). The only difference between the rotating and non-rotating flux distributions, however, is that doppler shifts are applied when the flux is summed over the disc of the star [Jauregi et al., 2005, Kurucz, 2005]. As the shape of the star, nor the atmospheric parameters as a function of

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\(^7\)http://archives.astro.ph/eclipse/
Figure 4.8: Kurucz 1221 point flux distributions (for $T_{\text{eff}} = 9000 \, \text{K}, \log g = 4.0, [M/H] = 0.0, \varepsilon_{\text{turb}} = 2.0 \, \text{km/s}$) showing the effect of poor sampling for $\lambda > 10 \, \mu\text{m}$ with different interpolation algorithms. After power law interpolation to points with 1 $\mu$m spacing the results of linear and natural cubic spline interpolation are identical on this scale.

latitude, are varied, the only effect is of spectral broadening. This was found to have negligible effect on photometric colours, even for the relatively narrow Strömgren bands. Rather than using the non-rotating flux distributions, flux distributions for $\varepsilon_{\text{rot}} = 10 \, \text{km/s}$ were used, as the smoothing helped to minimise differences between results due to different interpolation or integration algorithms.
4.3.2 Synthetic Photometry

Synthetic magnitudes were computed for all combinations of model flux distribution and photometric response function, taking care over whether the response functions were in photon counting or energy measuring units. The magnitudes were zero-pointed by subtracting synthetic magnitudes computed in the same way from a Vega reference flux distribution. The Vega reference used was from the Hubble Space Telescope CALSPEC package⁸ (alpha_lyr_stis_005, Feb 2010). This combines observed spectrophotometry from the International Ultraviolet Explorer (IUE) from 1256-1675 Å, HST’s Space Telescope Imaging Spectrograph (STIS) from 1675-5337 Å, and a specially constructed Kurucz model atmosphere at longer wavelengths. The program which implemented the synthetic photometry took three files as input: one containing the grid of model flux distributions, one containing all the response functions, and one containing the Vega reference flux distribution. The output was a table with a row for each synthetic magnitude, and columns containing the model parameters and response function used.

As the model flux distributions and response functions were all provided with their own wavelength samplings, interpolation of at least one of these was required to be able to perform the synthetic photometry integration. For greatest compatibility, all flux distributions and response functions were treated in the same way and were interpolated to a single wavelength sampling before they were used. Due to the range of over 3 decades in wavelength to be covered, and the fact that the width of spectral features is proportional to wavelength, a logarithmic wavelength sampling was used. The wavelength spacing used was $10^{-5}$ λ, corresponding to 1 Å at 10 μm, 0.1 Å at 1 μm etc., and $\sim 2.3 \times 10^5$ values per decade. The interpolation algorithm used was Akima spline interpolation, which was performed using routines from the GNU Scientific Library⁹. Akima spline interpolation was chosen as it allows smooth interpolation without the excessive overshooting typical of cubic splines for sparsely sampled data.

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⁸http://www.stsci.edu/hst/observatory/cdbs/calspec.html
⁹http://www.gnu.org/software/gsl/
4.3.3 Grid of Photometric Colours

The synthetic magnitudes were used to produce a finer spaced grid of calibrated photometric colours by interpolating in the three model parameters $T_{\text{eff}}$, log $g$, $[M/H]$. Each synthetic magnitude which was used to compute a colour had a zero point offset (ZPO) added to it. Ideally the ZPO for each band would be the magnitude of the photosphere of Vega (i.e. Vega without debris disc) in that band. In reality, the ZPOs deviate slightly (up to $\sim 0.05^m$ in the worst cases) from this to correct to zeroth order systematic errors in the model flux distributions, the Vega reference flux distribution, and the band response functions. Initial values for the ZPOs were taken from literature where available, and they were subsequently iteratively refined during the fitting process. The ZPOs are discussed in the following two subsections.

The colours were interpolated in $T_{\text{eff}}$, log $g$ and $[M/H]$ to produce a grid as described in Table 4.1. The interpolation algorithm used was natural cubic splines. The increase in sampling over the flux distribution grid is typically a factor of 25 in all three parameters. The spacing was chosen to be sufficiently fine that no further interpolation was required for fitting observed colours to this grid. To minimise storage requirements only a minimal set of colours was produced, and other colours were determined from linear combinations of these as required. The grid was stored as a table with a row for each $T_{\text{eff}}$, log $g$ and $[M/H]$ (1.1 x 10$^7$ in total), and a column for each colour.

Table 4.1: Properties of the grid of photometric colours used for fitting observed optical / near-IR photometry and predicting MIPS-24 and MIPS-70 magnitudes. The set of colours is minimal – others were computed from these as necessary e.g. $R_C - I_C = (V_J - I_C) - (V_J - R_C)$.

<table>
<thead>
<tr>
<th>Parameter</th>
<th>Step</th>
<th>Range</th>
</tr>
</thead>
<tbody>
<tr>
<td>$T_{\text{eff}}$</td>
<td>10 K</td>
<td>4000 – 13000 K</td>
</tr>
<tr>
<td>$[M/H]$</td>
<td>0.02 dex</td>
<td>-2.0 – +0.5 [sol]</td>
</tr>
<tr>
<td>log $g$</td>
<td>0.02 dex</td>
<td>3.0 – 5.0 [cm s$^{-2}$]</td>
</tr>
</tbody>
</table>
4.3.4 Photometric Bands

The photometric bands used in this work were chosen to be able to constrain the model photosphere parameters ($T_{\text{eff}}$, $\log g$ and $[M/H]$), and to be able to provide reliable normalisation of the models to produce the predicted MIPS photometry. Fitting the photosphere parameters required a wide range of optical colours, ideally combined with optical-IR colours to tightly constrain $T_{\text{eff}}$. Determining the MIPS photometry from the best fit model was ideally achieved from near-IR photometry, as the IR colours of early type stars are far less sensitive to the choice of photosphere parameters than colours involving optical bands.

To be able to use a given photometric band it was necessary to be able to obtain a reliable response function to compute the synthetic photometry. The bands used had to have measurements of a significant number of the stars of this survey, and it had to be shown that the photometry in each band would be homogeneous. These requirements were due to the need to determine, or at least check, the zero point offsets for each band, as the typical variations of the ZPOs were larger than the measurement uncertainties of the photometry. The photometry used therefore comes from a combination of large homogeneous surveys, and photometric systems which have been very well standardised.

The brightness of the stars of this survey, which includes many of the optically brightest stars in the sky, has been a source of many complications. In modern surveys these stars are either saturated or their photometry suffers from systematic errors. This has resulted in a reliance on legacy photometry, mostly obtained before the advent of CCD detectors, for many of the stars. A brief discussion of the each of the photometric systems which have been used is given below. Further issues are discussed in the following subsections.

**Johnson UBV**

The classic Johnson $UBV$ bands ($\lambda \sim 350, 430, 550\text{ nm}$) were used as photometry in these bands was available for almost all stars considered in this chapter, and most have been measured many times resulting in low uncertainties ($1\sigma \sim 0.01\text{ m}$). Response functions from Maíz Apellániz [2006] have been used for these bands. For the $B$ and $V$ bands the Maíz Apellániz [2006] response functions differ little from previously published versions. The $U$ band, however, has always been problematic as the blue edge is determined by the atmosphere and is thus variable. The $U$ response function from Maíz Apellániz [2006] is a compromise between previously published versions. ZPOs of $0.026 \pm 0.008\text{ m}$ for $V$, $0.036 \pm 0.009\text{ m}$ for $B$ and $0.056 \pm 0.016\text{ m}$ for $U$ from Maíz Apellániz [2006] were used without any modification.

The $UBV$ photometry was retrieved from the General Catalogue of Photometric Data
(GCPD\textsuperscript{10}). The photometry is presented as $V$ magnitudes and $B-V$ and $U-B$ colours. Where there are multiple measurements (usually the case for $UBV$), an uncertainty estimate based on the scatter of the values is given in the GCPD.

**Cousins $RI$**

The Cousins $RI$ bands ($\lambda \sim 600, 800\text{ nm}$) were chosen as these bands are well standardised [Bessell, 1979], and photometry in them was available for a significant number of stars which lacked high quality photometry at longer wavelengths (e.g. saturated in 2MASS). Response functions from the GCPD have been used for these bands\textsuperscript{11}. The ZPOs were not initially known, although they were expected to be similar to the ZPO of $\sim 0.03^m$ for Johnson $V$. Final ZPOs of $0.020^m$ and $0.011^m$ have been used for $R$ and $I$ respectively. The $RI$ photometry was obtained from the GCPD, which generally gave $V-R$, $R-I$ and $V-I$ colours. A conservative $1\sigma$ uncertainty of $0.03^m$ was adopted for these colours. A total of 56 sets of these colours were used.

**Tycho $B_T,V_T$**

The Tycho photometer onboard the *Hipparcos* satellite observed the whole sky in its $B_T$ and $V_T$ photometric bands, with a usable magnitude range of $\sim 2-12^m$. The $B_T$ and $V_T$ bands are similar to the Johnson $BV$ bands, however, the Tycho photometry has the advantage of being completely homogeneous (same instrument, no atmosphere) and well calibrated. A further advantage is that stars with separations as low as $\sim 0.5''$ are resolved in the Tycho-2 and TDSC catalogues. Response functions were taken from Bessell [2000]. These response functions were used by Maiz Apellániz [2006], where no modification was found necessary. ZPOs given by Maiz Apellániz [2006] are $0.034 \pm 0.006^m$ and $0.067 \pm 0.08^m$ for $V_T$ and $B_T$ respectively. These were found to need significant modification. Final values of $0.014^m$ and $0.037^m$ were adopted for $V_T$ and $B_T$ respectively. By examining figures of Maiz Apellániz [2006] it was seen that these values are consistent with the A type stars in that work ($-0.1 < B_T - V_T < 0.35$).

Tycho photometry was taken from the Tycho-2 and TDSC catalogues using the database developed in Chapter 3 of this thesis. The $1\sigma$ uncertainties are typically $0.009^m$ and $0.014^m$ for $V_T$ and $B_T$ respectively (calibration limited). Photometry for saturated stars, which are not included in Tycho-2, were available in the original Tycho catalogue, however, these were not found to be usable due to non-linearity and potentially different ZPOs from the Tycho-2/TDSC photometry.

\textsuperscript{10}http://obswww.unige.ch/gcpd/gcpd.html

\textsuperscript{11}http://obswww.unige.ch/gcpd/filters/fi354.html
**4.3. PHOTOSPHERE PHOTOMETRY PREDICTION**

**Hipparcos $H_p$**

The *Hipparcos* astrometry mission produced photometry for its targets in the $H_p$ band, which is a very wide band covering approximately 400–800 nm. The $H_p$ photometry generally has very high accuracy, with total 1σ uncertainties less than 0.01 m for most of the stars in this work. The actual uncertainties are so low that stellar variability is the limiting factor in many cases, and this is included in the published standard errors. The $H_p$ photometry was taken from F. van Leeuwen [2007].

There are complications for multiple star systems. If multiple stars fall within the detector field of view, of approximately 8'' diameter, and they are all detectable ($H_p \lesssim 9$ m, but contrast and separation dependant), then the published photometry is the sum of the components. Stars which fall outside the field are ignored, although they can cause contamination to the star in the field if especially bright (i.e., problematic for secondaries in binaries of separation $\sim 10''$). Stars which are within the field but are below the detection limit can contribute differing amounts to the published photometry depending on separation and contrast.

The response function was taken from Bessell [2000]. The ZPO was expected to be similar to the Tycho $V_T$ and Johnson $V$ bands, and a final value of 0.034 m was adopted. It was noted that there is nonlinearity for $H_p \lesssim 1$ m. For example, Vega has measured $H_p = 0.0868 \pm 0.014$ m, which is over 3σ fainter than the ZPO.

**Strömgren $uvby$**

The Strömgren system consists of medium ($uvby$) and narrow (Hβ × 2) optical bands. These were designed such that linear combinations of the magnitudes would be sensitive to stellar photosphere parameters. The $b$ and $y$ bands have similar central wavelengths to Johnson $B$ and $V$ bands, and the $b-y$ colour is primarily sensitive to $T_{\text{eff}}$, with less metallicity dependence than $B-V$. The $v$ band is centred on the Hδ line ($\sim 410$ nm), which is in a spectral region with strong line blanketing (metallicity dependence). The $u$ band is located between the Balmer discontinuity at $\sim 360$ nm and the atmospheric transmission cutoff at $\sim 320$ nm. Two standard indices are defined:

\[
m_1 = (v - b) - (b - y) = v - 2b + y \tag{4.4}
\]

\[
c_1 = (u - v) - (v - b) = u - 2v + b. \tag{4.5}
\]

$m_1$ is primarily intended to measure line blanketing (metallicity) although, for stars of mid A type and earlier, $m_1$ also depends strongly on surface gravity due to the deep Hδ absorption line at the centre of the $v$ band for these stars (the line profile varies with gravity). $c_1$ measures
the Balmer discontinuity, which has a maximum around mid A spectral type. $c_1$ is primarily a measure of surface gravity for stars cooler than mid A type, and a measure of temperature for hotter stars. For the stars of mid A type these indices are both complicated functions of $[M/H]$ and log $g$. The narrow H$\beta$ bands of the Strömgren system offer further measures of $T_{\text{eff}}$ and log $g$, which vary the depth and profile of the H$\beta$ line. The H$\beta$ bands have not been used here as standard response functions are not defined, so significant empirical calibration would be required to reliably determine observable values from synthetic photometry.

The relatively narrow bandwidths of the $uvby$ bands, combined with their well chosen edges with respect to atmospheric absorption, mean that photometric uncertainties are generally very low, typically with $1\sigma < 0.01^m$. Response functions and ZPOs from Maíz Apellániz [2006] have been used. Only index ZPOs are given in Maíz Apellániz [2006] $(0.007 \pm 0.003^m, 0.154 \pm 0.003^m$ and $1.092 \pm 0.004^m$ for $b - y$, $m_1$, $c_1$ respectively). Individual band ZPOs determined from these were used: $0.014^m (y)$, $0.021^m (b)$, $0.182^m (v)$, $1.435^m (u)$. These include an arbitrary additive constant. $uvby$ photometry was obtained from the GCPD, and also directly from several catalogues referenced in the GCPD compilation [Crawford and Barnes, 1970, Gronbech and Olsen, 1976, Oblák, 1978, Sowell and Wilson, 1993] where the GCPD values were questionable.

2MASS $JHK_s$

The 2 Micron All Sky Survey Point Source Catalogue [2MASS PSC; Cutri et al., 2003] contains near-IR photometry for $4.7 \times 10^8$ sources over the whole sky, and is generally considered the reference photometric catalogue in the near-IR. The bands are similar to historical Johnson $JHK$ bands (see below), with the exception that the $K_s$ band has a bluer red edge to reduce thermal noise. Response functions were taken from the 2MASS documentation. Final ZPOs of $-0.035^m$, $-0.001^m$ and $-0.070^m$ were adopted for $J$, $H$ and $K_s$ respectively. These are approximately $0.05^m$ more negative than was expected from previous works [Cohen et al., 2003, Rieke et al., 2008], which may be due to the different choices of Vega reference flux distributions or systematic errors in the model flux distributions which have been used here (e.g. a discontinuity between the Castelli and Kurucz [2003] and Munari et al. [2005] distributions).

There are a number of important limitations which apply to the stars used in this work. 2MASS took 2 lengths of exposure: ‘READ1’ mode 51 ms exposures, and ‘READ2’ mode 1.3 s exposures. Typically stars appear in six images of each mode. Almost all the stars in this work are saturated in the READ2 images, so the photometry in the 2MASS PSC comes from the READ1 images (rd_f1g = 1 or 3 in the PSC). Even in these images many of the stars are

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12http://web.ipac.caltech.edu/staff/vsm/2mass/spt_cal/index.html
saturated and no useful photometry is presented in the PSC (rd_f1g = 3). Generally J and H saturate more easily than Ks (approximate saturation magnitudes of 4.5, 4 and 3.5 for J, H and Ks respectively), so often Ks is all that was available.

The READ1 photometry in the PSC was obtained by aperture photometry with a source aperture of 4′′ radius and a sky annulus between 14 and 20′′. This causes problems for certain binary stars and in crowded regions. Binaries with separations of 1–30′′, depending on relative brightness, can be problematic. Even if ∆V is large for e.g. an M type companion to an A type star, ∆K is nowhere near as large, so this problem exists even for high optical contrast binaries.

Other limitations of the READ1 photometry are uncertainty of ~1–2% in the zero-point for each READ1 frame relative to the corresponding READ2 frame, and for sources which are saturated in some but not all of the READ1 frames there can be increased uncertainty due to not being able to average out spatial sampling effects and biases due to only the faintest detections being used. Such sources, which are only non-saturated in less than 3 frames, have ph_qual = 'E' in the PSC. For H band it was found necessary to ignore all magnitudes where ph_qual = 'E', as these appear to be systematically faint, with the worst cases being off by ~0.2 mag. For the J and Ks bands, magnitudes where ph_qual = 'E' were generally found to be usable within the elevated uncertainties for these given in the PSC.

The total numbers of usable J, H and Ks measurements (including secondary stars) for the 130 A type systems were 32, 33, and 83 respectively, with median 1σ uncertainties of 0.024 m, 0.033 m and 0.022 m.

**Johnson J K**

The use of legacy near-IR photometry is generally complicated by a lack of standardisation between observatories, due in part to the dependence of the photometric response functions on the atmosphere above each site. It was found that a large number of stars used here (42 out of 130) were observed in a single system defined in Johnson [1962, 1964] and Johnson et al. [1965]. Although this system includes several bands (J: 1.2 μm, H: 1.6 μm, K: 2.2 μm, L: 3.5 μm, M: 5 μm and N: 9 μm), targets of this survey mostly only have J and K photometry. As a result, only the J and K bands have been used here as there are insufficient stars with photometry in the other bands to determine ZP0s and uncertainties. Instrumental response functions were taken from Johnson [1964]. It was found necessary to multiply these by the atmospheric transmission. The atmospheric transmission function for the 2MASS north observatory was used for this\(^\text{13}\), as this site is at similar altitude (2306 vs. 2070 m) and is located only 500 miles

\(^{13}\text{http://web.ipac.caltech.edu/staff/waw/2mass/opt_cal/index.html}\)
from McDonald Observatory where the majority of the Johnson photometry was taken. The effect on the response functions is shown in Fig. 4.9.

The $JK$ photometry was obtained from Johnson et al. [1966] and Glass [1974] using the GCPD. In total 45 $JK$ pairs were used. Final ZPOs of $0.002^m$ and $-0.013^m$ for $J$ and $K$ respectively, and a typical $1\sigma$ uncertainty of $0.03^m$ for both bands, were determined.

![Combining instrumental $J$ (left) and $K$ (right) response functions from Johnson [1964] with typical atmospheric transmission to obtain total response functions for synthetic photometry. Note that the $J$ instrumental response has significant response in the gap between $J$ and $H$ atmospheric windows ($\sim 1.36\mu m$).](image)

**Figure 4.9:** Combining instrumental $J$ (left) and $K$ (right) response functions from Johnson [1964] with typical atmospheric transmission to obtain total response functions for synthetic photometry. Note that the $J$ instrumental response has significant response in the gap between $J$ and $H$ atmospheric windows ($\sim 1.36\mu m$).

**MIPS-24 and MIPS-70**

The response functions for the MIPS bands were taken from the *Spitzer* Science Centre web site. The ZPOs were set to $ZPO(K_s) + 0.022^m = -0.048^m$, following Rieke et al. [2008]. The zero point fluxes for MIPS-24 and MIPS-70 were assumed to be 7170 mJy and 778 mJy respectively [Rieke et al., 2008, Gordon et al., 2007]. The photometry predicted using these ZPOs and ZP fluxes are compared with the observed photometry in the following section.

**IRAS 12 \mu m**

The majority of the targets of this survey are bright enough to have been detected by *IRAS* at 12 $\mu m$ (see §2.2.1). Response functions for the *IRAS* bands were taken from Beichman et al. [1988, Table II.C.5]. The *IRAS* catalogues give instrumental fluxes rather than magnitudes, so a zero point flux had to be adopted in addition to a zero point offset – although these are

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14http://ssc.spitzer.caltech.edu/files/spitzer/MIPSfiltsumm.txt
equivalent. It was chosen to use the same ZPO for [12] as for the [24] and [70] MIPS bands (−0.048\textsuperscript{m}), so that the mid/far-IR colours of the Vega reference flux distribution would be zero. The zero point flux was adjusted instead of the ZPO, with a final value of 41.5 mJy being adopted. This is 1–3\% higher than published values, which range from 40.1 to 41.0 mJy [Beichman et al., 1988, Cohen et al., 1992, Rieke et al., 2008].

\textit{IRAS} photometry was generally only taken from the FSC (see §2.2.1), as the majority of stars which are in the PSC and not in the FSC are in highly confused regions. The FSC values for the stars were obtained using the database described in chapter 3, which includes cross identifications with the \textit{IRAS} catalogues determined by astrometric matching (including proper motion to the \textit{IRAS} epoch, and using the positional uncertainty ellipses in the \textit{IRAS} catalogues). Care was taken to ignore or down-weight [12] values which were suspected of being confused or in excess due to dust emission (e.g. stars with probable excess at 24 \textmu m). In total 117 [12] measurements have been used, with a median 1σ uncertainty of 0.06\textsuperscript{m}.

\subsection{4.3.5 Zero Point Offsets}

Accurately determining the ZPOs was found to be crucial in exploiting the accuracy of the observed optical and near-IR photometry. The uncertainties in measured photometry in several systems are < 0.01\textsuperscript{m}, which is far smaller than the absolute values of the ZPOs of ∼ 0.03\textsuperscript{m}. Some of the reasons for the range of ZPOs in Vega photometric systems are that Vega itself is often too bright to be reliably observed, and in the near-IR Vega’s debris disc contributes to its flux ∼ 0.014\textsuperscript{m} at K\textsubscript{s}; Rieke et al., 2008. For synthetic photometry, the ZPOs also act as zeroth order corrections for systematics in the flux distributions, response functions and numerical methods (interpolation and integration). As this work has concentrated on A type stars, it is possible that the ZPOs determined here are not generally applicable, as the systematic errors due to the model flux distributions and response functions will vary with the stellar parameters.

The adopted ZPOs are listed in Table 4.2. The Strömgren uvby and Johnson UBV ZPOs were taken from Maiz Apellániz [2006] without modification. As the response functions from the same paper have been used here, and the Vega reference is also similar, it is not surprising that these ZPOs work well here. Altering the Strömgren uvby ZPOs was also undesirable, as these bands are specifically designed to be sensitive to stellar parameters, so adjusting the ZPOs would systematically alter the fitted stellar parameters.

With the exception of the MIPS and \textit{IRAS} ZPOs, which were fixed at ZPO(K\textsubscript{s}) + 0.022\textsuperscript{m}, following Rieke et al. [2008], the ZPOs of the other bands (and the [12] zero point flux – see above) were iteratively adjusted to minimise the systematic offsets between [24] magnitudes determined from the different bands using the best fit models. The process of fitting the models
using a preliminary set of ZPOs and determining the adjustments necessary to the ZPOs was repeated several times until the adjustments were found to be negligible. As the Strömgren uvby and Johnson UBV ZPOs were fixed, the tendency will have been to make the best fit stellar parameters from the other colours consistent on average with those which would be determined from uvby and UBV alone. Whilst this must be borne in mind when considering the stellar parameters determined from the fitting, it does not have a negative impact on the predicted MIPS photometry. In fact, it has a positive effect by making photometry predictions based on different sets of optical/near-IR colours consistent.

Table 4.2: Adopted zero point offsets (ZPOs) for synthetic photometry

<table>
<thead>
<tr>
<th>System</th>
<th>Band</th>
<th>ZPO (mag)</th>
<th>System</th>
<th>Band</th>
<th>ZPO (mag)</th>
</tr>
</thead>
<tbody>
<tr>
<td>Tycho</td>
<td>$B_T$</td>
<td>0.037</td>
<td>Cousins</td>
<td>$R_C$</td>
<td>0.020</td>
</tr>
<tr>
<td>Tycho</td>
<td>$V_T$</td>
<td>0.014</td>
<td>Cousins</td>
<td>$I_C$</td>
<td>0.011</td>
</tr>
<tr>
<td>Hipparcos</td>
<td>$H_p$</td>
<td>0.034</td>
<td>Johnson</td>
<td>$J_3$</td>
<td>0.002</td>
</tr>
<tr>
<td>Strömgren</td>
<td>$b$</td>
<td>0.021</td>
<td>Johnson</td>
<td>$K_3$</td>
<td>−0.013</td>
</tr>
<tr>
<td>Strömgren</td>
<td>$y$</td>
<td>0.014</td>
<td>2MASS</td>
<td>$J$</td>
<td>−0.035</td>
</tr>
<tr>
<td>Strömgren</td>
<td>$v$</td>
<td>0.182</td>
<td>2MASS</td>
<td>$H$</td>
<td>−0.001</td>
</tr>
<tr>
<td>Strömgren</td>
<td>$u$</td>
<td>1.435</td>
<td>2MASS</td>
<td>$K_s$</td>
<td>−0.070</td>
</tr>
<tr>
<td>Johnson</td>
<td>$U_j$</td>
<td>0.056</td>
<td>IRAS</td>
<td>[12]</td>
<td>−0.048</td>
</tr>
<tr>
<td>Johnson</td>
<td>$B_j$</td>
<td>0.036</td>
<td>Spitzer/MIPS</td>
<td>[24]</td>
<td>−0.048</td>
</tr>
<tr>
<td>Johnson</td>
<td>$V_j$</td>
<td>0.026</td>
<td>Spitzer/MIPS</td>
<td>[70]</td>
<td>−0.048</td>
</tr>
</tbody>
</table>
4.3.6 Model Fitting

For each star the observed optical/IR colours were fitted by $\chi^2$ minimisation to the grid of synthetic colours which had been computed from the model photosphere flux distributions. The choice to fit colours rather than absolute photometry was partly practical, as it avoided the need for a normalisation constant to be computed for each considered grid point when fitting. The other justification was that in most photometric systems the colours are more accurately calibrated than the absolute photometry in individual bands.

Where photometry was presented as colours within systems these were used directly (Johnson $U - B, B - V;$ Johnson-Cousins $V - R_C, V - I_C; \text{ Strömgren } b - y, m_1, c_1$). For the Tycho photometry the $B_T - V_T$ colour was used (published values are individual $B_T, V_T$ magnitudes). For other photometry presented as individual magnitudes (Johnson $J, K, 2\text{MASS } JHK_s, \text{ IRAS } [12]$), the mean of $V - [\text{band}], V_T - [\text{band}]$ and $H_p - [\text{band}]$ was used. $V, V_T$ and $H_p$ all have similar effective wavelengths and the uncertainties in each are generally all $\sim 0.01^m$.

The fitting procedure performed a full search for the $\chi^2$ minimum within a specified range of $T_{\text{eff}}, \log g$ and $[M/H]$ in the grid of synthetic colours. A $T_{\text{eff}}$ range constraint was generally used to reduce the computing time (typically $\sim 2000^\circ K$ range). Ranges of $\log g$ and $[M/H]$ were only constrained to investigate multiple $\chi^2$ minima. It was found that most stars required some manual intervention in the fitting due to peculiarities or binarity of the stars, or erroneous (e.g. contaminated/confused) photometric values. As a result, the fitting program was implemented as a web-based application with a form for entering colours, magnitudes, fluxes and uncertainties in their native form, and parameter range constraints. The form was processed by a script which computed colours and uncertainties from magnitudes and fluxes, and then returned the entries from the colour grid with the lowest $\chi^2$ values. For each returned entry, the squared residuals (contribution to the $\chi^2$) for each colour, and the $[24]$ magnitude which would be determined from each photometric band, were given. These values were used to determine which colours or bands were discrepant. The uncertainties of the offending colours or magnitudes could then be tweaked or the values removed completely, and the fit performed again. In many cases constraints on $[M/H]$ were tried to check for the effects of chemical peculiarities, which often resulted in two $\chi^2$ minima with different metallicities (see below).

Care was required for colours with very small measurement uncertainties. When published uncertainties were below $0.01^m$ for Strömgren colours, or $0.02^m$ for Johnson $UBV$ colours, the uncertainties were generally set to these values. For colours computed from magnitudes or fluxes ([12] only) by the fitting script, the uncertainty used was $0.01^m$ added in quadrature with the magnitude uncertainties. Where published uncertainties were not available, typical values for each colour/band were determined from residuals to fits.
CHAPTER 4. PROPERTIES OF NEARBY A TYPE STARS WITH SPITZER

4.3.7 Predicting MIPS Photometry

Once the best fit model had been selected by fitting observed colours, the magnitude in any band could be estimated from the observed magnitude in another band and the appropriate colour from the model. To predict the MIPS photometry, [24] was determined from a weighted average of [24] values obtained using several observed bands. The model [24] − [70] colour was used to then determine [70].

As the IR colours of A type stars are small (\( \lesssim 0.1^m \)), near/mid-IR photometry was used with its natural uncertainties in the weighted averaging. For 2MASS and IRAS photometry, the uncertainties from the relevant catalogues were used. For Johnson JK bands, a typical 1σ uncertainty of 0.03\(^m\) was adopted. Although high accuracy optical photometry was generally available, a conservative 1σ uncertainty of 0.05\(^m\) was adopted for optical photometry used in the weighted average (c.f. typical measurement uncertainties of \( \sim 0.01^m \)). Only \( V_J, V_T \) and \( H_P \) optical bands have been used, as these are the best calibrated optical bands (photometry in all three was available for most stars), the predicted [24] from bluer bands is even more model dependant, and absolute \((RJ)_C\) photometry was generally not available.

The adopted 1σ uncertainty on [24] and [70] was the uncertainty of the weighted average added in quadrature with 0.02\(^m\). This extra 0.02\(^m\) contribution was determined by varying the model parameters within ranges giving acceptable \( \chi^2 \) values for the fit to the observed colours and noting the variation in predicted [24]. In summary, the predicted MIPS-24 magnitude and uncertainty of a star was determined by:

\[
[24] = \sum_i \frac{[i] - ([i] - [24])}{\sigma_i^2} / \sum_i \frac{1}{\sigma_i^2}, \quad \sigma_{[24]}^2 = 0.02^2 + 1/ \sum_i \frac{1}{\sigma_i^2}, \quad (4.6)
\]

\[
\text{where } \sigma_i = \begin{cases} 
\text{published } \sigma & \text{if } i = \{[12], J, H, K_s\}; \\
0.03^m & \text{if } i = \{J_J, K_J\}; \\
0.05^m & \text{if } i = \{V_J, V_T, H_P\}. 
\end{cases} \quad (4.7)
\]

The uncertainty on the predicted MIPS photometry is thus in the range 0.02–0.05\(^m\), or 2–5% in flux. Improving on this would require either more realistic photosphere models (see the following subsections), and sufficient observational data to fit them, or high accuracy near-IR photometry. Although 2MASS provides photometry with 1σ uncertainties down to \( \sim 0.015^m \) for stars which are not saturated in any detections, a large fraction of the stars in this sample are saturated in some or all detections. 2MASS photometry is also not usable for binaries with separations of a few arcseconds as the photometry of bright stars is performed using aperture photometry (4′ source radius and 14–20′′ sky annulus).
4.3.8 Peculiar Stars

"I do not now, nor have I ever, seen a normal A0 star."

– W. P. Bidemann, quoted in Eggen [1984].

Many A type stars have spectra which are peculiar in some way, primarily exhibiting abnormal element abundances and effects due to rapid stellar rotation. As none of these peculiarities are present in the model photosphere flux distributions which have been used in this work (or any grids of model photosphere flux distributions of which this author is aware), there are inherent problems with fitting the photometric colours for these stars. Some A stars also exhibit variability (δ Scuti, α CVn and γ Doradus types), which reduces the accuracy of the photometry as it was not all performed simultaneously.

Chemical Peculiarities

There are several commonly used classifications of chemically peculiar (CP) A stars [Preston, 1974]: Am (CP1, metallic line) stars, which have high metallicity, but are deficient in Calcium and/or Scandium; Ap (CP2) stars, which exhibit abnormally high magnetic fields and enhanced abundances of Silicon, Chromium, Strontium, Europium and other rare earth elements; HgMn (CP3) stars, which have enhanced Mercury and Manganese abundances; and λ Bootis stars [Venn and Lambert, 1990], which are rapid rotators with roughly solar abundances of Carbon, Nitrogen, Oxygen and Sulphur, but are deficient in iron peak elements (Iron, Manganese etc.). In reality there is considerable blurring between these classes, and the majority of A type stars exhibit a mix of these peculiarities at some level.

When fitting the Am stars it was usually found that there were two relatively poor $\chi^2$ minima, corresponding to metallicities of roughly solar and to $+0.5$ (the maximum in the grids). The low metallicity solution would generally be a better match to the $B - V$ and $B_T - V_T$ colours, and the high metallicity solution would be a better fit to $U - B$ and $m_1$. This makes sense, as the Calcium H and K lines lie in the overlap between the $U$ and $B$ bands and don’t strongly affect any Strömgren bands. The $U - B$ colour, like the $m_1$ colour, is thus relatively unaffected by the calcium abundance and these colours reflect the high metallicity. The $B - V$ colours are bluer than would be expected for the metallicity due to the decreased $B$ band absorption. Generally the high metallicity solutions have been adopted. Where the metallicity would need to be $>+0.5$, the $m_1$ and $U - B$ uncertainties have been increased to avoid them dominating the $\chi^2$ and skewing log $g$ and $T_{\text{eff}}$.

The other CP types were less common in the sample. The symptoms of the Ap stars were a general inability to produce a simultaneous good fit to $c_1$, $m_1$ and the broadband optical colours. The strong Silicon, Chromium and Strontium absorption features of these stars primarily occur
CHAPTER 4. PROPERTIES OF NEARBY A TYPE STARS WITH SPITZER

in the $v$ and $B/B_T$ bands. Unlike the Am stars where the $U-B$ colour is conveniently unaffected by the peculiarity, all of $U-B$, $B-V$, $B_T-V_T$, $m_1$ and $c_1$ are affected to differing degrees. In general for these stars the bluest colours have had their uncertainties increased for the fitting to ensure that the model prediction has meaningful colours for $V$ and longer bands for the MIPS photometry prediction, although the fitted parameters are likely to be incorrect (mostly $[M/H]$ is incorrectly high).

To the author’s knowledge there are no HgMn stars in this sample. The λ Bootis stars in the sample have generally ended up with metallicities which are likely in between the light metal and iron peak abundances. Generally there was no sign in the fitting that there was an abundance peculiarity, although the fitted metallicities are all lower than solar.

**Rotation effects**

A type stars are seen with rotational velocities from a few 10s of km/s up to very near their breakup velocity (typically ~ 200-300 km/s). Rapid rotation causes stars to become oblate, with the radius at their equator larger than at the poles. The centrifugal force and the difference in radius causes the atmospheric parameters (mostly log $g$ and $T_{\text{eff}}$) to vary from pole to equator. This effect has been observed directly by near-IR interferometry observations of nearby A stars such as Vega and Altair [Peterson et al., 2006a,b]. Variations of $T_{\text{eff}}$ and log $g$ from equator to pole can be 2000 K and 0.5 dex respectively [Peterson et al., 2006a]. This inevitably causes problems for fitting models with a single $T_{\text{eff}}$ and log $g$.

The primary effect of oblateness on the fitting was that the best fit $T_{\text{eff}}$ and log $g$ would vary systematically by weighting down either blue optical colours or red/near-IR colours. The blue optical colours ($c_1$, $m_1$, $B-V$, $U-B$, $b-g$, $B_T-V_T$) would favour a hotter temperature and higher log $g$, while the redder colours would favour lower temperatures and lower log $g$.

The fitted parameters for Altair from this work ($T_{\text{eff}} = 7650$ K, log $g = 3.88$ [cm s$^{-2}$]) are closer to the equator parameters ($T_{\text{eff}} = 6890$ K, log $g = 3.85$ [cm s$^{-2}$]) than the non-rotating model parameters ($T_{\text{eff}} = 8200$ K, log $g = 4.26$ [cm s$^{-2}$]) from Peterson et al. [2006a]. This is due to the use of several red/IR colours in the fit here (using $RI$, $JK$ and $[12]$ bands), which are less biased towards emission from the hot poles than the optical colours. The fitted parameters from this work better reflect the red/IR colours, which were used to predict the MIPS photometry, than parameters fitted from optical colours alone.
4.3. PHOTOSPHERE PHOTOMETRY PREDICTION

4.3.9 Binaries

Multiple star systems with separations up to several tens of arcseconds were problematic for the fitting and flux prediction due to the methods used to produce the observed photometry at all wavelengths. This is unfortunate, as the MIPS-24 PSF fit photometry reliably resolves stars with separations of less than 5″.

The legacy photometry \((UBV, uvby, RI, JK)\) was generally performed with photomultiplier tubes or similar, which measured flux within a circular aperture of typical radius 10–20″. The legacy photometry thus provides reasonable combined photometry for components with separations \(\lesssim 10″\). For separations of \(\sim 10–30″\) the photometry generally contains a less than equal contribution from the secondary star. The situation is similar for the \(IRAS\) photometry, although due to the limited spatial resolution rather than measurement aperture. The 2MASS PSC photometry for bright stars is problematic for separations between \(\sim 2–20″\) (see above), and the \(H_p\) photometry for binaries is generally complicated (see above).

The Tycho photometry was generally the only photometry for components with all separations \(\gtrsim 1″\). The Tycho photometry is resolved for binaries of low contrast down to separation of less than 0.5″, however, it was found that the magnitude differences between components with separations of \(\lesssim 1″\) suffered from significant systematic errors. The Tycho photometry of multiple components could be reliably summed to produce combined photometry for all separations where both stars were detected.

As it was not possible to constrain the stellar parameters with the \(B_T - V_T\) colour alone, it was generally necessary to fit unresolved photometry for multiple star systems with separations up to \(\sim 10–20″\). It was considered out of the scope of this work to attempt to fit multiple star model flux distributions. This would have significantly increased the complexity of the fitting process as the number of parameters would be increased from three to six or seven depending on whether \([M/H]\) would be set to the same for both components. The resulting fits of unresolved photometry to the single stars models generally have \(T_{\text{eff}}\) somewhere between those of the individual components. The situation is similar to the case of rapidly rotating stars, where it is incorrect to characterise the flux distribution by a single set of atmosphere parameters. Generally for these unresolved binaries the uncertainties of the blue optical colours (primarily \(U - B, c_1\)) were increased during the fitting so that the resulting best fit was consistent with the red and near-IR colours.

In general, the fitted colours for the unresolved binaries were not that dissimilar from the colours of the primary stars. For low contrast binaries the stars have similar colours anyway, and for high contrast binaries the secondary contributes relatively little to the photometry (especially in the optical). Where some resolved photometry in the bands used for predicting
the MIPS photometry was available for the primary stars, predicted MIPS photometry was produced from this as well as the unresolved photometry for the sum of the stars. The unresolved predicted magnitudes were expected to be systematically too faint, as the relative brightness of the secondary to the primary is higher in the mid/far-IR than in the optical and near-IR. The resolved predicted magnitudes were expected to be too bright as the model is cooler (redder) than the primary star.

It was possible to quantify the average systematics for such binary systems by comparing the average differences between measured and predicted MIPS-24 photometry for single stars, combined multiple stars and resolved primary stars with models fitted to combined optical/near-IR photometry. Weighted mean differences between measured and predicted magnitudes of $-0.0135 \pm 0.0040^m$, $-0.0163 \pm 0.0067^m$ and $+0.0093 \pm 0.0095^m$ were found for single stars, combined multiples and primaries respectively. Corrections of $-0.0028^m$ and $+0.0228^m$ have been applied to the predicted photometry for the combined multiples and primaries respectively, to bring them in line with the calibration of single stars on average.
4.4 Calibration

Zero point offsets to be subtracted from the magnitudes measured by DAOPHOT, to produce magnitudes in a standard Vega system, were determined using the zero point fluxes of 7170 and 778 mJy for MIPS-24 and MIPS-70 respectively from Rieke et al. [2008] and Gordon et al. [2007]. The post-BCD images have units of MJy/sr, with pixel scales of 2.45′′/pix and 4.0′′/pix for MIPS-24 and MIPS-70 respectively. The units are converted from instrumental units in the SSC pipeline using the flux calibrations of Engelbracht et al. [2007] and Gordon et al. [2007]. The DAOPHOT magnitudes are zero pointed such that an integrated flux of 1 ADU ≡ 1 MJy/sr has a magnitude of 25. Aperture corrections of 1.082 and 1.22 have been used [Engelbracht et al., 2007, Gordon et al., 2007]. The zero point offsets are then,

\[ ZP_{24} = 25^m - 2.5 \log \left( 1.082 \times \frac{2.45''^2}{1 \text{sr}} \times \frac{1 \text{ MJy}}{7170 \text{ mJy}} \right) = 13.320475^m \]  
\[ ZP_{70} = 25^m - 2.5 \log \left( 1.22 \times \frac{4.00''^2}{1 \text{sr}} \times \frac{1 \text{ MJy}}{778 \text{ mJy}} \right) = 16.926625^m. \]

With these zero point offsets subtracted, the DAOPHOT photometry should be in the same system as the predicted MIPS photometry for the stellar photospheres (see above). This assumption was tested by comparing measured and predicted photometry.

4.4.1 MIPS-24

It was expected that the measurement uncertainties, both due to noise and repeatability, for the measured MIPS-24 photometry would be negligible compared to the uncertainties in the predicted photosphere photometry of typically 2–5%. This was indeed found to be the case. The low rate of excess detection at 24 μm, combined with the low rate of contamination of the photometry, meant that it was straightforward to test the calibration by comparing measured and predicted photometry. Initial comparisons showed any offset to be no more than ~0.01 m.

Fig. 4.10 shows the difference between measured and predicted magnitudes as a function of measured magnitude for all the stars in this work with both measurements and predictions. The weighted mean of differences within ±0.09 m of zero (102 values) is \(-0.0080 \pm 0.0032^m\), with a reduced \(\chi^2\) value of 1.1. As it is conceivable that this offset is skewed negative by net excess emission from dust, the offset was not considered significant and no correction for it has been applied. There is no significant sign of nonlinearity over the nine magnitude range covered (factor of 4000 in flux). The reduced \(\chi^2\) value of 1.1 indicates that the adopted uncertainties are reasonable. The uncertainties used were the DAOPHOT measurement uncertainty added in quadrature with the adopted predicted magnitude uncertainty, where the latter generally
dominates.

![Graph](image)

**Figure 4.10:** Comparison of observed PSF fit photometry and predicted photometry for MIPS-24. Predicted photometry has had a zero point offset of 13.3205\textsuperscript{m} added. The weighted average remaining offset has been determined from points with $|([24]\text{PSF} - [24]\text{predicted}| < 0.09\text{m}$.

### 4.4.2 MIPS-70

For MIPS-70 the measurement uncertainties generally dominate over the photosphere prediction uncertainties. The total uncertainty on the measured photometry is the measurement error determined from the PSF fitting added in quadrature with a random calibration uncertainty, which accounts for both repeatability on particular stars and variation of the calibration between stars. Gordon et al. [2007] show the repeatability uncertainty for individual stars to be $\sim 0.04-0.05\text{m}$. The observations of the repeatability standard stars (HD 180711 and HD 163588) have also been reduced here, yielding standard deviations of 0.0236\textsuperscript{m} and 0.0305\textsuperscript{m}. The measurement uncertainties produced by DAOPHOT have been tested using Al tinha, which has eight observations performed consecutively (not significantly affected by long term calibration variation). The standard deviation of the eight measurements is 0.0080\textsuperscript{m} and the average uncertainty reported by DAOPHOT was 0.0087\textsuperscript{m}. The DAOPHOT uncertainties are thus considered to be realistic, at least for bright stars.

The measured and predicted MIPS-70 photometry are compared in Fig. 4.11. The comparison is complicated by the high incidence rate and typical sizes of excess, in addition to a large scatter. Based on various attempts at this comparison, using different selections of stars, an uncertainty on each measurement of the DAOPHOT measurement uncertainty added in quadrature with 0.10\textsuperscript{m} has been adopted. Determining the offset between measured and
predicted magnitudes has been problematic. Various sensible selections of stars yielded offsets from approximately $-0.08^m$ to $+0.01^m$. It has not been possible to robustly show that there is a significant offset from zero. In Fig. 4.11 a somewhat arbitrary selection, of all stars with measured magnitude no more than 1.25$\sigma$ brighter than the predicted magnitude, gives an uncertainty weighted average of almost exactly zero ($0.004\pm0.018^m$). There appears to be a slight non-linearity, however, it has not been possible to show that this is statistically significant. Fits of a straight line have typically yielded a slope of $\sim0.02 \pm 0.01$. After much investigation, it was decided not to apply any offset or linearity correction. With the adopted uncertainties the minimum 3$\sigma$ excess detection threshold is 0.3$^m$, which is far larger than the size of any offset or the effect of the fitted non-linearity over the relevant range of magnitudes.

### 4.4.3 PSF fitting vs. Aperture Photometry

The PSF fitting photometry was compared with aperture photometry for single, uncontaminated stars, as shown in Fig. 4.12. This confirmed that DAOPHOT’s calibration of the PSF fitting photometry to the aperture photometry performed with calibration aperture configurations was correct, and showed that there is no significant non-linearity between the PSF and aperture photometry over the relevant magnitude ranges in each band.
Figure 4.12: Comparison of daophot PSF fit and aperture photometry measurements for the stars used to construct the PSF models (single, uncontaminated). All values are in instrumental magnitudes, with no aperture correction applied. For MIPS-24 (left) the aperture photometry shown was produced with the calibration aperture configuration from Engelbracht et al. [2007, 35" source, 40-50'' sky], and the PSF magnitudes are implicitly calibrated against these by DAOPHT. For MIPS-70 (right), aperture photometry produced using two sets of aperture configurations from Gordon et al. [2007] is shown. The PSF fit magnitudes are calibrated against the larger aperture configuration (35'' source, 39-65'' sky), as this is used for the official MIPS-70 calibration, and has lower systematic uncertainty due to the aperture correction being much smaller than for the more compact configuration (16'' source, 18-39'' sky). The smaller aperture configuration is shown here as the S/N is higher than for the calibration aperture configuration. There is no significant sign of non-linearity in the PSF fit magnitudes when compared to the aperture magnitudes, and the mean offset between PSF fit photometry and the corresponding calibration aperture photometry is essentially zero as it should be.
4.5 Results

4.5.1 Observations Summary

The targets of this survey comprise a volume limited sample of 130 stellar systems with A type
main sequence primary stars (the A type sample in chapter 3 of this thesis and Phillips et al.
[2016]). The observations performed specifically for this survey (Spitzer programme \#50771, PI: Phillips) targeted all systems which did not already have MIPS-24 or MIPS-70 observations
scheduled in November 2007 (49 systems). By the cessation of Spitzer cryogenic operations at
the end of March 2009, a total of 116 systems had been observed by MIPS (114 with both bands
and 2 with only MIPS-24). With the exception of fine-scale MIPS-70F mode observations for
five systems, and MIPS-24 observations of Vega, all observations have been reduced and had
photometry performed as described above. There are seven systems which were not detected
in MIPS-70 observations due to either galactic cirrus emission or insufficient integration time.

It must be noted that many of the observations not obtained specifically for this survey (i.e.
in programmes other than \#50771) have been used in previously published works including
Rieke et al. [2005], Su et al. [2006] and Trilling et al. [2007]. The photometry presented here,
however, has the advantage that it has, as far as possible, all been produced and calibrated in
a homogeneous manner.

4.5.2 Stellar Parameters, Photometry and Excesses

Table 4.3 presents the raw results of this work: the stellar parameters \( (T_{\text{eff}}, \log g \) and \( |M/H| \))
determined from the photosphere flux distribution fitting, the predicted and observed MIPS
photometry in Vega magnitudes, and notes for the MIPS observations and the binarity and
chemical peculiarities of the stars. The photometric uncertainties presented are as described in
§4.3.7, §4.4.1 and §4.4.2. The significance of excess in each band, defined by,

\[
\chi^2_{24} = \frac{[24]_{\text{pred}} - [24]_{\text{obs}}}{\sigma^2_{[24]_{\text{pred}}} + \sigma^2_{[24]_{\text{obs}}}} \quad \text{and} \quad \chi^70 = \frac{[70]_{\text{pred}} - [70]_{\text{obs}}}{\sigma^2_{[70]_{\text{pred}}} + \sigma^2_{[70]_{\text{obs}}}},
\]

(4.10)

is also given. For multiple star systems all combinations of components for which photometry
was measured or predictions could be made are included.

There are many cases for which stars in MIPS-24 images could be reliably resolved by PSF
fitting, although only unresolved predicted and MIPS-70 photometry could be obtained. The
smallest separations resolved in the MIPS-24 photometry are approximately 2\" (e.g. UNS A107
with \( \rho = 1.7\" \) and \( \Delta[24] \approx 0.4\" \), and UNS A112 with \( \rho = 2.8\" \) and \( \Delta[24] \approx 1.4\" \)). The smallest
separation resolved in the MIPS-70 photometry is 14.4\" for UNS A028, although the S/N was

131
CHAPTER 4. PROPERTIES OF NEARBY A TYPE STARS WITH SPITZER

rarely sufficient to detect secondary stars (the primary stars were typically only \( \sim 1'' \) above the detection limit). For separations below \( \sim 10'' \), the measured MIPS-70 single star photometry has been assigned to the sum of the components (c.f. the MIPS-70 PSF FWHM of \( \sim 20'' \)).

4.5.3 Stellar Parameter Comparison

The purpose of the photosphere model spectrum fitting in this work was to predict MIPS photometry by extrapolating from near-IR and red optical photometry. For this, the quality of the fit between the model and observed colours was the primary concern, rather than the physical meaning of the best-fit model parameters. Indeed, given the complications due to chemical peculiarities, rapid rotation and unresolved binaries (§4.3.8, 4.3.9), the physical interpretation of the fitted model parameters is dubious at best.

For comparison, the photosphere parameters determined here for the single stars in the sample were compared with those compiled in chapter 3 from Gray et al. [2003, 2006], and with other literature values retrieved using SIMBAD\(^{15}\) and from Saffe et al. [2008]. These comparisons are shown in Fig. 4.13. For the general literature comparison, all literature values found for a star are shown, creating rows of points in Fig. 4.13. The spread in the literature values of all three parameters is quite considerable. In general there is good agreement on \( T_{\text{eff}} \) in all sources. Values of \([M/H]\) determined here are systematically higher than values from both Gray et al. [2003, 2006] and other literature, although this is largely due to the fitting of Am peculiar stars which have often been fitted with a very high metallicity here (\( \sim +0.5 \)). \( \log(g) \) values determined here are systematically higher than those from Gray et al. [2003, 2006], but are in good agreement with the other literature values, although with a very large scatter.

4.5.4 Dust Properties

For stars with 3\( \sigma \) or greater excess in both bands (\( \chi_{24} \geq 3 \) and \( \chi_{70} \geq 3 \)), a dust temperature and mass were determined from the star-subtracted photometry. The dust temperature combined with the stellar luminosity was used to determine a characteristic orbital radius. Where only one band showed 3\( \sigma \) or greater excess, limits on these parameters were determined.

As both MIPS bands are shortward of the typical wavelengths at which the dust opacity starts to decrease for debris discs (\( \lambda_0 \) in §1.5.2), a simple black body was sufficient to determine the dust temperature. Converting the temperature to a radius, however, requires more specific assumptions about the dust properties (e.g. Equns. (1.13), (1.14) and (1.15)). For simplicity, and to allow comparison with other works, the radii have been computed assuming perfect black body grains i.e. using Eqn. (1.13).

\(^{15}\)http://simbad.u-strasbg.fr/

132
4.5. RESULTS

**Figure 4.13:** Comparison of fitted photosphere parameters with literature values for the single stars in the sample. Top: comparison with values from Gray et al. [2003, 2006]. Bottom: comparison with values from SIMBAD and Saffe et al. [2008].

The determined dust temperatures, characteristic orbital radii and masses are given in Tables 4.4 and 4.5, for stars with significant excess in both bands and only one band respectively. The methods for determining these parameters are described below. Stars for which there is a MIPS-24 excess, but no MIPS-70 photometry could be obtained are listed in Table 4.6.

**Dust Temperatures**

A relationship between [24] – [70] colour and temperature was obtained by performing synthetic photometry (see §4.3.2) on a set of black body flux distributions (Eqn. (1.1)) logarithmically spaced in $T$. An analytical function was fitted to allow $T$ to be determined from [24] – [70] without needing a look-up table and interpolation. The fitted function was,

$$
\log T = (2.4467 \pm 0.0004) - (0.8616 \pm 0.0009) \log ([24] - [70]) \\
+ (0.05052 \pm 0.0014) \log ([24] - [70])^2 - (0.0537 \pm 0.0011) \log ([24] - [70])^3 \\
- (0.00449 \pm 0.0010) \log ([24] - [70])^4.
$$

(4.11)
The $T$ vs. $[24] - [70]$ relationship and residuals of the fit are shown in Fig. 4.14. The residuals in $T$ are less than 1% over the relevant temperature range ($\sim$30–1000 K).

![Graph](image)

**Figure 4.14:** Temperature vs. $[24] - [70]$ relationship. Left: $T$ vs. $[24] - [70]$ colour computed from synthetic photometry. Right: fractional residuals of fitted function. Points with $T < 30$ K were excluded from the fit.

For stars with significant excess in both MIPS bands, the $[24] - [70]$ colour of the dust emission was determined from the flux excesses (measured minus predicted fluxes):

$$([24] - [70])_{\text{dust}} = 2.5 \log (\Delta F_{70}) - 2.5 \log (\Delta F_{24}) .$$

Temperature uncertainties were determined by varying the measured $[70]$ and predicted $[24]$ by $\pm 1 \sigma$ before computing $\Delta F_{70}$ and $\Delta F_{24}$. For stars with significant excess in only one band, a $3 \sigma$ upper limit on the excess in the other band was used to compute a temperature limit.

**Characteristic Orbital Radii**

Stellar luminosities were computed using absolute $V$ band magnitudes and bolometric corrections as in §3.5.3. The stellar effective temperatures in Table 4.3 were used to determine the bolometric correction. The stellar luminosities are given alongside the dust properties in Tables 4.4 and 4.5.

The characteristic orbital radii were computed assuming perfect black body grains in thermal equilibrium using Eqn. (1.13). Radius uncertainties were determined by computing $r$ for the $1\sigma$ upper and lower temperature values. These radii are generally lower limits, as realistic grains are hotter at any given distance. The discrepancy between these black body radii and radii in steady-state collisional debris disc models of Wyatt et al. [2007b], as a function of radius.
and stellar spectral type (proxy for luminosity), is shown in Fig. 4.15. The quoted radius uncertainties in Table 4.4 do not include this model dependant uncertainty.

![Graph showing orbital radius ratios](image)

**Figure 4.15:** Ratio of real orbital radius in steady-state collision models of Wyatt et al. [2007b] to orbital radius computed from [24] – [70] colour assuming black body grains. Figure from Bonsor and Wyatt [2010].

**Dust Masses**

Dust masses were computed from Eqn. (1.17) using the determined dust temperatures. An opacity of $\kappa_{70 \mu m} = 1.0 \text{ m}^2/\text{kg}$ was used, which was determined by assuming typical debris disc values of $\kappa_{850 \mu m} = 0.17 \text{ m}^2/\text{kg}$, $\beta = 1$ and $\lambda_0 = 150 \mu m$ [e.g. Wyatt, 2008]. It must be noted that these dust masses are highly uncertain, especially for discs with significant masses of dust with temperatures below $\sim 70$ K. Determining the masses of discs is one of the primary goals of longer wavelength surveys with *Herschel* and ground based sub-mm instruments.

**Orbit Configurations in Multiple Star Systems**

Where dust was detected in multiple star systems, the orbital configuration of the dust – circumbinary, circum-primary, circum-secondary etc. – was determined. The projected separations of visual binaries and the orbital periods of spectroscopic binaries were compared with the dust radii. The semimajor axis of an elliptic orbit of period $P$ is,

$$\frac{a}{\text{AU}} = \left( \frac{P}{\text{yr}} \right)^{2/3} \left( \frac{M_A + M_B}{M_\odot} \right)^{1/3} \approx 1.5 \left( \frac{P}{\text{yr}} \right)^{2/3},$$

(4.13)
CHAPTER 4. PROPERTIES OF NEARBY A TYPE STARS WITH SPITZER

for an A type primary and a lower mass secondary. The orbital configurations are presented in Table 4.7. The statistical properties of dust around binary stars is considered in the following section.
Table 4.3: Fitted stellar parameters and observed and predicted MIPS photometry. For binary systems, various combinations of components for which photometry was available are given. The ‘Stat’ column indicates which entry for each system was used in the statistical analyses. The ‘Bin.’ and ‘Pec.’ columns indicate binarity (see end of table for values) and chemical peculiarity (Am, Ap or λ Boo). The $T_{\text{eff}}$, $\log g$ and $[M/H]$ values are from the photosphere flux distribution fits to optical/near-IR photometry. These were used to produce the predicted MIPS photometry. The observed and predicted MIPS photometry are in Vega magnitudes, and are given with 1σ uncertainties as discussed in the text. The $\chi^2_{24}$ and $\chi^2_{70}$ values are the significance of excess, defined as e.g. $\chi^2_{24} = \frac{(24)_{\text{obs}} - (24)_{\text{pred}}}{\sqrt{\sigma_{(24)_{\text{obs}}}^2 + \sigma_{(24)_{\text{pred}}}^2}}$. The F24 and F70 columns contain flags for the MIPS observations and photometry (see end of table for values).

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<th>Pec.</th>
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<th>$\log g$</th>
<th>$[M/H]$</th>
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<th>$(24)_{\text{pred}}$ (mag)</th>
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<td>2.269 ± 0.005</td>
<td>3.261 ± 0.032</td>
<td>29.1</td>
<td>1.247 ± 0.141</td>
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<td>8380</td>
<td>4.06</td>
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<td>3.083 ± 0.005</td>
<td>3.072 ± 0.026</td>
<td>-0.4</td>
<td>3.020 ± 0.104</td>
<td>3.075 ± 0.026</td>
<td>0.5</td>
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log g

4.06

5080

4.50

4.12

8330
10130

3.70

3.82

3.86

7100

9280

7380

3.98

3.98

9310

3.92

9310

3.92

7750

7750

3.88

4.02

9210

4.02

3.90

4.00

3.98

3.84

4.20

4.30

4.10

4.06

3.90

4.02

3.94

3.94

3.94

4.30

7955

[M/H]

0.14

0.50

0.48

−0.04

0.42

0.50

6.221 ± 0.006

3.067 ± 0.006

3.429 ± 0.005

2.367 ± 0.002

3.520 ± 0.005

2.893 ± 0.007

7.005 ± 0.085

2.918 ± 0.007

−0.50

3.502 ± 0.006

−0.50

0.0

1.2

0.4

1.0

0.6

0.6

6.211 ± 0.023 −0.4

3.078 ± 0.026

3.476 ± 0.032

3.461 ± 0.032

2.382 ± 0.026

3.542 ± 0.035

2.887 ± 0.026 −0.2

2.918 ± 0.035

3.554 ± 0.041

3.578 ± 0.041 −0.1

3.577 ± 0.035 −0.2

0.2

0.1

2.249 ± 0.041 −0.7
2.018 ± 0.041

3.581 ± 0.006

0.5

0.1

9.8

0.0

1.837 ± 0.027 −1.4

3.066 ± 0.032

1.757 ± 0.031

2.345 ± 0.026

3.195 ± 0.026

2.010 ± 0.010

6.380 ± 0.043

1.8

3.9

3.1

4.915 ± 0.054 −2.9

3.929 ± 0.033

3.535 ± 0.033

4.061 ± 0.026

1.4

2.792 ± 0.027 −1.7
2.442 ± 0.026

3.663 ± 0.041

3.585 ± 0.006

12.0

D

χ24 :

2.824 ± 0.035 −1.0

2.193 ± 0.026

0

[24]pred

3.661 ± 0.007

2.278 ± 0.004

1.876 ± 0.001

3.049 ± 0.005

1.753 ± 0.011

2.068 ± 0.005

3.194 ± 0.025

5.463 ± 0.178

3.337 ± 0.025

3.867 ± 0.004

3.401 ± 0.006

3.966 ± 0.013

2.404 ± 0.003

2.841 ± 0.005

2.860 ± 0.005

1.852 ± 0.004

0

[24]obs

−0.10

−0.10

0.00

0.50

0.50

0.50

0.46

−0.38

−2.00

0.50

−0.70

0.00

0.46

−0.06

−0.10

0.36

−1.74

−1.74

0.00

[cm/s2 ] [sol]

8400

9440

9150

8570

9000

9440

8600

6355

7490

8030

6590

8190

8490

8490

9560

N

Teff

2.912 ± 0.103

2.671 ± 0.141

2.194 ± 0.101

2.597 ± 0.112

3.216 ± 0.106

3.434 ± 0.107

1.806 ± 0.101

3.496 ± 0.119

2.063 ± 0.101

1.960 ± 0.101

2.830 ± 0.102

1.663 ± 0.100

0.666 ± 0.224

3.377 ± 0.107

3.637 ± 0.112

2.579 ± 0.101

2.360 ± 0.103

2.796 ± 0.104

0.569 ± 0.102

0

[70]obs

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15.5

2.6

9.0

0.9

7.5

2.2

0.8

1.7

5.5

1.9

2.6

3.0

1.3

2.1

1.4

1.8

 )  E 1

6.207 ± 0.023

3.095 ± 0.026

3.488 ± 0.032

3.462 ± 0.032

2.396 ± 0.026

3.551 ± 0.035

2.892 ± 0.026

2.923 ± 0.035

3.557 ± 0.041

3.581 ± 0.041

3.583 ± 0.035

2.032 ± 0.041

3.676 ± 0.041

2.264 ± 0.041

1.850 ± 0.027 −1.1

3.070 ± 0.032

1.748 ± 0.031

2.360 ± 0.026

3.196 ± 0.026 −1.6

4.911 ± 0.054

3.938 ± 0.033

3.540 ± 0.033

4.058 ± 0.026

2.452 ± 0.026

D

χ70 :/#

2.783 ± 0.027 −0.1

2.815 ± 0.035

2.203 ± 0.026

0

[70]pred

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<th>Bin., Pec.</th>
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<th>log $g$</th>
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<th>$[24]_{\text{pred}}$</th>
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<td>O</td>
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Table 4.3: Fitted stellar parameters and observed and predicted MIPS photometry (continued)

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Table 4.3: Fitted stellar parameters and observed and predicted MIPS photometry (continued)

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Table 4.3: Fitted stellar parameters and observed and predicted MIPS photometry (continued)

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### Table 4.3: Fitted stellar parameters and observed and predicted MIPS photometry (continued)

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**Binarity flag values (Bin. column):**

- **B** – entry includes multiple stars (both combined entries for visual binaries, and spectroscopic/astrometric/eclipsing/occultation binaries etc.).
- **C** – entry is for a single component star in a multiple star system.

**MIPS-24 and MIPS-70 observation flag values (F24 and F70 columns):**

- **O** – not observed – there are 14 systems not observed with either band, and a further two (UNS A010 and A081) observed with MIPS-24 only.
- **C** – strong galactic cirrus emission voided MIPS-70 observation. Cirrus generally also visible with MIPS-24 but PSF fit photometry unaffected.
- **S** – photometry from Su et al. [2008] used here. This is the case for five systems with only fine-scale MIPS-70 observations (UNS A063, A012, A014, A016, A024).

**Notes for individual stars and systems:**

- UNS A003 (Vega, HD 172167) is saturated in the MIPS-24 images, and as the disc is highly resolved it is not possible to fit the wings of the PSF.
- UNS A066 (λ Gem, CCDM 07 181+832) appears to have a third star (X’ here) located ~1.4′′ from the A component in the MIPS-24 image.
- HD 163588 and HD 180711 are the MIPS-70 repeatability calibrators. These were used to construct the MIPS-70 PSF model.
Table 4.4: Physical disc parameters for stars with significant excess at both MIPS-24 and MIPS-70. The luminosities were computed from the fitted \( T_{\text{eff}} \) values, absolute \( V \) magnitudes and bolometric corrections as in §3.5.3. Dust temperatures were computed from the \([24] – [70]\) colour of the excess (see text). The disc orbital radii were computed from the temperatures and stellar luminosities assuming black body grains. The disc masses were computed from the excess MIPS-70 flux and the temperature, assuming an opacity of \( 1 \text{ m}^2/\text{kg} \). Note that some photometry for UNS A003 (Vega), A014 (\( \beta \) Pic), A016 (\( \zeta \) Lep) and A024 (\( \beta \) UMa) come from Su et al. [2006] rather than this work.

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<th>( d ) (pc)</th>
<th>( L ) ((L_\odot))</th>
<th>( \Delta[24] ) (mag)</th>
<th>( \Delta[70] ) (mag)</th>
<th>( T ) (K)</th>
<th>( r ) (AU)</th>
<th>( M ) (( M_\odot ))</th>
</tr>
</thead>
<tbody>
<tr>
<td>A003 HD 172167</td>
<td>7.7</td>
<td>55.1</td>
<td>0.299 ± 0.031</td>
<td>2.921 ± 0.225</td>
<td>80 ± 6</td>
<td>87 ± 13</td>
<td>1.1 × 10^{-2}</td>
</tr>
<tr>
<td>A004 HD 219956</td>
<td>7.7</td>
<td>17.8</td>
<td>0.288 ± 0.030</td>
<td>3.760 ± 0.105</td>
<td>71 ± 3</td>
<td>66 ± 5</td>
<td>1.4 × 10^{-2}</td>
</tr>
<tr>
<td>A005 HD 109647</td>
<td>11.0</td>
<td>14.8</td>
<td>0.365 ± 0.030</td>
<td>2.052 ± 0.104</td>
<td>115 ± 7</td>
<td>23 ± 3</td>
<td>7.0 × 10^{-4}</td>
</tr>
<tr>
<td>A013 HD 115892</td>
<td>18.0</td>
<td>23.8</td>
<td>0.191 ± 0.028</td>
<td>0.461 ± 0.105</td>
<td>260 ± 167</td>
<td>6 ± 4</td>
<td>5.3 × 10^{-5}</td>
</tr>
<tr>
<td>A014 HD 39000</td>
<td>19.4</td>
<td>9.2</td>
<td>3.532 ± 0.030</td>
<td>6.499 ± 0.174</td>
<td>110 ± 6</td>
<td>19 ± 2</td>
<td>3.7 × 10^{-2}</td>
</tr>
<tr>
<td>A016 HD 38678</td>
<td>21.6</td>
<td>15.3</td>
<td>0.993 ± 0.034</td>
<td>2.018 ± 0.145</td>
<td>211 ± 29</td>
<td>7 ± 2</td>
<td>2.7 × 10^{-4}</td>
</tr>
<tr>
<td>A018 HD 139006</td>
<td>23.0</td>
<td>66.5</td>
<td>0.341 ± 0.028</td>
<td>1.634 ± 0.105</td>
<td>131 ± 30</td>
<td>37 ± 5</td>
<td>1.3 × 10^{-3}</td>
</tr>
<tr>
<td>A021 HD 2262</td>
<td>23.8</td>
<td>12.6</td>
<td>0.135 ± 0.035</td>
<td>0.961 ± 0.106</td>
<td>125 ± 20</td>
<td>18 ± 5</td>
<td>2.4 × 10^{-4}</td>
</tr>
<tr>
<td>A024 HD 95418</td>
<td>24.5</td>
<td>64.9</td>
<td>0.277 ± 0.028</td>
<td>1.690 ± 0.225</td>
<td>118 ± 16</td>
<td>45 ± 11</td>
<td>1.6 × 10^{-3}</td>
</tr>
<tr>
<td>A042 HD 87606</td>
<td>28.2</td>
<td>10.7</td>
<td>0.249 ± 0.031</td>
<td>0.792 ± 0.112</td>
<td>193 ± 30</td>
<td>7 ± 2</td>
<td>8.4 × 10^{-4}</td>
</tr>
<tr>
<td>A033 HD 172555</td>
<td>28.5</td>
<td>8.3</td>
<td>2.051 ± 0.033</td>
<td>3.011 ± 0.105</td>
<td>265 ± 28</td>
<td>3 ± 1</td>
<td>3.1 × 10^{-4}</td>
</tr>
<tr>
<td>A047 HD 27045</td>
<td>28.9</td>
<td>7.3</td>
<td>0.101 ± 0.029</td>
<td>0.787 ± 0.120</td>
<td>125 ± 25</td>
<td>13 ± 5</td>
<td>1.3 × 10^{-4}</td>
</tr>
<tr>
<td>A053 HD 125162</td>
<td>30.4</td>
<td>17.1</td>
<td>0.438 ± 0.031</td>
<td>3.166 ± 0.105</td>
<td>89 ± 3</td>
<td>41 ± 3</td>
<td>4.4 × 10^{-3}</td>
</tr>
<tr>
<td>A057 HD 88955</td>
<td>31.1</td>
<td>25.2</td>
<td>0.121 ± 0.032</td>
<td>0.752 ± 0.107</td>
<td>137 ± 29</td>
<td>21 ± 8</td>
<td>2.3 × 10^{-4}</td>
</tr>
<tr>
<td>A059 HD 161868</td>
<td>31.5</td>
<td>29.3</td>
<td>0.644 ± 0.036</td>
<td>3.905 ± 0.107</td>
<td>84 ± 3</td>
<td>59 ± 4</td>
<td>1.4 × 10^{-2}</td>
</tr>
<tr>
<td>A064 HD 20320</td>
<td>33.6</td>
<td>11.2</td>
<td>0.119 ± 0.028</td>
<td>1.972 ± 0.104</td>
<td>84 ± 6</td>
<td>37 ± 5</td>
<td>1.6 × 10^{-3}</td>
</tr>
<tr>
<td>A073 HD 14005</td>
<td>31.4</td>
<td>28.0</td>
<td>0.555 ± 0.030</td>
<td>3.874 ± 0.105</td>
<td>84 ± 2</td>
<td>62 ± 4</td>
<td>1.3 × 10^{-2}</td>
</tr>
<tr>
<td>A072 HD 31295</td>
<td>35.7</td>
<td>15.5</td>
<td>0.433 ± 0.034</td>
<td>3.894 ± 0.105</td>
<td>76 ± 3</td>
<td>53 ± 2</td>
<td>1.1 × 10^{-2}</td>
</tr>
<tr>
<td>A076 HD 110411</td>
<td>36.3</td>
<td>13.1</td>
<td>0.463 ± 0.030</td>
<td>3.461 ± 0.104</td>
<td>84 ± 3</td>
<td>40 ± 3</td>
<td>4.7 × 10^{-3}</td>
</tr>
<tr>
<td>A082 HD 71155</td>
<td>37.5</td>
<td>30.0</td>
<td>0.578 ± 0.028</td>
<td>2.545 ± 0.104</td>
<td>116 ± 6</td>
<td>36 ± 3</td>
<td>2.0 × 10^{-3}</td>
</tr>
<tr>
<td>A085 HD 172523</td>
<td>38.4</td>
<td>30.5</td>
<td>0.341 ± 0.035</td>
<td>1.413 ± 0.109</td>
<td>147 ± 16</td>
<td>20 ± 4</td>
<td>4.4 × 10^{-4}</td>
</tr>
<tr>
<td>A086 HD 13161</td>
<td>38.9</td>
<td>79.9</td>
<td>0.222 ± 0.028</td>
<td>2.486 ± 0.103</td>
<td>86 ± 4</td>
<td>93 ± 9</td>
<td>1.4 × 10^{-2}</td>
</tr>
<tr>
<td>A087 HD 95008</td>
<td>39.0</td>
<td>23.6</td>
<td>0.569 ± 0.032</td>
<td>0.718 ± 0.140</td>
<td>778 ± 366</td>
<td>1 ± 2</td>
<td>1.8 × 10^{-5}</td>
</tr>
<tr>
<td>A102 HD 146624</td>
<td>41.3</td>
<td>20.4</td>
<td>0.215 ± 0.027</td>
<td>1.277 ± 0.181</td>
<td>128 ± 19</td>
<td>21 ± 6</td>
<td>2.9 × 10^{-4}</td>
</tr>
<tr>
<td>A103 HD 1404</td>
<td>41.3</td>
<td>23.6</td>
<td>0.184 ± 0.035</td>
<td>1.252 ± 0.111</td>
<td>122 ± 15</td>
<td>25 ± 5</td>
<td>4.6 × 10^{-4}</td>
</tr>
<tr>
<td>A109 HIP 117452 AB</td>
<td>42.1</td>
<td>26.3</td>
<td>0.486 ± 0.030</td>
<td>1.774 ± 0.111</td>
<td>146 ± 12</td>
<td>19 ± 3</td>
<td>4.5 × 10^{-4}</td>
</tr>
<tr>
<td>A109 HD 223340 C</td>
<td>42.1</td>
<td>0.4</td>
<td>0.262 ± 0.028</td>
<td>3.732 ± 0.116</td>
<td>69 ± 3</td>
<td>10 ± 1</td>
<td>1.5 × 10^{-3}</td>
</tr>
<tr>
<td>A113 HD 23281</td>
<td>42.4</td>
<td>8.6</td>
<td>0.451 ± 0.026</td>
<td>1.710 ± 0.105</td>
<td>144 ± 11</td>
<td>11 ± 2</td>
<td>2.7 × 10^{-4}</td>
</tr>
<tr>
<td>A115 HD 37594</td>
<td>42.6</td>
<td>5.9</td>
<td>0.123 ± 0.026</td>
<td>4.471 ± 0.105</td>
<td>54 ± 2</td>
<td>64 ± 5</td>
<td>4.0 × 10^{-2}</td>
</tr>
<tr>
<td>A117 HD 197950</td>
<td>42.7</td>
<td>8.8</td>
<td>0.082 ± 0.026</td>
<td>1.800 ± 0.104</td>
<td>80 ± 3</td>
<td>36 ± 8</td>
<td>1.1 × 10^{-3}</td>
</tr>
<tr>
<td>A123 HD 213308 A</td>
<td>43.8</td>
<td>35.2</td>
<td>0.295 ± 0.032</td>
<td>1.625 ± 0.106</td>
<td>124 ± 10</td>
<td>30 ± 4</td>
<td>7.3 × 10^{-4}</td>
</tr>
<tr>
<td>A125 HD 150492</td>
<td>44.5</td>
<td>13.2</td>
<td>0.939 ± 0.025</td>
<td>3.180 ± 0.104</td>
<td>118 ± 5</td>
<td>20 ± 2</td>
<td>2.2 × 10^{-3}</td>
</tr>
</tbody>
</table>
Table 4.5: Physical disc parameter limits for stars with significant excess in only one band. The limits computed for stars with only MIPS-70 excess use a [24] magnitude of \([24]_{\text{predicted}} - 3\sigma_{\Delta[24]}\). Similarly, for stars with only MIPS-24 excess, a \([70]_{\text{predicted}} - 3\sigma_{\Delta[70]}\) was used.

<table>
<thead>
<tr>
<th>UNS</th>
<th>name</th>
<th>(d) (pc)</th>
<th>(L) ((L_\odot))</th>
<th>(\Delta[24]) ((\text{mag}))</th>
<th>(\Delta[70]) ((\text{mag}))</th>
<th>(T) (K)</th>
<th>(r) (AU)</th>
<th>(M) ((M_\odot))</th>
</tr>
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<tbody>
<tr>
<td></td>
<td>MIPS-70 excess only</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>A003</td>
<td>HIP 12828 AB</td>
<td>25.8</td>
<td>10.8</td>
<td>0.033 ± 0.034</td>
<td>0.791 ± 0.145</td>
<td>&lt; 134</td>
<td>&gt; 14</td>
<td>&gt; 2.2 \times 10^{-4}</td>
</tr>
<tr>
<td>A051</td>
<td>HIP 105319 AB</td>
<td>30.3</td>
<td>13.3</td>
<td>0.014 ± 0.030</td>
<td>0.540 ± 0.109</td>
<td>&lt; 167</td>
<td>&gt; 10</td>
<td>&gt; 1.1 \times 10^{-4}</td>
</tr>
<tr>
<td>A082</td>
<td>HD 159560 A</td>
<td>30.4</td>
<td>8.6</td>
<td>-0.017 ± 0.026</td>
<td>0.550 ± 0.121</td>
<td>&lt; 119</td>
<td>&gt; 16</td>
<td>&gt; 1.6 \times 10^{-4}</td>
</tr>
<tr>
<td>A061</td>
<td>HD 186228</td>
<td>32.2</td>
<td>29.2</td>
<td>0.015 ± 0.033</td>
<td>1.380 ± 0.107</td>
<td>&lt; 96</td>
<td>&gt; 45</td>
<td>&gt; 6.9 \times 10^{-4}</td>
</tr>
<tr>
<td>A073</td>
<td>HD 16754 A</td>
<td>35.7</td>
<td>14.0</td>
<td>0.026 ± 0.029</td>
<td>0.534 ± 0.127</td>
<td>&lt; 156</td>
<td>&gt; 12</td>
<td>&gt; 1.0 \times 10^{-4}</td>
</tr>
<tr>
<td>A074</td>
<td>HD 79439</td>
<td>35.8</td>
<td>12.4</td>
<td>0.004 ± 0.028</td>
<td>0.367 ± 0.111</td>
<td>&lt; 173</td>
<td>&gt; 9</td>
<td>&gt; 8.3 \times 10^{-5}</td>
</tr>
<tr>
<td>A106</td>
<td>HD 210049</td>
<td>41.6</td>
<td>24.7</td>
<td>0.015 ± 0.031</td>
<td>0.639 ± 0.133</td>
<td>&lt; 155</td>
<td>&gt; 16</td>
<td>&gt; 1.6 \times 10^{-4}</td>
</tr>
<tr>
<td>A108</td>
<td>HIP 31167 AB</td>
<td>41.8</td>
<td>8.3</td>
<td>0.027 ± 0.025</td>
<td>0.670 ± 0.119</td>
<td>&lt; 132</td>
<td>&gt; 13</td>
<td>&gt; 1.4 \times 10^{-4}</td>
</tr>
<tr>
<td>A111</td>
<td>HD 28527 A</td>
<td>42.5</td>
<td>18.4</td>
<td>0.025 ± 0.032</td>
<td>0.596 ± 0.134</td>
<td>&lt; 150</td>
<td>&gt; 15</td>
<td>&gt; 1.6 \times 10^{-4}</td>
</tr>
<tr>
<td>A120</td>
<td>HD 186219</td>
<td>43.6</td>
<td>10.8</td>
<td>0.064 ± 0.025</td>
<td>0.625 ± 0.119</td>
<td>&lt; 158</td>
<td>&gt; 10</td>
<td>&gt; 1.2 \times 10^{-4}</td>
</tr>
<tr>
<td>A122</td>
<td>HD 48807</td>
<td>43.6</td>
<td>13.8</td>
<td>-0.007 ± 0.026</td>
<td>1.496 ± 0.124</td>
<td>&lt; 83</td>
<td>&gt; 42</td>
<td>&gt; 8.2 \times 10^{-4}</td>
</tr>
<tr>
<td></td>
<td>MIPS-24 excess only</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
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</tr>
<tr>
<td>A009</td>
<td>HD 203280</td>
<td>15.0</td>
<td>19.4</td>
<td>0.100 ± 0.028</td>
<td>0.060 ± 0.104</td>
<td>&gt; 218</td>
<td>&lt; 7</td>
<td>&lt; 8.5 \times 10^{-5}</td>
</tr>
<tr>
<td>A048</td>
<td>HD 125161 A</td>
<td>29.1</td>
<td>8.8</td>
<td>0.007 ± 0.031</td>
<td>0.130 ± 0.122</td>
<td>&gt; 191</td>
<td>&lt; 6</td>
<td>&lt; 4.5 \times 10^{-5}</td>
</tr>
<tr>
<td>A099</td>
<td>HD 6961</td>
<td>41.0</td>
<td>25.9</td>
<td>0.129 ± 0.034</td>
<td>0.147 ± 0.129</td>
<td>&gt; 222</td>
<td>&lt; 8</td>
<td>&lt; 1.0 \times 10^{-4}</td>
</tr>
<tr>
<td>A111</td>
<td>HD 123998</td>
<td>42.3</td>
<td>16.0</td>
<td>0.150 ± 0.029</td>
<td>0.059 ± 0.150</td>
<td>&gt; 218</td>
<td>&lt; 7</td>
<td>&lt; 7.7 \times 10^{-5}</td>
</tr>
</tbody>
</table>

Table 4.6: Stars with MIPS-24 excess but no MIPS-70 photometry.

<table>
<thead>
<tr>
<th>UNS</th>
<th>name</th>
<th>(d) (pc)</th>
<th>(L) ((L_\odot))</th>
<th>(\Delta[24]) ((\text{mag}))</th>
<th>Notes</th>
</tr>
</thead>
<tbody>
<tr>
<td>A020</td>
<td>HD 130819 B</td>
<td>23.2</td>
<td>3.9</td>
<td>0.096 ± 0.031</td>
<td>F4 V, (p = 271''), excess probably not real</td>
</tr>
<tr>
<td>A055</td>
<td>HD 135379</td>
<td>30.5</td>
<td>19.2</td>
<td>0.565 ± 0.037</td>
<td>MIPS-70 non-detection due to shallow observation</td>
</tr>
<tr>
<td>A100</td>
<td>HD 198639</td>
<td>41.0</td>
<td>13.2</td>
<td>0.235 ± 0.031</td>
<td>very strong cirrus emission</td>
</tr>
<tr>
<td>A126</td>
<td>HD 37507</td>
<td>44.5</td>
<td>20.1</td>
<td>0.203 ± 0.027</td>
<td>strong cirrus, unable to fit star</td>
</tr>
</tbody>
</table>
Table 4.7: Orbit configurations for discs detected in multiple star systems. Configuration abbreviations: B: circum-binary, P: circum-primary, S: circum-secondary, T: circum-tertiary.

<table>
<thead>
<tr>
<th>UNS</th>
<th>name</th>
<th>P (days)</th>
<th>pd (AU)</th>
<th>rdust (AU)</th>
<th>Conf.</th>
<th>Note</th>
</tr>
</thead>
<tbody>
<tr>
<td>A018</td>
<td>HD 130006</td>
<td>17</td>
<td>(0.2)</td>
<td>37^{+6}_{-5}</td>
<td>B</td>
<td></td>
</tr>
<tr>
<td>A033</td>
<td>HIP 12820 AB</td>
<td>3</td>
<td>&gt; 14</td>
<td>B</td>
<td></td>
<td></td>
</tr>
<tr>
<td>A043</td>
<td>HD 172555</td>
<td>2026</td>
<td>3^{+1}_{-1}</td>
<td>P</td>
<td></td>
<td></td>
</tr>
<tr>
<td>A047</td>
<td>HD 27045</td>
<td>5206</td>
<td>13^{+5}_{-4}</td>
<td>P</td>
<td></td>
<td></td>
</tr>
<tr>
<td>A048</td>
<td>HD 125161 A</td>
<td>1125</td>
<td>&lt; 6</td>
<td>P</td>
<td></td>
<td></td>
</tr>
<tr>
<td>A051</td>
<td>HIP 105319 AB</td>
<td>181</td>
<td>&gt; 10</td>
<td>?</td>
<td></td>
<td></td>
</tr>
<tr>
<td>A052</td>
<td>HD 159560 A</td>
<td>38</td>
<td>(0.3)</td>
<td>16</td>
<td>B</td>
<td>Aa–Ab, ( \rho d(B) = 1881 \text{ AU} )</td>
</tr>
<tr>
<td>A064</td>
<td>HD 20320</td>
<td>18</td>
<td>(0.2)</td>
<td>37^{+6}_{-5}</td>
<td>B</td>
<td></td>
</tr>
<tr>
<td>A073</td>
<td>HD 16754 A</td>
<td>850</td>
<td>&gt; 12</td>
<td>P</td>
<td></td>
<td>circum-binary not ruled out</td>
</tr>
<tr>
<td>A086</td>
<td>HD 13161</td>
<td>31</td>
<td>(0.3)</td>
<td>93^{+10}_{-9}</td>
<td>B</td>
<td></td>
</tr>
<tr>
<td>A108</td>
<td>HIP 31167 AB</td>
<td>176</td>
<td>&gt; 13</td>
<td>?</td>
<td></td>
<td></td>
</tr>
<tr>
<td>A109</td>
<td>HIP 117432 AB</td>
<td>164</td>
<td>19^{+3}_{-3}</td>
<td>P</td>
<td></td>
<td>MIPS-24 photometry resolved</td>
</tr>
<tr>
<td>A109</td>
<td>HD 223340 C</td>
<td>3143</td>
<td>10^{+1}_{-1}</td>
<td>T</td>
<td></td>
<td>second disc in system</td>
</tr>
<tr>
<td>A114</td>
<td>HD 28827 A</td>
<td>10623</td>
<td>&gt; 15</td>
<td>P</td>
<td></td>
<td></td>
</tr>
<tr>
<td>A122</td>
<td>HD 48097</td>
<td>483</td>
<td>(1.8)</td>
<td>&gt; 42</td>
<td>B</td>
<td></td>
</tr>
<tr>
<td>A123</td>
<td>HD 21388 A</td>
<td>1327</td>
<td>30^{+5}_{-4}</td>
<td>P</td>
<td></td>
<td></td>
</tr>
<tr>
<td>A126</td>
<td>HD 37507</td>
<td>446</td>
<td>(1.7)</td>
<td>?</td>
<td>B?</td>
<td>no MIPS-70 due to cirrus</td>
</tr>
</tbody>
</table>
4.5.5 Statistics

The primary motivation for the volume limited samples presented in Chapter 3 was to allow unbiased statistics for debris disc incidence rates and properties to be determined. Within the A star sample studied here it has been possible to explore debris properties as a function of binarity, metallicity, effective temperature, chemical peculiarities and separation in binary systems. There are very few known planets around A type stars, so correlating the presence of debris with the presence of planets was not possible\textsuperscript{16}.

A significant factor which is not investigated here is system age. Measuring the ages of individual stars is notoriously difficult, and the peculiarities and rotation of A type stars especially complicate this. The evolution of debris around A stars has previously been investigated in Su et al. [2006], which primarily targeted members of stellar kinematic groups and clusters with known ages. The direct investigation of evolution within the sample of field stars studied here is considered outside the scope of this work.

In the following statistical analyses a single entry for each system is used, as indicated by the 'Stat.' column in Table 4.3. Generally the entry used is for the sum of all components within approximately 10\arcsec of the primary star. This corresponds to the resolution of the MIPS-70 photometry and most of the legacy optical/near-IR and IRAS 12\,$\mu$m photometry used in the photosphere photometry prediction.

It must be noted that the robustness of the statistical conclusions which can be drawn from this work, whilst greater than previous works with more biased sample selections, is still limited by various systematic factors. Firstly, as described in Chapter 2, the dust mass sensitivity of MIPS observations is a strong function of dust temperature, and hence orbital radius and stellar luminosity ($M \propto 1/B_\nu(T_g), T_g \propto L^{1/4}r^{-1/2}$ – Eqs. (1.13) and (1.17)). The luminosity range in this sample is relatively small (approximately 5-100$L_\odot$, a factor of two in $L^{1/4}$), however, the rapid fall-off in dust sensitivity for dust temperatures below $\sim$50 K means that there may be significant masses of dust at $\gtrsim 100$ AU which have escaped detection here. A second factor is that not all stars in the volume limited sample were observed before the end of the Spitzer cryogenic mission, and for MIPS-70 several stars are not detected. These missing data are likely to increase the excess detection rates, as many of the programmes which have observed stars in this sample have selected stars with previously known excess, and systems with excess are more likely to be detected in the MIPS-70 observations.

The basic statistics of various combinations of subsamples of systems, which are discussed below, are summarised in Table 4.8. For each subsample the total number of systems ($N_{\text{tot}}$) and the numbers observed ($N_{\text{obs}}$) and detected ($N_{\text{det}}$) in each band are given, to allow the effects of

\textsuperscript{16}The only planets detected around A type stars of which this author is aware were detected in imaging of scattered starlight from debris discs.
missing data on the excess fractions and average excess magnitudes to be judged. Two excess fractions are given for each band, one computed ignoring all missing data \((f = N_{\text{ex}}/N_{\text{det}})\), where \(N_{\text{ex}}\) is the number of systems with significant excess), and the other computed assuming all missing systems do not have excess \((f_{\text{min}} = N_{\text{ex}}/N_{\text{tot}})\). In all cases, significant excess in each band is defined as \(\chi \geq 3.0\), where \(\chi\) is the excess significance as in Eqn. (4.10). The uncertainties on the excess fractions obey Poisson statistics, and in the following sections uncertainties on the excess fractions of \(\sqrt{N_{\text{ex}}/N}\) are assumed (this is a good approximation where \(N_{\text{ex}} \geq 5\).

**The Whole Sample**

Histograms of the number of stars as a function of excess significance in the two bands are shown in Fig. 4.16. Stars with \(\chi \geq 3.0\) are highlighted. The typical uncertainties on the excesses are approximately 0.03\(m\) and 0.11\(m\) for MIPS-24 and MIPS-70 respectively, with minimum values of 0.02\(m\) and 0.10\(m\) (see above). Note that it is useful to define the excess in terms of magnitudes (log excess ratio) as opposed to the excess ratio – as typically used in other works [e.g. Su et al., 2006] – as the uncertainty distribution is expected to be normal in magnitudes as opposed to log-normal in excess flux ratio.

The excess rates for the detected stars in the sample as a whole are 33 ± 5\% and 41 ± 6\% for MIPS-24 and MIPS-70 respectively. The MIPS-70 figure is especially likely to be an overestimate due to bias of the detected systems toward systems with excess. The lower limit on MIPS-70 excess rate, assuming all the missing systems have no excess, is 32 ± 5\%. These results are very similar to the excess rates found by Su et al. [2006], of 32±5\% and ≥33±5\% for MIPS-24 and MIPS-70, for their sample of ~160 A type stars in clusters and stellar kinematic groups. The criteria for determining excess used here and in Su et al. [2006] are essentially the same for MIPS-24, however, the MIPS-70 excess criterion is slightly more relaxed here, with a typical threshold of ~0.33\(m\) compared with the 0.45\(m\) threshold in Su et al. [2006]. Setting a MIPS-70 excess detection threshold of \(\Delta[70] \geq 0.45m\) here instead of the \(\chi_{70} \geq 3.0\) threshold yields 41 systems with excess, which is only one less than previously. As the excess detection rate is effectively unchanged, the overall detection rates found here are considered to be the same as in Su et al. [2006], despite the different sample selections.

**Binarity**

The entries for the systems used for the statistics were split into single stars, entries representing multiple stars (unresolved or the sum of photometry for individual stars), and entries for individual primary stars in wide multiple systems. The separation threshold between the combined and wide binaries was approximately 10\(\prime\prime\) (~100-500 AU for all relevant cases) as
Table 4.8: Subsample excess properties. \( N_{\text{tot}} \) is the total number of systems in each subsample. For each band: \( N_{\text{obs}} \) is the number of systems observed; \( N_{\text{ex}} \) is the number of systems with \( \geq 3\sigma \) excesses; \( f \) is the excess fraction, \( N_{\text{ex}}/N_{\text{det}} \), as a percentage; \( f_{\text{min}} \) is the excess fraction, \( N_{\text{ex}}/N_{\text{tot}} \), assuming all non-observed and non-detected systems do not have significant excess; \( \langle \Delta_m \rangle \) and \( \langle \Delta_r \rangle \) are the mean excess magnitudes for detected systems.

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<th>Subsample</th>
<th>( N_{\text{tot}} )</th>
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<th>( N_{\text{ex}} )</th>
<th>( f )</th>
<th>( f_{\text{min}} )</th>
<th>( \langle \Delta_m \rangle )</th>
<th>( \langle \Delta_r \rangle )</th>
<th>( N_{\text{obs}} )</th>
<th>( N_{\text{ex}} )</th>
<th>( f )</th>
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4.5. RESULTS
described above. In addition to the already compiled information from chapter 3, a literature search was performed for each system to identify further multiplicity. The work of chapter 3 had extensively searched for visual and common proper motion (CPM) binaries, however, identification of unresolved multiple stars was less extensive. The catalogue of Eggleton and Tokovinin [2008], which compiles multiplicity information for bright stars (\(H_p < 6.0\)), has been especially useful for finding extra multiple systems and additional stars in systems.

The strongest factor found here to affect the dustiness of systems in this sample is binarity. As shown in Table 4.8, the incidence rates for the combined binary stars (separations less than a few hundred AU) are a factor of 2–5 lower than for the single stars and the individual primaries in wider binaries. Histograms of the excess significance in the two bands for the combined binaries, and single stars and primaries of wide binaries, are shown in Fig. 4.17. The excess incidence rates for MIPS-24 and MIPS-70 (for detected stars) are 11 ± 5% and 21 ± 7% for the combined binaries, and 48 ± 8% and 57 ± 10% for single stars and primaries of wide binaries. The rates for these samples thus differ by \(\sim 3-4\sigma\). If the primaries of wide binaries are added to the binary sample rather than the single star sample, then the rates are 15 ± 5% and 26 ± 7% for all binaries, and 50 ± 9% and 57 ± 11% for the single stars. The result still holds at the 3\(\sigma\) level.

This significant result differs to that from a previous work which surveyed 69 F-type and late A-type binary stars with MIPS [Trilling et al., 2007]. Trilling et al. [2007] stated that the excess rate for binary stars was the same as, or even higher, than for single stars of similar spectral types. The MIPS-24 and MIPS-70 excess detection rates they determined for the binary stars are 9 ± 4% and 40 ± 7% respectively. Comparison of these rates with the ones determined in
4.5. RESULTS

![Histograms](image)

**Figure 4.17:** Excess significance histograms for single and binary stars. The primaries of wide binaries are included with the single stars here, as their excess statistics are essentially the same. Note that β Pic (UNS A014, single star) has a MIPS-24 excess of 118σ (not shown).

this work is not straightforward due to the excesses thresholds being defined in different ways. Trilling et al. [2007] define excess in a band as $\chi \geq 2.0$ and $\Delta[24] \geq 0.15^m$ and $\Delta[70] \geq 0.28^m$. In this work simply $\chi \geq 3.0$ is used to define excess, resulting in typical thresholds of approximately $\Delta[24] \geq 0.09^m$ and $\Delta[70] \geq 0.33^m$ (minimum thresholds of 0.07$m$ and 0.30$m$). Allowing for this difference, the MIPS-24 detection rate in Trilling et al. [2007] and the rate here are essentially the same (there is only one binary system here with $0.07^m < \Delta[24] < 0.15^m$). For MIPS-70 the comparison is more dubious, as there are many systems near the excess thresholds, and the aperture photometry used in Trilling et al. [2007] is more prone to contamination than the PSF fit photometry used here. Applying the MIPS-70 excess criteria used here ($\chi \geq 3.0$, with minimum 1σ photometric uncertainty of 0.10$m$) to the photometry and predictions of Trilling et al. [2007] yields 11 binaries with MIPS-70 excess out of 50 detections, i.e. 22 ± 7%, which is essentially the same as found here. The reason Trilling et al. [2007] found a higher rate for binary stars than other authors found for single stars is simply because Trilling et al. [2007] use
more lenient MIPS-70 excess criteria (e.g. Su et al. [2006] use $\chi \geq 3.0$ or $\Delta[70] \geq 0.45^m$).

**Binary Separation**

To investigate the effects in binary systems further, estimates of orbital radii were obtained. For small separations ($\rho \ll 1''$) these were generally determined from measured orbital periods, converted to semi-major sizes using Eqn. (4.13). For larger separation binaries, where orbital periods were not available, the projected linear separation, $a[\text{AU}] \approx \rho[''] \times d[\text{pc}]$, was used as the separation. The excess MIPS-24 and MIPS-70 magnitudes for the systems are plotted against binary separation in Fig. 4.18, with systems possessing an excess in either band highlighted. There is a noticeable gap between $\sim 3\text{AU}$ and $\sim 150\text{AU}$, in which there are no systems with detected excess.

To quantify the significance of the difference between the binary separation distributions of systems with and without dust, statistical tests were performed on the cumulative distributions shown in the bottom plot of Fig. 4.18. First the Kolmogorov-Smirnov (K–S) test was applied. This quantifies the probability of observing the maximum vertical separation between sample cumulative distributions if they are drawn from the same underlying distribution. Here the probability of the measured maximum separation between the cumulative distributions for binary systems with and without excess, if they have the same underlying distribution, was 14%.

The K–S probability showed that a real difference in the distributions was likely, however, it was not a robust indication. The K–S test is primarily sensitive to differences in the cumulative distributions near the median, where the difference is small here. A related test known as Kuiper's test was also applied. This is essentially the same as the K–S test, except that the sum of the maximum positive and negative distances between the cumulative distributions is used instead of just the larger of these. The Kuiper test gave a probability of 0.3% for the systems with and without excess to yield the measured sum of absolute maximum positive and negative differences, if they are drawn from the same distribution. This shows that this result, of excess fraction depending on binary separation, is significant.

It is quite possible that the systems with separations of order 10 AU possess circum-binary dust at distances of a few hundred AU, and simply this is not detected here as the temperature is too low. The question of whether these systems are completely lacking massive debris discs or they possess large circum-binary discs will be answered by longer wavelength surveys with *Herschel* and sub-millimetre instruments.
Figure 4.18: Magnitude excesses as a function of physical binary separation. Systems with dust detected in either band are shown in dark red, and systems with no significant excess detection are shown in grey. The pale green band highlights separations of 3-100 AU, in which there are no systems with significant excess in either band. The bottom plot shows the cumulative distributions used to perform the K-S and Kuiper tests.
Metallicity

Trends with metallicity were investigated using the \([M/H]\) values from the photosphere flux distribution fits (Table 4.3). The distribution of metallicities for the 130 systems are shown in Fig. 4.19. The large bin sizes of 1 dex were chosen to produce large samples, between which excess detection rates could meaningfully be compared. Of the systems in the lowest metallicity bin \((-2 \leq [M/H] < -1)\), all but one are binaries where the fitted metallicity is an artifact due to the composite spectrum (the exception is the \(\lambda\) Boo star HD 19260). Due to the dubious nature of the metallicities, and the small number of systems, in the lowest metallicity bin, only the two higher metallicity bins are considered here. The upper bound of +1 on the highest metallicity bin is beyond the maximum of +0.5 which could be fitted, however, there are a considerable number of stars for which +0.5 appeared to be too low to correctly fit the observed optical/near-IR colours (primarily Am peculiar stars; in total there are 18 systems with fitted \([M/H] = +0.5\)).

![Figure 4.19: Distribution of metallicities ([M/H]) from fitting photosphere flux distributions.](image)

With reference to the values in Table 4.8, it is seen that the excess rates and mean excess magnitudes are greater for the lower metallicity systems. For the sample as a whole there is a \(~1\sigma\) distinction between the excess rates, with MIPS-24 and MIPS-70 excess rates of 38 \pm 8\% and 47 \pm 10\% for \(-1 \leq [M/H] < 0\), and 31 \pm 7\% and 36 \pm 9\% for \(0 \leq [M/H] < +1\). This does not appear to be strongly related to the trend with binarity, as the fraction of combined binaries in the two metallicity samples are approximately equal (24/63 and 26/61 for low and high metallicity respectively).

There have been previous studies of the metallicity dependence of debris disc incidence [e.g.
Greaves et al., 2006, Saffe et al., 2008], motivated in part by the robust correlation between high metallicity stars and ‘hot jupiter’ planets [e.g. Santos et al., 2004]. These studies have showed no significant difference between the metallicity distributions of stars with and without debris discs. The results in this work are also statistically consistent with there being no trend of debris disc incidence with metallicity.

It must be noted that the measured chemical abundances and total metallicities for A type stars do not necessarily reflect their true (primordial) values due to poor mixing of the outer atmosphere.

Effective Temperature (Spectral Type)

The sample was split into two effective temperature subsamples with approximately equal numbers of observed systems by dividing about $T_{\text{eff}} = 8200$ K. For the single stars there is a $\sim 1.5\sigma$ distinction between the two $T_{\text{eff}}$ groups, with MIPS-24 and MIPS-70 excess rates of $39 \pm 12\%$ and $41 \pm 14\%$ for $T_{\text{eff}} \leq 8200$ K, and $60 \pm 14\%$ and $70 \pm 16\%$ for $T_{\text{eff}} > 8200$ K. For the binary stars, both combined and wide, the trend is the opposite, although with less significance due to the smaller detection rates and sample sizes. As a result there is little significance in the differences for the two temperature ranges for the sample as a whole.

The slightly increased excess rates for the hotter stars can be explained as a selection effect due to the dust mass sensitivity increasing with $T_{\text{eff}}$, as shown for the MIPS bands in Fig. 2.4.

Chemical Peculiarities

Subsamples of Am, Ap and $\lambda$ Boo peculiarities were created based on spectral types and a literature search. These subsamples are listed in Table 4.9. All other primary stars are considered ‘normal’. Interpreting the statistics for the peculiar stars is complicated due to the small numbers of stars (17 Am, 9 Ap and 4 $\lambda$ Boo), and degeneracies with other parameters.

The Am stars are the most numerous peculiar stars in the sample, however, they form a continuum of ‘Am-ness’, which can be defined as the difference between the Ca II K-line and metal-line spectral types [Gray and Corbally, 2009]. Separate Ca II K-line, metal-line and Hydrogen spectral types were only available for 6 of the 17 stars indicated as being Am, so it has not been possible to evaluate the effect of the strength of the Am pecularity on dust emission. The Am phenomenon is generally believed to be caused by slow stellar rotation. In the absence of rapid rotation there is very little mixing in the atmospheres of A type stars, so chemical separation can occur due to the different balance of radiation pressure and gravity for different ions. A primary cause of slow rotation is tidal braking in binary systems, so unsurprisingly a large fraction of Am stars are found to be in binary systems (9 out of 17 here, although this is a
lower limit). For single stars slow rotation either implies that the stars are old or were formed slowly rotating. For both the binary and single Am stars in this sample the excess statistics are essentially indistinguishable from the normal stars. In future, improved spectral classifications may allow more robust and meaningful conclusions to be drawn for the Am stars.

The Ap stars in this sample are mostly found in binary systems (7 out of 9), and their excess rates are consistent with those of the binary stars in general. It is not obvious why the binary fraction of Ap stars should be high, as the Ap peculiarities are generally attributed to chemical separation caused by strong magnetic fields [Gray and Corbally, 2009].

The four λ Boo stars in the sample are all considered to be single. These four stars have previously been studied with pre-Spitzer photometry in Paunzen et al. [2003], where the excesses of HD 125162, HD 31295 and HD 110411 were identified. It has been hypothesised that the λ Boo peculiarities could be due to accretion of circumstellar matter with the abundances of interstellar gas (i.e. minus the dust which contains the heavier metals). It must be noted that the debris discs of HD 125162, HD 31295 and HD 110411 all share essentially the same temperatures, masses and radii as derived from the MIPS photometry (Table 4.4), so it is tempting to suggest that the disc properties are related to the λ Boo phenomenon. HD 192640 does not fit this pattern, although it is not possible to rule out the presence of cold dust due to the MIPS-70 image being swamped by galactic cirrus emission. Generally the presence of only four λ Boo stars in the sample makes it impossible to do anything more than speculate on the relationship between the λ Boo phenomenon and circumstellar (or interstellar) dust.
Table 4.9: The chemically peculiar stars and their excesses. The spectral type references are as given in Table 3.3.

<table>
<thead>
<tr>
<th>UNS</th>
<th>Name</th>
<th>Spectral type</th>
<th>Ref.</th>
<th>Δ[24] (mag.)</th>
<th>χ24</th>
<th>Δ[70] (mag.)</th>
<th>χ70</th>
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**Am stars**

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<th>UNS</th>
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<th>Spectral type</th>
<th>Ref.</th>
<th>Δ[24]</th>
<th>χ24</th>
<th>Δ[70]</th>
<th>χ70</th>
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<tr>
<td>A013</td>
<td>HD 115892</td>
<td>kA1.5kA3mA3m A3 Va</td>
<td>gray06</td>
<td>0.191</td>
<td>6.7</td>
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<td>A015</td>
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<td>0.3</td>
<td></td>
<td></td>
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<td>A020</td>
<td>HD 130841 A</td>
<td>kA2hA5mA4 IV-V</td>
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<td>0.092</td>
<td>0.9</td>
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<td>A031</td>
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<td>A100</td>
<td>HD 198639</td>
<td>A4me...</td>
<td>HIP</td>
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<td>1.710</td>
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<td>A3MA3-A8</td>
<td>hook</td>
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<td>A120</td>
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<td>HIP</td>
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Mean: 0.109 | 0.373

**Ap stars**

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<th>Ref.</th>
<th>Δ[24]</th>
<th>χ24</th>
<th>Δ[70]</th>
<th>χ70</th>
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<tr>
<td>A010</td>
<td>HD 128898 A</td>
<td>A7 Vp SrCeEu</td>
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<td>0.049</td>
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<td>A024</td>
<td>HD 95418</td>
<td>A1 IVp (Sr II)</td>
<td>gray03</td>
<td>0.277</td>
<td>9.8</td>
<td>1.693</td>
<td>7.5</td>
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<tr>
<td>A027</td>
<td>HD 40183</td>
<td>A1 IV-Vp kA1mA1.5 (Sr)</td>
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<td>A033</td>
<td>HIP 12828 AB</td>
<td>A9 IIp</td>
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<td>A065</td>
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<td>A8 V: SrCeEu</td>
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<tr>
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<td>A0 II-HII SrEuCr</td>
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<td>A092</td>
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<td>A7P Sr</td>
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<td>0.329</td>
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<tr>
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<td>Ap (A0 IVm III)</td>
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<td>0.1</td>
<td>0.079</td>
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<tr>
<td>A112</td>
<td>HIP 11569 ABC</td>
<td>A5p Sr</td>
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<td>2.0</td>
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Mean: 0.050 | 0.228

**λ Boo stars**

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<th>Ref.</th>
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<th>χ24</th>
<th>Δ[70]</th>
<th>χ70</th>
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<tr>
<td>A053</td>
<td>HD 125162</td>
<td>A3 Va kB9mB9 λ Boo</td>
<td>gray03</td>
<td>0.438</td>
<td>14.2</td>
<td>3.166</td>
<td>30.1</td>
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<tr>
<td>A072</td>
<td>HD 31295</td>
<td>A3 Va kB9.5mA9.5 λ Boo</td>
<td>gray03</td>
<td>0.433</td>
<td>12.8</td>
<td>3.894</td>
<td>36.9</td>
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<tr>
<td>A076</td>
<td>HD 110411</td>
<td>A3 Va kB9.5mA0 λ Boo</td>
<td>gray03</td>
<td>0.463</td>
<td>15.3</td>
<td>3.461</td>
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</tr>
<tr>
<td>A116</td>
<td>HD 192640</td>
<td>A2V</td>
<td>BSC5</td>
<td>−0.030</td>
<td>−0.9</td>
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</tr>
</tbody>
</table>

Mean: 0.326 | 2.630
4.5.6 Interesting Objects

Of the 35 systems observed specifically for this work (out of 49 proposed), there are 12 with detected excess in at least one band, and five with excess in both (UNS A085, A113, A115, A117 and A123). With the exception of UNS A085 [HD 178253; Patten and Willson, 1991], all these detected discs are new discoveries. HD 37594 (UNS A115) stands out as an especially massive disc. A further interesting discovery, from observations performed by other authors, is a second debris disc in the CCDM 23480-2808 ABC (UNS 109) triple star system.

**HD 37594 (UNS A115)**

With the largest dust mass of any system in this volume limited sample – which includes such well known systems as β Pic, Vega, Fomalhaut, β Leo and β UMa (UNS A014, A003, A004, A003 and A024 respectively) – HD 37594 is a striking discovery. The excess from this star would have been easily within the detection limits of IRAS at 60 μm, but the low spatial resolution of IRAS (see §2.2.1) meant that it was swamped by emission from nearby nebulosity, as seen in the lower left corner of the MIPS-24 image in Fig. 4.20. The nebulosity, IRAS 05371-0338 / SFO 21, is a bright rimmed cloud [Sugitani et al., 1991], which is visible at optical wavelengths due to hydrogen Balmer emission as shown in Fig. 4.21.

HD 37594 was observed as a potential optical long-baseline interferometer calibration target by van Belle et al. [2008], however, they excluded it based on having an abnormal visibility distribution compared to known good calibrators. These abnormal visibilities are potentially due to circumstellar dust, as there are no indicators of HD 37594 being a binary system. Such observations have been used in other studies to detect circumstellar dust [e.g. Absil et al., 2008].

Spectroscopically, HD 37594 is seen as a low $v \sin i$ A8–9 V spectral type star with approximately ‘normal’ element abundances [Houk and Swift, 1999, Bikmaev et al., 2002]. The observed $v \sin i$ of 17 ± 2 km/s [Bikmaev et al., 2002] is especially low for normal A type stars, with 100–200 km/s being typical for an A7–A9 type star [Royer et al., 2007]. As HD 37594 doesn’t exhibit Am type abundance peculiarities, which are thought to be a consequence of slow rotation, it is likely that the star is observed pole-on ($i \lesssim 10^\circ$ if $v \geq 100$ km/s). Bikmaev et al. [2002] have noted that the spectral line profiles of HD 37594 have a slightly peculiar shape, which is also seen in δ Sct pulsating stars. The photometric variation measured by Hipparcos in only $\sim 0.008^{\text{m}}$ [F. van Leeuwen, 2007], so HD 37594 is at most a mild δ Sct pulsator.

The dust temperature and orbital radius determined from the MIPS [24] – [70] colour and stellar luminosity are $T = 54^{+3}_{-2}$ K and $r = 64^{+6}_{-5}$ AU (Table 4.4). This orbital radius assumes perfect black body dust grains, and it is possible that this $r$ is an underestimate by as much a factor of five or more [Fig. 4.15 and Bonsor and Wyatt, 2010]. The distance to HD 37594
**Figure 4.20:** MIPS images of HD 37594. Left: MIPS-24 post-BCD image displayed with log scale, right: MIPS-70 filtered and corrected post-BCD image displayed with linear scale. North is up, and both images are at the same scale. The bright-rimmed cloud SFO 21 is seen at the bottom-left of the MIPS-24 image.

is $42.55 \pm 0.67 \text{ pc}$ [F. van Leeuwen, 2007], so a lower-limit radius of 64 AU would result in a minimum angular diameter of $3''$. Even though the angular size was expected to be small compared to the $\sim 20''$ FWHM MIPS-70 PSF, the very high S/N in the MIPS-70 image has allowed the marginally resolved nature to be confirmed by PSF subtraction. Fig. 4.22 shows the MIPS-24 and MIPS-70 images after subtracting the PSF model scaled such that the resulting central pixel value is equal to the average background value. The low dust flux in the MIPS-24 band results in there being relatively little to see, however, in the MIPS-70 image, where almost all of the observed flux is from the dust, there is seen to be considerable extension. It must, however, be noted that some of the extension will be due to the redder than stellar spectrum, and the considerable spectral width of the MIPS-70 band.

The low temperature, large dust mass, and angular extent of several arcseconds, make this a potentially very interesting target to follow-up at longer wavelengths. Unfortunately HD 37594 is not included in any of the present *Herschel* debris disc surveys, due to the *IRAS*-based confusion estimates used to exclude targets indicating that this region would be highly confused. HD 37594 is, however, included in the SUNS survey with SCUBA-2, which will allow the dust mass to be more reliably determined. $450 \mu\text{m}$ observations with SCUBA-2, with a
**Figure 4.21:** 15 × 15′ Composite of UK Schmidt Telescope B and R band images obtained from the SuperCosmos Sky Survey (http://www-rfau.roe.ac.uk/sss/), centred between HD 37594 and SFO 21. The fields of view of the MIPS observations are overlaid. The nebulosity of SFO 21 is clearly seen in hydrogen emission.

FWHM beamsize of 7′, will potentially allow the size of the disc at sub-mm wavelengths to be measured. HD 37594 is potentially an excellent target for the ALMA sub-mm interferometer, although the angular extent is likely to be on the upper limit of the usable field of view (the antenna beams have FWHM beam sizes of 18′ at 850 μm, and there will be pointing uncertainties of a few arcseconds).
4.5. RESULTS

**Figure 4.22:** PSF subtracted MIPS images of HD 37594 showing resolved nature, especially for MIPS-70. Left: MIPS-24, right: MIPS-70. Both images are $160 \times 160''$ in size, and have the instrumental orientation. The subtracted PSF flux has been chosen to yield approximately the image background value at the centre of the image after subtraction.

**CCDM 23489-2808 ABC (UNS 109)**

This triple system (CCDM 23489-2808 ABC, UNS 109, HD 223352 AB + 223340) is shown in this work to contain at least two circumstellar discs (see Tables 4.4 and 4.7), with a circum-primary disc around the A component, and a circum-tertiary disc around the C component. The primary disc has previously been noted in Morales et al. [2009], however, the C component appears to have gone unnoticed. Fig. 4.23 shows the MIPS-24 and MIPS-70 images, with the AB and C components indicated.

The primary star is of A0 spectral type, with the B component, of late K or early M type, separated on the sky by $\sim 3.9''$ ($\geq 164$ AU). The C component, separated by $\sim 75''$ ($\geq 3150$ AU), is of K1 V spectral type [Houk, 1982], with effective temperature of 5270 K measured here from the photosphere flux distribution fit. It is interesting to note that the dust mass implied from the MIPS photometry (Table 4.4) for the C component is a factor of three larger than for the primary, which is likely due to the outer edge of the circum-primary disc being truncated by the B component.

This system will be observed by both the DEBRIS *Herschel* survey and the SUNS survey with SCUBA-2, which will allow the outer disc properties and the dust masses to be reliably determined. Probing the truncation of the circum-primary disc is a task ideally suited to ALMA, with the maximum size scale of $\sim 8''$ well matching the usable ALMA field of view at 850 $\mu$m.
Figure 4.23: MIPS images of the CCDM 23489-2808 ABC (UNS 109) system. Left: MIPS-24 post-BCD image displayed with linear scale, right: MIPS-70 filtered and corrected post-BCD image displayed with linear scale. North is up, and both images are at the same scale. The AB–C separation vector is $\rho = 74.64''$, PA = 296.5° (from TDSC).
4.6 Summary

This chapter has presented a volume limited survey of 130 A type star systems at 24 and 70 μm using the MIPS instrument on the Spitzer space telescope. The target systems were the A type sample presented in chapter 3 of this thesis and Phillips et al. [2010]. Observations of 49 systems were proposed in November 2007 [Phillips et al., 2008], to complement previous or already scheduled observations of the other 81 members of the sample. All systems observed in either MIPS-24 or MIPS-70 bands are included here. As far as possible all data reduction, photometric measurements and analysis have been performed in a homogeneous manner. Photometry measured elsewhere has only been used here for five systems observed in an uncommon MIPS-70 observing mode. In total MIPS photometry is presented here for 116 systems at 24 μm and 102 systems at 70 μm. Where possible PSF fit photometry has been used to maximise S/N, minimise the effects of contamination and confusion, and resolve stars in multiple systems.

In order to determine the presence of excess flux due to emission from dust, the photometry for the stellar photospheres was predicted using model photosphere flux distributions. Observed optical and near-IR photometric colours were fitted by \( \chi^2 \) minimisation to a grid of synthetic colours computed for photosphere models with varying \( T_{\text{eff}} \), \( \log g \) and \([M/H]\). The MIPS photosphere photometry was then computed using selected near-IR and optical photometry and synthetic colours from the best fit \( T_{\text{eff}} \), \( \log g \) and \([M/H]\) model. Despite various complications for the systems in this survey – including saturation in modern photometric surveys such as 2MASS, inconvenient binary separations, chemical peculiarities and rotational oblateness – prediction uncertainties of \( 1\sigma \sim 0.02^m-0.04^m \) were achieved, equal to the best previously achieved for A type stars [e.g. Su et al., 2006].

The observed and predicted MIPS photometry, and the stellar parameters from the photosphere model fitting, have been presented (Table 4.3). This has significant legacy value, as these results are directly usable in the analysis of observations performed by the DEBRIS and SUNS surveys which share the same A type sample as this work. The significance of photometric excesses were determined, and stars with \( \geq 3\sigma \) excess in either band were identified. Twelve new debris discs were discovered. The [24] – [70] colour of the excess emission was used in combination with stellar luminosities to determine characteristic dust temperatures, orbital radii and masses (Tables 4.4 and 4.5). For the multiple star systems with excess, the configuration of the dust orbits relative to the stars (circum-binary vs. circum-primary etc.) were determined using the star separations and the characteristic dust orbital radii (Table 4.7).

The volume limited and hence largely unbiased nature of this survey has allowed a meaningful statistical investigation of debris incidence rates with various system parameters. For the sample as a whole the incidence rates of significant (\( \geq 3\sigma \)) excess at 24 and 70 μm are \( 33 \pm 5\% \)
and 41 ± 6% respectively, with lower limits, assuming all non-observed and non-detected systems do not possess excess, of ≥ 29 ± 5% and ≥ 32 ± 5% respectively. These global detection rates were shown to be in perfect agreement with rates determined by Su et al. [2006] for a differently selected sample of A type stars (although many of the systems observed here are also included in the sample studied by Su et al. [2006]). It has been shown that the excess incidence rates are higher for hotter stars within the sample (~1.5σ significance), although this is likely to be a selection effect due to the influence of stellar luminosity on dust mass sensitivity.

The excess rates have been shown to be essentially independent of metallicity, which is in agreement with previous works [Greaves et al., 2006; Saffé et al., 2008], and quite distinct from the established correlation between radial velocity detected planets and high stellar metallicity. Both the giant planet – metallicity correlation, and the non-correlation of the presence of debris with metallicity support the core accretion model of planet formation. In this model essentially all systems would be expected to build planetesimals, which are the source of the observed debris, but only systems with large dust masses (i.e. high metal content) would be expected to be able to form sufficiently massive cores to accrete gas before the gas is dispersed from the protoplanetary disc (see §1.2.2).

The dominant system property affecting the excess detection rates was found to be binarity. A thorough literature search was performed to identify multiple stars, in addition to the work in chapter 3. The proximity of the stars in this sample means that the information on binarity is as complete as possible for stars of this type (with the exception of spectroscopic and eclipsing binaries, sensitivity to binarity depends on distance). The detection rates for binary systems with separations below a few hundred AU (a threshold determined by practical considerations of MIPS-70 and legacy photometry resolutions) have been shown to be lower than single stars and wider binaries, with a significance of at least 3σ. This disagrees with a previously published, and much cited, work of Trilling et al. [2007], which claimed that debris discs are more commonly detected around binary stars than single stars. It has been shown here that the result of Trilling et al. [2007] is due to their use of different excess detection criteria for single and multiple stars (they perform their own analysis for binary stars, and compare their results with the analysis of Su et al. [2006] for single stars). The results here are free from such bias, and also free from any possible bias due to different selection criteria for the samples of multiple and binary star systems.

Approximate orbital separations for the binary systems in the sample were compiled from the literature. This showed there to be a noticeable lack of excess detections for systems with separations of approximately 3–150 AU. This agrees with what was found by Trilling et al. [2007] for their sample of mid-A and F type binary stars. Statistical tests comparing
the distributions of systems with and without excess as a function of separation have shown that the distributions are significantly different. It is, however, hard to rule out this being a selection effect. The dust detected in the systems with separations of $\lesssim 3$ AU is generally circum-binary, whereas the dust in the systems with separations of $\gtrsim 150$ AU generally orbits one of the components (circum-primary, circum-secondary, circum-tertiary). It is conceivable that there is circum-binary dust around systems with separations of $\sim 3$–150 AU, but that the low temperature makes detection with MIPS unlikely. Whether such large circum-binary discs are common is a question for Herschel and sub-mm surveys.

In summary, this chapter presents the pre-Herschel debris disc incidence statistics for the 130 A type systems within $\sim 45$ pc of the Sun. The photometry and other data compiled in this work is a significant contribution to the DEBRIS and SUNS surveys with Herschel and SCUBA-2, which will observe large subsets of the sample studied here (83 and 100 systems respectively). Whilst significant statistical conclusions and discoveries of interesting systems have been made by this survey, the need for longer wavelength observations to minimise selection effects and better determine dust temperatures, masses and orbital distances has also been highlighted.
Chapter 5

LABOCA Data Reduction &
Observations of Southern
Circumstellar Discs

5.1 Introduction

This chapter describes the reduction of data from the Large APEXBOlometer CAmera [LABOCA, Siringo et al., 2009] instrument, and the application of this to determine the dust masses and (sub-)millimetre SED properties of circumstellar discs observed in support a large *Herschel* Key Programme.

LABOCA is a bolometer array with 295 DC-coupled bolometers, arranged in a 12′ diameter hexagonal grid of 2 × FWHM spacing, operating at an effective wavelength of 870μm. Since May 2007, LABOCA has operated at the cassegrain focus of the 12 m Atacama Pathfinder EXperiment (APEX) telescope\(^1\), situated at 5105 m altitude on Llano de Chajnantor in the Atacama desert in Chile. This site exhibits an exceptionally dry atmosphere, with consistently low opacity during the months April to December [Otárola et al., 2005]. An overview of the instrument, its performance, and typical observing and data reduction procedures is given in Siringo et al. [2009]. The FWHM beam size of LABOCA is approximately 19″, and the typical point-source sensitivity (1σ) obtained in one hour of the most efficient scanning mode is ~5 mJy.

The observations presented in this chapter were taken to complement the Gas in Proto-planetary Systems (GASPS) *Herschel* Key Programme [Mathews et al., 2010, PI: Dent], which is described below. With the exception of the original observing proposals (PIs: Dent and

\(^1\)http://www.apex-telescope.org/
Liseau), the author has solely been responsible for the whole project. This has included the
observation planning (phase two) and data reduction. During the data reduction an improved
scan reduction pipeline for faint compact sources was written, which reduced noise in maps by
up to a factor of two over the example scripts which were shipped with the reduction software.
A problem with the instrument which affected one of the two observing runs was identified,
and as far as possible has been corrected for.

5.1.1 The GASPS Survey

GASPS aims to characterise the gas in discs of ages 1–30 Myr, a range which spans the epochs
of gas giant planet formation, and the transition from gas-rich protoplanetary discs to gas-
poor debris discs. To understand the gas in discs it is necessary to also understand the dust
properties, as the dust dominates the radiative transfer within discs. The ratio of gas to dust
mass in discs is often a matter of great debate and uncertainty, as the evolution of dust and gas in
discs can potentially alter the ratio by orders of magnitude from the typical ISM ratio of \( \sim 100 \).
Until recently, the only readily detectable gas features from discs have been sub-mm rotational
transitions of molecules such as CO, HCN and HCO\(^+\), which are primarily emitted from the
cold outer regions of discs. Total gas masses derived from these lines are highly uncertain due
to significant optical thickness, freeze-out in the cold disc midplane, and dissociation in the
upper disc layers [Kamp et al., 2010]. **Herschel** offers the ability to observe other gas features
in the far-IR, which are emitted from warmer regions at radii of \( \sim 1–100 \text{ AU} \) in discs.

The primary aims of GASPS are to:

- Trace gas and dust in the planet formation region of discs across an extensive multivariate
  parameter space
- Provide the first definitive measurement of the gas dissipation timescale
- Study the evolutionary link between protoplanetary and debris discs
- Investigate the extent of warm H\(_2\)O in the planet forming regions of discs, with implica-
tions for the volatile content of developing planets
- Provide an extensive database of disc observations and models with long lasting legacy
  value for follow up observations

GASPS targets a total of 240 systems, the majority of which are members of associations with
ages of 1–30 Myr within 200 pc of the Sun (Taurus, Upper Scorpius, \( \eta \) Chamaeleontis, TW
Hydra, \( \beta \) Pictoris and Tucana-Horologium). Several isolated HAEBeS with ages of 0.1–30 Myr
are also included. The targets are generally not embedded within molecular clouds and have
Lada SED classes of II-III. Stellar masses are approximately in the range 0.5–4$M_\odot$. Target systems were chosen to possess discs spanning a wide range of evolutionary states, based on previous near-IR to millimetre photometry and accretion indicators. Some association members with no previously detected disc are also included.

Systems are typically to be observed with spectroscopy of fine structure lines of OI (neutral oxygen atom) at 63.2 and 145.5 $\mu$m, and CII (singly ionised carbon atom) at 157.7 $\mu$m, as well as broadband photometry at 70 and 160 $\mu$m. Systems with [OI] or [CII] detections will be followed up with spectroscopy of lines of CO, H$_2$O, OH, CH$^+$ and DCO$^+$ in the wavelength range 70-160 $\mu$m. The [OI] and [CII] lines are the most readily detectable far-IR spectral features of gas in circumstellar discs, and their line fluxes and ratios can be used to directly determine disc gas masses and outer radii. For example, $[\text{OI} 63]/[\text{OI} 145]$ is sensitive to gas mass and is insensitive to outer radius, whereas $[\text{CII} 158]$ is highly sensitive to the outer radius [Kamp et al., 2010].

The *Herschel* observations, combined with ancillary observations such as those in this chapter, are to be used to fit detailed disc models in order to determine realistic disc parameter constraints (e.g. gas mass, gas:dust ratio, radial structure and outer radius). Photometry at (sub-)millimetre wavelengths is of particular importance for modelling the dust properties, especially for constraining the properties of the grain size distribution and measuring the total mass in dust grains up to sizes of a few millimetres.
5.2 LABOCA Data Reduction

All observations with LABOCA are taken in a scanning mode where the telescope is continuously moving\(^2\). Scanning is used to modulate the signals of compact astronomical objects to frequencies above the 1/\(f\) knee\(^3\), to reduce dependence on individual bolometers, and to provide dense sampling and even coverage of fields of various sizes. In this work three scan patterns based on a raster of spirals are used. The positional offsets of the array in azimuth and elevation (AzEl) coordinates as a function of time for these patterns are shown in Fig. 5.1. The compact raster spiral pattern (Fig. 5.1, top) is used for all observations with the exception of the \(\eta\) Chamaeleontis cluster, for which scans in the larger area patterns are also used. The spiral patterns combine high time efficiency with dense spatial sampling.

All routine observations with APEX are performed in service mode. The data products which the astronomer receives are time series data of the sampled bolometer voltages and other instrument parameters for each scan, in MB-FITS format\(^4\). The data reduction process for each scan generally involves processing of the time series data, and rendering a map from the time series data. The maps from individual scans are then stacked by weighted averaging, and analysis such as photometry is then performed on the stacked maps.

All the data reduction is performed in the Bolometer array data Analysis (BoA) software package\(^5\), which is written in a combination of Fortran90 and Python languages. The extensive use of Python makes BoA inherently script-able, and the code is clear, and easily modified and extended. BoA is shipped with several scripts which provide instrument specific functions, and pipelines for the reduction of various types of LABOCA scans. During the course of this work, several improved pipeline scripts have been developed based on the shipped scripts. Some of the improvements are primarily for convenience, such as automatically dealing with array parameters and flags based on the observing date, however, improvements in the reduction of science scans of weak sources has resulted in significantly reduced noise. For science observations, which usually consist of several scans, it is typical to write a script which runs an appropriate pipeline script on each scan, produces maps, and stacks them. The version of BoA used throughout this thesis is revision 2473 (2008-01-17).

The following subsections describe the reduction process from the raw time-series data to stacked maps, and photometric measurements on the maps.

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\(^2\)A chopping mode for point-sources, using the ‘wobbler’ secondary mirror has also been commissioned in 2010

\(^3\)The power spectrum of instrumental noise is generally well approximated by white noise (no frequency dependence) plus noise with a 1/\(f\) spectrum. The ‘knee’ refers to the frequency below which deviation from white noise becomes significant.

\(^4\)\url{http://fits.gsfc.nasa.gov/registry/m勃fits.html}

\(^5\)\url{http://www.apex-telescope.org/bolometer/laboca/boa/}
Figure 5.1: LABOCA scan patterns used in this work. Top: compact raster spiral pattern, which uses four spirals with small offsets to achieve very dense spatial sampling. Bottom-left: $3 \times 3 + 2 \times 2$ raster spiral pattern giving approximately $15' \times 15'$ square with even depth. Bottom-left: $4 \times 4 + 3 \times 3$ raster spiral pattern giving approximately $25' \times 25'$ square with even depth. The patterns shown in red and black in the bottom two plots are taken as separate scans.

5.2.1 Flagging

Alongside the time series data values, BoA stores a set of flags for each channel, each time-stamp (set of samples with the same time), and each sample within each channel. These flags are primarily used to indicate that particular data should be ignored by following processing routines.

One of the first steps in the reduction process is flagging channels which do not contain valid astronomical data. These are primarily channels which are not connected to bolometers (there are 320 channels in total), or are connected to bolometers which are not functioning at
all, are very noisy, or exhibit strong cross-talk with other bolometers. Usually $\sim 50$ of the 295 bolometers are flagged in this manner. For some observations used in this thesis, a number of bolometers need to be flagged due to a support structure in the receiver cabin shadowing them from the incident radiation. In general, lists of channels to be flagged are downloaded from the APEX website\(^6\), and they can be automatically loaded along with a corresponding array flat-field, by using the start time of each scan to determine which files to use. The channel flag files from the APEX website tend to span quite a long period in time (nominally a month), so it is often worth inspecting the signal in science scans from bolometers listed as intermittently bad to see if they are working at the time of each scan.

Times when the telescope scan speed is excessively low or high, or when the acceleration is too high, are flagged and are not used at all during reduction. Data taken at such times suffer from excessive $1/f$ noise, beam smearing, and poor pointing accuracy respectively. In the compact raster spiral observing mode used for the majority of observations here, this results in the loss of only $9\%$ of samples.

### 5.2.2 Calibrating the Time Series Data

The first step in the time domain reduction is to scale the data to calibrated units of flux density. This process is broken down as follows:

- Scale data from ADU to Volts using the back-end gain
- Scale from Volts per beam to Janskys per beam by using a nominal factor of $6.3 \text{Jy}/\mu\text{V}$ [Siringo et al., 2009]
- Scale the data for each bolometer based on its sensitivity relative to the nominal value by using an array flat-field
- If required, scale the data by a constant determined from observed calibrator fluxes (the instrumental calibration is generally stable to within $10\%$ [Siringo et al., 2009])
- Correct for atmospheric transmission by multiplying the data by $\exp(\tau_z \text{sec} z)$ using the median elevation of the scan

The first three of these are handled automatically within scan reduction pipeline scripts; however, as the latter two points require the calibration correction and $\tau_z$ to be specified, these operations are performed for each scan before running a scan reduction pipeline script.

An extra correction can be applied to compensate for the effect of temperature variations of the array (strictly the $^3$He cooler temperature). Varying bolometer temperatures cause signals

\(^6\)http://www.apex-telescope.org/bolometer/laboca/calibration/array/
which are indistinguishable from variations in incident radiation. The temperature variation is primarily correlated with telescope elevation, as LABOCA is mounted at the cassegrain focus and thus tilts with the telescope. During science scans the temperature varies very little due to the small range of elevations covered and the relatively short timescales, so correction for this is not necessary. If there is any effect in science scans, it will be removed during the reduction process anyway by common-mode noise removal and baseline subtraction. The only scenario in which the temperature drift does need to be corrected is for sky-dip scans. For these, the telescope slews through a large range in elevation, and it is the total power received from the sky which must be measured.

5.2.3 Correlated Noise Removal

The vast majority of the signal produced by the bolometers is due to the brightness of the Earth’s atmosphere, which is essentially uniform over the field of view at any instant. Other signal variations caused within the instrumentation are also common between channels. Due to LABOCA’s design there are three main scales on which noise is correlated between channels:

- Whole array

- 12 groups of 26 bolometers, which are each connected to a single board of bias resistors and a board of low temperature current amplifiers

- 4 boxes of 3 groups of 26 bolometers, which correspond to boxes containing 3 boards of low temperature amplifiers. Each cold amplifier box feeds a separate warm amplifier unit, which in turn feeds a data acquisition unit.

The process of removing the correlated noise from a selection of channels (group, box or all) consists of computing the gain of each channel to the average signal of the selected channels, and then from each channel subtracting a fraction of the average signal proportional to the gain of the channel to the average signal.

\[
F[\text{channel}, t] = F[\text{channel}, t] - fG[\text{channel}](F)[t], \quad 0 < f < 1,
\]

(5.1)

where \(G\) is the channel gain and \(f\) is typically set close to one. It is necessary to compute gains, as the sensitivities to large common mode signals, which can be introduced at various points within the instrument, are not generally the same as the sensitivities to small astronomical signals, which is what these data have been calibrated for (see above). Within the reduction scripts shipped with BoA, this process is iterated between 2 and 5 times, with \(f = 0.8\).
The complete correlated noise treatment in the default reduction scripts for weak sources (\(<1\,\text{Jy}\)) is as follows,

1. flag excessively noisy channels \((\sigma_{\text{channel}} > 5\sigma)\)

2. whole array: 5 iterations, \(f = 0.8\)

3. despike data by \(\sigma\)-clipping each channel to \(\pm3\sigma\) (non-iterative)

4. boxes: 2 iterations, \(f = 0.8\)

5. groups: 2 iterations, \(f = 0.8\)

It was found that the S/N could be improved by modifying this algorithm. The primary changes are the addition of despiking (\(\sigma\)-clipping) in between iterations of correlated noise removal, and repeating the whole process a number of times. The motivation for the careful despiking and iteration of the process is that after each stage of correlated noise removal the data get flatter in time. Initial despiking passes, when the data contain large low frequency common-mode components, tend to flag valid data at large scale maxima and minima, and miss actual spikes as the \(\sigma\) is high. Before each despiking, all samples previously flagged as spikes are unflagged. The despiking alternates between non-iterative and iterative \(\sigma\)-clipping on successive iterations (non-iterative first for each scale). The result is to allow more reliable mean signals to be computed and subtracted on the passes with iterative \(\sigma\)-clipping, whilst still making maximal use of samples which occur around common-mode transients which would get clipped in the iterative \(\sigma\)-clipping. The measured improvement from this algorithm is between a factor of 1.2 to 2.0 lowering of the RMS in maps. The only major downside to the improved pipeline is an increase by a factor of \(\sim 5\) in the computing time for each scan, however, for the reduction required in this thesis this was only a minor inconvenience. Large scale structure which would also be correlated between bolometers may be removed to a greater extent than in the default pipeline, however, both pipelines are detrimental to such structure. It is conceivable that there are systematic effects on bright compact sources. Comparison of photometry produced using this pipeline with photometry from the pipeline used for flux calibrator observations, for the bright Herbig Ae/Be sources in this chapter, shows little systematic effect. In fact, the despiking before correlated noise removal iterations, followed by unflagging the spike samples afterwards, behaves in a similar manner to the flagging and unflagging by proximity to the source position in the calibrator reduction pipeline.
5.2.4 Frequency Domain Filtering

The channels of LABOCA exhibit $1/f$ noise, which becomes significant below 0.1 Hz [Siringo et al., 2009]. The duration of raster spiral scans is approximately 450 s, so frequency components down to $\sim 0.004$ Hz are sampled. To suppress $1/f$ noise, for observations where large scale structure is not of interest, BoA provides routines for altering the noise power spectrum. In this work, a whitening filter is used when such filtering is acceptable. This filter scales Fourier components below a threshold frequency such that they have the same power as components at a reference frequency. The threshold and reference frequencies used here are 0.30 and 0.35 Hz. These are chosen based on the slowest scanning speed in the spiral mode of 3.3 beams/s.

The benefits of the whitening filter after using the improved correlated noise removal outlined above are relatively small, suggesting that the majority of the $1/f$ noise is correlated between channels. The improvement due to whitening is typically to reduce the map RMS by 5–30%.

5.2.5 Baseline Subtraction

A polynomial baseline is typically fit to, and subtracted from, the data for each channel. This essentially removes the lowest frequency components from the data. In this work a polynomial of order one is used, which is also used in the default reduction scripts.

5.2.6 Final despiking ($\sigma$-clipping)

The final time domain processing is another despiking by $\sigma$-clipping. In the improved weak-source reduction pipeline developed in this work, this is an iterative $\sigma$-clipping. The thresholds need to be chosen with care to avoid throwing away an excessive fraction of samples and clipping the flux from bright sources. For maps containing only sources of interest fainter than $\sim 100$ mJy, thresholds of $\pm 3\sigma$ were found to be optimal, typically only clipping 1.5% of samples. For maps containing brighter sources of interest, these limits need to be increased. For the brightest sources in this work, thresholds of $-4\sigma$ and $+9\sigma$ have been used (note that these asymmetric limits produce maps with non-zero mean).

5.2.7 Map Projection

After the time domain reduction, each channel is assigned a weight of $1/\sigma^2$, where $\sigma^2$ is the variance of the time series for the channel. The samples for each channel are multiplied by the channel weight. A map is constructed by starting with two blank maps—an intensity map and a weight map. For each sample from each channel, the corresponding map pixel coordinate is computed using the array AzEl offset for the time of the sample, and the relative bolometer
offsets within the array. The value of the sample (which has been weighted) is then added to the corresponding pixel in the intensity map, and the channel weight is added to the same pixel in the weight map. BoA also provides the option of distributing the sample value and weight amongst the four closest pixels, based on their distances from the real sample position (the neighbor option). Once all samples have been plotted like this, the intensity map is normalised by dividing by the weight map. The pixel values in maps produced like this have the same unit as the time series data, which is Jy/beam in this case. With this normalisation, the peak value of a point source in a map is equal to the total flux density of the source.

Maps can either be produced in AzEl coordinates, or in equatorial coordinates by using the telescope pointing. Projection in AzEl coordinates is generally only suitable for individual scans of relatively short duration, due to field rotation. The BoA scripts for reducing and analysing calibration observations use AzEl maps, as they meet these criteria. For science observations where maps from multiple scans need to be stacked, an equatorial projection is necessary. For equatorial maps, BoA uses a Sanson-Flamsteed projection (GLS in FITS terminology). This is an equal area projection, in which parallels (lines of constant declination) are horizontal, equally spaced, and have their true length ($\propto \cos \delta$) [Calabretta and Greisen, 2002].

The pixel size in maps can be chosen arbitrarily. The BoA scripts for calibration observations use a pixel size equal to a quarter of the beam FWHM assumed by BoA, which is 18.22'''. The calibration maps are produced with the neighbour option turned off, which can result in a small number of blank pixels for short scans which only provide a fairly sparse spatial sampling. Maps with a pixel size of 1'' produced with the neighbour option enabled, have also been extensively used in this work (the ~1 pixel smoothing due to the neighbour option has negligible effect on the effective beam size with this pixel scale).

5.2.8 Map Smoothing

A common technique applied to maps is to smooth them by convolving with a Gaussian kernel. This allows arbitrarily small pixel sizes to be used, as the noise is determined by the size of the smoothing Gaussian rather than the pixel size. Smoothing results in an increase of the beam size of a map. If the instrumental beam is well approximated by a Gaussian (as is the case for LABOCA), then the effective beam size in a smoothed map is the quadrature sum of the instrumental and smoothing kernel beam sizes. In this work a beam size for LABOCA of FWHM = 18.6'' is generally assumed, and maps have often been smoothed with Gaussians of the full, and half, the beam size, resulting in effective beam sizes of FWHM = 26.3'' and 20.8'' respectively.

7The value of 18.6'' comes from the commissioning report for LABOCA, which states a value of 18.6 ± 1.0''. More recently a value of 19.2 ± 1.0'' has been stated in Siringo et al. [2009].
Care must be taken with the normalisation of the smoothing kernel. If a kernel of unit integral were used then the flux calibration of the map would remain in units of flux density per *instrumental* beam. Thus the kernel or the map must be multiplied by the ratio of the effective to the instrumental beam sizes to produce a map in units of flux density per *effective* beam. This preserves the condition that the peak value in the map for a point source is equal to the total flux density of the source.

The case of smoothing with a kernel equal to the instrumental beam corresponds to matched filtering, which can be shown to be the optimum linear filter for maximising the S/N. For this reason, maps of marginally detected point sources are typically smoothed with a full beam sized Gaussian for display.
5.3 Calibration Observations

A number of types of calibration observations are required to make use of the science data.

5.3.1 Pointing and Focus

Pointing and focus observations are made regularly throughout each night, typically at \( \sim 2 \) hour intervals, and corrections are applied after each when significant. For LABOCA, the pointing observations are spiral scan maps, just like the majority of science observations. The fitted position of the pointing object in a map, projected in AzEl coordinates, is used as a correction to the telescope control system’s pointing model. The pointing model gives rms pointing errors of \( 2^\prime \) and \( 4^\prime \) in azimuth and elevation respectively\(^8\), and with the use of pointing calibration observations the pointing for science observations is generally correct to within \( 1-2^\prime \).

Focus observations are performed with one bolometer staring at a source which is significantly brighter then the atmosphere (only Mars, Jupiter, Saturn and Uranus), or chopping on a fainter object using the telescope’s wobbling secondary mirror (until 2010 this was the only LABOCA observing mode using chopping). Subscans are taken in 10 focus positions, and by interpolation, the position yielding the peak flux is determined. The pointing and focus corrections are applied at the telescope, and hence are generally not required for later data reduction.

5.3.2 Sky-dips & Radiometer PWV Measurements

For ground based (sub-)millimeter observations flux calibration is especially complicated due to the limited, and often highly variable, transmission of the Earth’s atmosphere. Although continuum observations will usually not be performed when the atmospheric conditions are especially variable, it is still typical for the transmission to vary significantly on timescales of a few scan durations. As it is extremely rare to have calibration sources within the science field that can be significantly detected in a small number of scans, other means must be used for calibrating on the necessary timescales. Observations of bright flux calibrators are an important part of the calibration procedures, but they can only be performed infrequently without seriously reducing observing efficiency, and the need for an air-mass correction to the science field elevation requires the atmospheric effects to be disentangled from the instrument sensitivity.

The primary method of characterising the effects of the atmosphere is through sky-dip observations, which measure the sky brightness temperature as a function of elevation. With

\(^8\)http://www.apex-telescope.org/bolometer/laboca/calibration/
5.3. CALIBRATION OBSERVATIONS

LABOCA this is achieved by first performing a calibration scan which measures the sky brightness followed by a warm load of known temperature and emissivity. A second scan is then performed where the telescope tips continuously through a large range in elevation. Using the brightness temperature calibration from the first scan, and compensation for variation of the $^3$He temperature in LABOCA with elevation, the second scan yields the sky brightness temperature as a function of elevation. This is then fit to a model incorporating the temperature of the telescope and atmosphere, the total optical coupling and forward efficiency, and the opacity as a function of air-mass. It is acceptable to use a single atmospheric temperature model, as the majority of water vapour occurs at low altitude, so the effects of higher altitude, lower temperature layers can be neglected. As well as providing the opacity at zenith, $\tau_z$, which can be used for calibrating scans as in § 5.2.2, comparison of the observed brightness temperature profile to the fitted model is a useful diagnostic for the stability of the atmosphere.

The model which is used within BoA to fit the sky-dips is,

$$T_{\text{sky}}(z) = (1 - \eta_{\text{opt}})T_{\text{cabin}} + \eta_{\text{opt}}T_{\text{atm}}(\eta_F (1 - e^{-\tau_z \sec z}) + (1 - \eta_F)), \tag{5.2}$$

where $T_{\text{sky}}$ is the measured sky brightness temperature, $\eta_{\text{opt}}$ is the optical coupling ($\sim 80\%$), $\eta_F$ is the forward efficiency ($\sim 97\%$), $T_{\text{cabin}}$ is the receiver cabin temperature, which is assumed to equal the telescope temperature, and $T_{\text{atm}}$ is the physical atmosphere temperature.

Due to the large slews involved in performing sky-dips, they take several minutes and can potentially disturb the telescope pointing. Consequently, sky-dips are usually only performed roughly once per hour, along with other calibration observations.

To allow monitoring of the atmosphere on shorter timescales without negatively impacting observing efficiency, a dedicated instrument is used to frequently perform sky-dips and measure $\tau_z$. At APEX, a tipping radiometer operating at bands either side of the 183.3 GHz (1.64 mm) line of H$_2$O is used$^9$. The radiometer is aligned in azimuth with the telescope and performs sky-dips once per minute. The radiometer data are fitted to an atmospheric model for Chajnantor to determine the precipitable water vapour (PWV) above the site in mm. This is archived and made available on-line$^{10}$, and in the observing log shipped with the data from each observing run. Under the assumption that the variability in the opacity at the observing frequency is dominated by water vapour, $\tau_z$ can approximately be determined by a linear relationship from the PWV. This relationship is instrument and site dependent due to the instrument spectral response, and opacity due to molecules other than water, which is assumed to be constant or vary directly with PWV.

$^9$http://archive.ess.org/ess/meteo_apex.html
$^{10}$http://archive.ess.org/web/web/ess/meteo_apex/form
A linear fit of \( \tau_z(870 \mu m) \) to the archived PWV measurements, determined from LABOCA sky-dips and flux calibrator observations performed on the nights of observations used in this work (including those in the following chapter), is shown in Fig. 5.2. The adopted relationship is,

\[
\tau_z(870 \mu m) = 0.06 + 0.32(\text{PWV}/\text{mm}).
\]  

(5.3)

Only sky-dips where the sky brightness profile is well fitted by the sky-dip model have been used. This relationship is used to compute \( \tau_z(870 \mu m) \) on a scan by scan basis from the archived PWV values during data reduction.

**Figure 5.2:** Empirical \( \tau_z(870 \mu m) \) vs. PWV relationship from flux calibrators and sky-dips observed around the times of observations used in this work. Only sky-dips with good fits have been included. Points from calibrators observed with very low PWV are given reduced weighting as they are very sensitive to the expected fluxes and the instrument sensitivity. No points from December 2008 are included due to the reduced sensitivity problem. The linear relationship \( \tau_z(870 \mu m) = 0.32 \text{PWV} + 0.06 \) is used throughout this work.

### 5.3.3 Observations of Flux Calibrators

Observations of flux calibrators are used to monitor the instrumental sensitivity, and also provide a cross-check on the atmospheric opacity determined either from sky-dips or PWV measurements. The sensitivity of LABOCA, as determined from the flux calibrator observations, is generally consistent to within 5%. Thus deviations of much more than 5% indicate that either the assumed opacity is erroneous or that there is a problem with instrument (as shown in §5.5.1).
The primary calibrators are planets, for which the flux at any epoch must be computed. In this work, fluxes for the planets at the epoch of observation were computed using the ASTRO\textsuperscript{11} program within GILDAS\textsuperscript{12}. The frequency and FWHM beam size used were 345 GHz and 18.6″. The beam size is necessary to compute the peak flux which should be measured in a map, or the measured flux if only using a single bolometer pointed at the planet for simple chopped observations.

There is also a collection of \(\sim\) 20 secondary calibrators which are observed regularly with LABOCA. The flux for these has been measured relative to the primary calibrators, and their fluxes are tabulated on the APEX website\textsuperscript{13} and in Siringo et al. [2009]. Note that here the values from the APEX website have been used, which differ slightly in some cases from those in Siringo et al. [2009].

As described above, the reduction of calibration observations within BoA produces maps in AzEl coordinates, with quarter beam pixels (4.5″). The calibrator reduction scripts flag all samples within a 50″ radius of the target before performing despiking and correlated noise removal on the smaller instrument scales to avoid flux being removed from the target. An elliptical Gaussian is fit to the map (not smoothed), and the peak value of the Gaussian is reported as the source flux. Whilst this does not give the total flux of resolved sources, this is the same procedure used to produce the list of calibrator fluxes (Siringo et al. [2009] also give total fluxes as well as the peak fluxes, although these are not used here).

\textsuperscript{12}http://iram.fr/IRAMFR/GILDAS/
\textsuperscript{13}http://www.apex-telescope.org/bolometer/laboca/calibration/
5.4 Photometric Analysis

5.4.1 Measurement Types

There are three general methods for performing photometric measurements on images – aperture photometry, PSF (beam) fitting, and matched filtering. Aperture photometry estimates the total flux of an object by summing pixel values within some aperture centred on an object, subtracting off an estimated background level, and scaling by an aperture correction factor to compensate for the finite aperture size and any contribution of the source flux to the background estimate. PSF fitting performs a $\chi^2$ minimisation of a scaled model of the PSF by varying the amplitude, and often the position and background level. The background level is often determined separately, as for aperture photometry, to avoid degeneracies in the $\chi^2$ fitting. Matched filtering cross-correlates the image with the PSF, and then the value at the position of an object, or the peak nearest to an object, is a measurement of the object’s flux. Again the background must be estimated and subtracted. It can be shown that PSF fitting and matched filtering are equivalent if the sky contribution and noise for each pixel is constant [Naylor, 1998, eqn. 25 with $V_{i,j}$ and $S_{i,j}$ constant].

For aperture photometry it can be shown [see Naylor, 1998] that the aperture radius which maximises $S/N$ for a Gaussian beam is $\frac{\ln 2}{1.6}$FWHM. The aperture correction assuming a Gaussian beam and negligible source contribution to the background estimate, for an aperture radius of $n \times$ FWHM is,

$$f_{ap} = \left( \frac{1.0}{\text{erf}(\frac{2}{\sqrt{2\text{ln} n}})} \right)^2 \quad (5.4)$$

5.4.2 Measurements in this Work

The APEX/LABOCA beam is well approximated by a Gaussian, with slight variability and ellipticity (both $\leq 5\%$). The photometric procedures provided within BoA fit a Gaussian to maps, using the corresponding weight map to determine the uncertainty of each pixel. Typically the peak of the fitted Gaussian, minus a constant background, is taken as the photometric measurement. When the $S/N$ is sufficiently high, an elliptical Gaussian with variable FWHMs and orientation can be fitted (the default for flux calibration observations). The position can be fitted or fixed. As this is the method by which the LABOCA data are calibrated, this procedure is used in this work where feasible. For very faint sources and non-detections, the position has been fixed and a circular Gaussian of FWHM $= 18.6''$ has been used. Some minor modifications within BoA were required to make this possible.
5.4.3 Measuring Noise (RMS) in Maps

There are many methods by which noise can be determined. The relevant quantity for photometry of point sources is the root-mean-square (RMS) noise per beam. The method which has been used here to estimate the noise, is to perform the same measurement as for a faint / non-detected source, at several positions around each target. Generally maps produced with quarter beam pixels (4.5") have been used for this. A rectangular grid of 5 × 5 positions centred on each target has been used, and the sample standard deviation of the 24 off-source points (the source flux is excluded) is taken as the RMS noise per beam. A spacing of 2 × FWHM = 37.2" has generally been used. This spacing means that measurements at adjacent points are not strongly correlated on the scale of the beam. The weight can vary significantly over the largest scales of the grid (diagonal of 3.5"), especially if a target is near the edge of a map, so the noise computed in this manner can be pessimistic as it is dominated by the variations in the noisiest regions.

To check the validity of using the Gaussian fitting routines in BoA for this, measurements at positions on the same grid have been performed by two other methods: Pixel values from 1"/px maps smoothed by a beam size Gaussian, and by averaging over apertures in unsmoothed maps. For the aperture case, an aperture radius of 4FWHM was used, with a corresponding aperture correction computed from Eqn. (5.4). The sample standard deviations of values determined by these two alternative methods were found to agree to within 10%. They were, however, generally only about 70–80% of the values determined by the beam fitting within BoA. This discrepancy is attributed to the additional uncertainty inherent in the background determination, which is included in the beam fit results, but not in the other two methods.

5.4.4 Extended Sources

Photometry of extended sources cannot simply be determined by fitting the peak flux in a map, but must generally be determined by measuring the total flux over a large area. To convert from integrated fluxes measured in maps with units of flux per beam to physical fluxes requires the integral of a beam with unity peak flux to be known. For a Gaussian this is simply $2\pi \sigma^2$ or $\pi \text{FWHM}^2/(4 \ln 2)$. Thus the flux of an object can be determined from the integral over the object in the map divided by $\pi \text{FWHM}^2/(4 \ln 2)$, with the FWHM in units of pixels. The measurement of the total flux can either be achieved by fitting a model of the source or performing aperture photometry with a large aperture. With a Gaussian beam like that of LABOCA, sources which are only slightly extended can often be well fitted by an elliptical Gaussian.
5.5 LABOCA Observations of Southern GASPS Targets

As a precursor to the GASPS Herschel survey, a programme was undertaken to obtain millimeter and sub-millimeter continuum flux measurements for targets lacking these data. Whilst the primary interest of the GASPS survey is in the gas components of circumstellar discs, continuum observations are necessary to constrain the dust masses and dust grain properties of the discs. Even though the dust in protoplanetary discs contributes only $\sim 1\%$ to the total mass, it plays a very important role in the radiative transfer within these discs [e.g., Pinte et al., 2006]. Models of discs designed to allow physical interpretation of gas observations [e.g. Woitke et al., 2009, and references therein] contain continuum radiative transfer due to dust as a key component.

In this programme Northern targets were observed at 1.3 mm with the MAX-Planck Millimeter Bolometer array [MAMBO-2, Kreysa et al., 1998] bolometer array on the Institut de Radioastronomie Millimétrique (IRAM) 30 m telescope on Pico Veleta, Spain. Southern targets were observed at 870 $\mu$m with LABOCA on APEX. The LABOCA observations are the subject of this chapter. The LABOCA targets have ages of 5–30 Myr, and span a wide range of stellar mass and evolutionary stage. Included are the 18 confirmed members of the $\sim 6$ Myr old $\eta$ Chamaeleontis cluster, and a selection of Herbig Ae/Be and T-Tauri stars from a variety of other nearby young stellar associations.

These LABOCA observations are summarised in Table 5.1. They were carried out in two runs, comprising October 19–22 and December 26–28 2008. Publicly available archive data from August 2008 has also been used. The proposed observing strategy was to reach a uniform point-source flux sensitivity of $3\sigma = 20$ mJy, which corresponds to a total disc mass (assuming $\kappa_\nu(870 \mu$m) = 0.35 m$^2$/kg, $T_{\text{dust}} = 20$ K and gas to dust ratio of 100) of $\sim 0.001 M_{\odot}$ at the median target distance of $\sim 100$ pc. This is the same mass limit as achieved in previous surveys of Northern targets e.g. Andrews and Williams [2005].

5.5.1 Instrument Problem During December Run

During the December run there appears to have been a problem with the instrument, which has caused the sensitivity to be reduced to approximately 63% of nominal, equivalent to reducing observing time by a factor of 2.5. This problem was identified by measuring fluxes for calibrators observed on the same nights as the science observations. The calibrator measurements for both runs are shown in Table 5.2. There appears to be no significant trend with elevation or PWV, prompting the conclusion that the sensitivity is generally impaired. Scans taken in the December run have been multiplied by average calibration factors of 1.654 and 1.554 for the nights of the 26th and 28th respectively during data reduction to compensate for this (there
Table 5.1: Summary of LABOCA observations of GASPS targets. The durations are approximate, determined by multiplying the number of usable scans by typical scan durations. The overhead factor for LABOCA observations is 2.0, making the total telescope time used here approximately 46 hours. The observations performed in August are from another programme.

<table>
<thead>
<tr>
<th>Target</th>
<th>Assoc.</th>
<th>Pointing position J2000</th>
<th>Scan pattern</th>
<th>Aug. hours</th>
<th>Oct. hours</th>
<th>Dec. hours</th>
</tr>
</thead>
<tbody>
<tr>
<td>GSC 00056+00482</td>
<td></td>
<td>02 36 51.54 -52 03 04.4</td>
<td>compact</td>
<td>0</td>
<td>1.8</td>
<td>0</td>
</tr>
<tr>
<td>HD 16978</td>
<td></td>
<td>02 39 35.49 -68 16 01.0</td>
<td>compact</td>
<td>0</td>
<td>0.3</td>
<td>0.8</td>
</tr>
<tr>
<td>HD 44627</td>
<td>Carina</td>
<td>06 19 12.91 -58 03 15.5</td>
<td>compact</td>
<td>0</td>
<td>0.9</td>
<td>0</td>
</tr>
<tr>
<td>HD 5081</td>
<td>β Pic</td>
<td>06 18 28.19 -72 02 40.8</td>
<td>compact</td>
<td>0</td>
<td>0.1</td>
<td>1.1</td>
</tr>
<tr>
<td>HD 53842</td>
<td>Tuc-Hor</td>
<td>06 46 13.64 -83 59 29.0</td>
<td>compact</td>
<td>0</td>
<td>0.9</td>
<td>0</td>
</tr>
<tr>
<td>HD 55270</td>
<td>Carina</td>
<td>07 00 30.49 -79 41 45.4</td>
<td>compact</td>
<td>0</td>
<td>0.9</td>
<td>0</td>
</tr>
<tr>
<td>RECX 18</td>
<td>η Cha</td>
<td>08 36 10.73 -79 08 18.4</td>
<td>compact</td>
<td>0</td>
<td>0.9</td>
<td>0</td>
</tr>
<tr>
<td>RECX 1</td>
<td>η Cha</td>
<td>08 36 56.23 -78 56 45.7</td>
<td>compact</td>
<td>0</td>
<td>0</td>
<td>1.0</td>
</tr>
<tr>
<td>RECX 17</td>
<td>η Cha</td>
<td>08 38 51.50 -79 16 13.7</td>
<td>compact</td>
<td>0</td>
<td>0.9</td>
<td>0</td>
</tr>
<tr>
<td>RECX 1-15,17,18</td>
<td>η Cha</td>
<td>08 42 43.20 -78 59 24.0</td>
<td>4 x 4 + 3 x 3</td>
<td>2.4</td>
<td>0</td>
<td>0</td>
</tr>
<tr>
<td>RECX 2-9,13-15</td>
<td>η Cha</td>
<td>08 42 43.20 -78 59 24.0</td>
<td>3 x 3 + 2 x 2</td>
<td>0</td>
<td>0</td>
<td>3.3</td>
</tr>
<tr>
<td>RECX 16</td>
<td>η Cha</td>
<td>08 44 09.15 -78 33 09.2</td>
<td>compact</td>
<td>0</td>
<td>0.9</td>
<td>0</td>
</tr>
<tr>
<td>RECX 10</td>
<td>η Cha</td>
<td>08 44 31.88 -78 46 31.2</td>
<td>compact</td>
<td>0</td>
<td>0.9</td>
<td>0</td>
</tr>
<tr>
<td>RECX 11,12</td>
<td>η Cha</td>
<td>08 47 29.20 -78 57 13.9</td>
<td>compact</td>
<td>0</td>
<td>0.9</td>
<td>0</td>
</tr>
<tr>
<td>TWA 21</td>
<td>TW Hya</td>
<td>10 13 14.72 -52 30 53.8</td>
<td>compact</td>
<td>0</td>
<td>0</td>
<td>0.9</td>
</tr>
<tr>
<td>HD 97048</td>
<td>Cha I</td>
<td>11 08 03.26 -77 39 17.5</td>
<td>compact</td>
<td>0</td>
<td>0</td>
<td>0.7</td>
</tr>
<tr>
<td>HD 100543</td>
<td>LCC</td>
<td>11 33 05.54 -54 19 28.6</td>
<td>compact</td>
<td>0</td>
<td>0</td>
<td>0.9</td>
</tr>
<tr>
<td>HD 100546</td>
<td>LCC</td>
<td>11 33 25.37 -70 11 41.2</td>
<td>compact</td>
<td>0</td>
<td>1.0</td>
<td>0</td>
</tr>
<tr>
<td>HD 104237</td>
<td>ε Cha</td>
<td>12 00 04.97 -78 11 34.6</td>
<td>compact</td>
<td>0</td>
<td>0</td>
<td>0.9</td>
</tr>
</tbody>
</table>

Total hours: 2.4, 8.4, 12.3

are no observations from the 27th). The noise of the scans is seen to be elevated by the same factor, indicating that the cause is not just a simple change in back-end gain.

Upon further investigation it was found that the DC offsets, which are subtracted from the bolometer signals in the warm amplifier units to avoid ADC saturation, are all significantly lower during the December run than at other times. These offsets are related to the bolometer AC bias amplitude, so it is likely that the bias voltages were lower than normal in December. Looking at scans taken over a period of weeks prior to the December observations used here, shows that the voltages were steadily declining as shown in Fig. 5.3, with channels within boxes showing similar values to each other, but values differing significantly between boxes. The times at which the DC offsets start to drop significantly differ by as much as 10 days between boxes of channels, so the effect is unlikely to be due to thermal issues or the incident radiation, which would affect all channels at the same time. The effect could potentially be explained by discharging of the bias power supply batteries in the amplifier boxes. Each box contains an AA
Table 5.2: Flux calibrator measurements taken on nights of GASPS target observations showing the poor sensitivity during the December run. There is no trend with elevation or PWV, indicating that this is an instrument issue. Correction factors of 1.654 and 1.554 have been used for 26th and 28th December respectively.

<table>
<thead>
<tr>
<th>Scan</th>
<th>Object</th>
<th>PWV (mm)</th>
<th>( \tau_e ) (PWV)</th>
<th>Elevation (°)</th>
<th>( F_{\text{expected}} ) (Jy)</th>
<th>( F_{\text{measured}} ) (Jy)</th>
<th>Correction</th>
</tr>
</thead>
<tbody>
<tr>
<td>2008-12-19T00:07:38</td>
<td>G34.3</td>
<td>1.04</td>
<td>0.3928</td>
<td>45.4</td>
<td>56.1</td>
<td>56.68</td>
<td>0.990</td>
</tr>
<tr>
<td>2008-12-19T00:08:38</td>
<td>G45.1</td>
<td>1.04</td>
<td>0.3928</td>
<td>42.8</td>
<td>8.8</td>
<td>8.23</td>
<td>0.984</td>
</tr>
<tr>
<td>2008-12-19T00:10:47</td>
<td>Ursinus</td>
<td>1.03</td>
<td>0.3906</td>
<td>57.6</td>
<td>66.05</td>
<td>65.46</td>
<td>1.009</td>
</tr>
<tr>
<td>2008-12-19T01:24:16</td>
<td>G34.3</td>
<td>0.58</td>
<td>0.2436</td>
<td>29.0</td>
<td>56.1</td>
<td>56.42</td>
<td>0.994</td>
</tr>
<tr>
<td>2008-12-19T03:18:31</td>
<td>V880Lori</td>
<td>0.47</td>
<td>0.2104</td>
<td>23.5</td>
<td>1.31</td>
<td>1.34</td>
<td>0.978</td>
</tr>
<tr>
<td>2008-12-20T06:57:35</td>
<td>N2071R</td>
<td>0.33</td>
<td>0.1656</td>
<td>58.8</td>
<td>9.1</td>
<td>9.25</td>
<td>0.983</td>
</tr>
<tr>
<td>2008-12-20T08:12:31</td>
<td>N2071R</td>
<td>0.33</td>
<td>0.1656</td>
<td>66.5</td>
<td>9.1</td>
<td>9.38</td>
<td>0.970</td>
</tr>
<tr>
<td>2008-12-20T09:30:12</td>
<td>N2071R</td>
<td>0.29</td>
<td>0.1526</td>
<td>61.4</td>
<td>9.1</td>
<td>9.51</td>
<td>0.957</td>
</tr>
<tr>
<td>2008-12-22T01:56:11</td>
<td>Ursinus</td>
<td>0.36</td>
<td>0.1752</td>
<td>71.9</td>
<td>65.83</td>
<td>65.89</td>
<td>0.999</td>
</tr>
<tr>
<td>2008-12-22T01:33:36</td>
<td>CRL618</td>
<td>0.66</td>
<td>0.2712</td>
<td>27.8</td>
<td>4.4</td>
<td>2.70</td>
<td>1.633</td>
</tr>
<tr>
<td>2008-12-22T01:34:45</td>
<td>N2071R</td>
<td>0.50</td>
<td>0.2200</td>
<td>47.8</td>
<td>9.1</td>
<td>5.24</td>
<td>1.738</td>
</tr>
<tr>
<td>2008-12-22T02:59:35</td>
<td>VYCma</td>
<td>0.40</td>
<td>0.1880</td>
<td>50.1</td>
<td>1.5</td>
<td>0.95</td>
<td>1.577</td>
</tr>
<tr>
<td>2008-12-22T05:29:44</td>
<td>CW-LEO</td>
<td>0.26</td>
<td>0.1432</td>
<td>38.4</td>
<td>4.2</td>
<td>2.46</td>
<td>1.707</td>
</tr>
<tr>
<td>2008-12-22T07:54:52</td>
<td>B13134</td>
<td>0.38</td>
<td>0.1816</td>
<td>36.7</td>
<td>12.7</td>
<td>8.18</td>
<td>1.553</td>
</tr>
<tr>
<td>2008-12-22T09:51:09</td>
<td>B13134</td>
<td>0.48</td>
<td>0.2136</td>
<td>46.8</td>
<td>12.7</td>
<td>7.41</td>
<td>1.714</td>
</tr>
<tr>
<td>2008-12-22T03:48:34</td>
<td>N2071R</td>
<td>0.98</td>
<td>0.3736</td>
<td>66.6</td>
<td>9.1</td>
<td>5.60</td>
<td>1.624</td>
</tr>
<tr>
<td>2008-12-22T06:28:47</td>
<td>CW-LEO</td>
<td>0.75</td>
<td>0.3000</td>
<td>48.5</td>
<td>4.2</td>
<td>2.77</td>
<td>1.515</td>
</tr>
<tr>
<td>2008-12-22T07:12:18</td>
<td>B13134</td>
<td>0.40</td>
<td>0.1880</td>
<td>36.2</td>
<td>12.7</td>
<td>8.36</td>
<td>1.520</td>
</tr>
<tr>
<td>2008-12-22T08:52:36</td>
<td>VYCMa</td>
<td>0.47</td>
<td>0.2104</td>
<td>43.2</td>
<td>1.5</td>
<td>0.96</td>
<td>1.556</td>
</tr>
</tbody>
</table>

battery to provide clean power for biasing the bolometers which the box is connected to.

Thankfully, the relative values of the DC offsets between boxes are approximately nominal on the 26th and 28th, even though they are all reduced to approximately half their nominal values. This is significant for targets which are not centred on the array, and hence fall on different bolometers to flux calibrators. The relative array flat-field which is used during the data reduction for the December data\(^\text{14}\) was computed from beam-map observations performed on November 19th and December 25th, and thus should be valid on the 26th and 28th, although based on Fig. 5.3, it is unlikely to be valid for the nights of 13–22 December. In the channel flags file\(^\text{15}\) it is noted that there are 16 extra bolometers seem to be bad in the December 25th beam-map, which is potentially related to the reduced DC offsets. These bolometers are flagged as bad prior to reducing the December scans in this work.


Figure 5.3: DC offsets subtracted before A/D conversion, averaged over boxes of channels during December 2008. The points are the nightly median of the box averaged values from observations. Unused channels, with DC offset values below 1.0, are not included in the averages. There were no observations performed after the 28th until April 2009.
5.5.2 η Chamaeleontis

The η Chamaeleontis association is an open cluster at a distance of 94 pc, approximately 1 pc in extent, with 18 known members and an estimated age of 4–9 Myr (see below). This cluster is of particular significance due to the combination of its youth and relative proximity, as shown by a comparison with other clusters within 5 kpc (Fig. 5.4). The age of the cluster is comparable to the lifetime of protoplanetary discs, so the members can give important insight into the final stages of protoplanetary disc evolution and the transition to lower mass gas-poor debris discs.

![Figure 5.4: Distance and age of the η Cha cluster in relation to other open clusters, showing a rare combination of proximity and youth. Figure from Mamajek et al. [2000].](image)

**History & Cluster Membership**

The η Cha cluster was discovered by ROSAT X-ray observations which showed a clustering of T-Tauri stars [Mamajek et al., 1999]. Initially 12 X-ray detected members, referred to as *ROSAT* η Cha X-ray (RECX) 1–12 [Mamajek et al., 1999, 2000], and a co-moving X-ray quiet member, HD 75505, were reported. A further 5 members have since been identified [Lawson et al., 2002, Lyo et al., 2004, Song et al., 2004]. Here, as in other works [e.g. Luhman and Steeghs, 2004], the additional members are assigned membership numbers 13–18 starting with HD 75505 as 13. The RECX identifier is used with these for convenience (and is recognised by SIMBAD), even though they are not X-ray detected. Further surveys, either covering larger areas [Luhman, 2004], or with the ability to detect lower mass stars [Lyo et al., 2006] have not found any other members.

There are three intermediate mass members with B/A spectral types, at least one of which (RS Cha, RECX 8) is a binary. The other members all have spectral types of late K to M6. The lack of F/G-type stars is to be expected at the age of the cluster due to the temperature
5.5. LABOCA OBSERVATIONS OF SOUTHERN GASPS TARGETS

evolution of pre-main-sequence stars. At the age of this cluster the higher mass (\(\gtrsim 1.5 \, M_\odot\))
stars are already on the main-sequence, however, the lower mass stars will continue to contract
and increase in temperature. For example, RECX 1, 7 and 11, with spectral types \(\sim K6\),
have masses of approximately \(1.0 \, M_\odot\) [Lyo et al., 2004]. The lack of stars later than M6 type,
corresponding to around \(0.15 \, M_\odot\), has been considered abnormal compared to other clusters
[Luhman, 2004, Lyo et al., 2004, Murphy et al., 2010]. The implication of this is that either the
initial mass function of the cluster is abnormal, or that lower mass stars and brown dwarfs have
been dynamically expelled to large distances. This latter scenario has been shown to be possible
by modelling of the dynamical evolution of the cluster [Moraux et al., 2007]. Recently Murphy
et al. [2010] have discovered four probable low mass (0.08–0.3 \(M_\odot\)) members with separations up
to \(5\degree\) (8 pc projected) from the cluster centre, which suggest that ejection has indeed occurred.

**Wider Associations**

Shortly after its discovery, Mamajek et al. [2000] suggested a relationship between the \(\eta\) Cha
cluster and the \(\epsilon\) Cha association, as well as potentially the TW Hya association. This was based
on tracing back the trajectories of stars in these associations, and finding that they converge
within a few tens of parsecs around 10 Myr ago. At that time they would fall near what is
now the Lower-Centaurus-Crux (LCC) region of the large Scorpius-Centaurus OB association.
Jilinski et al. [2005] performed a more detailed study, retracing the dynamical centres of the \(\eta\)
Cha cluster and the \(\epsilon\) Cha association, and find that they converge to within 3 pc at 6–7 Myr
ago. This is consistent with the ages of both associations, so it is highly likely that the stars in
both clusters were formed together and have since diverged. Jilinski et al. [2005] also determine
distances from the LCC association of \(\sim 25\) pc at the corresponding epoch, suggesting that \(\eta\)
Cha and \(\epsilon\) Cha were formed on the outskirts of the LCC rather than inside it.

**Distance and Age**

The distance to the \(\eta\) Cha cluster can be determined from Hipparcos parallaxes. There are
two cluster members with Hipparcos measurements, \(\eta\) Cha (RECX 2) and HD 75747 (RECX
8), with a variance weighted mean parallax of \(10.61 \pm 0.13\) mas [F. van Leeuwen, 2007]. This
corresponds to a distance of \(94.2 \pm 1.2\) pc.

Estimates of the cluster age range from approximately 4 to 9 Myr based on photometry and
spectroscopy of the stars [Mamajek et al., 1999, Lawson et al., 2001, Luhman and Steeghs,
2004]. As noted above, Jilinski et al. [2005] have shown that the \(\eta\) Cha cluster and the \(\epsilon\) Cha
association were at a common point 6–7 Myr ago, so it is tempting to take this as the age of
both groups. Treating the \(\eta\) Cha cluster as part of the \(\epsilon\) Cha association, Torres et al. [2008]
and da Silva et al. [2000] adopt a common age of 6 Myr.

Disc Population

Based on classification of optical spectra, the three early type members (η Cha, RS Cha AB and HD 75505) are main-sequence stars, and the majority of the lower mass members are weak-lined or classical T-Tauri stars [Lawson et al., 2002, Sicilia-Aguilar et al., 2009]. The SEDs of members from optical to mid-IR wavelengths have been studied by several authors [Lyo et al., 2003, Megeath et al., 2005, Gautier et al., 2008, Sicilia-Aguilar et al., 2009]. The three early type members show no significant near/mid-IR excess indicative of protoplanetary discs. The late type members, however, feature a wide variety of SED types. In total eight of the fifteen late type members show near/mid-IR excess indicating the presence of protoplanetary discs. Of these, two have typical Lada class II SED profiles corresponding to flared discs, and another two appear to have flattened discs which also extend in close to the central star. The remaining four show only mid-IR excess, indicating a lack of material in the inner region of their discs. Given that they show signs of ongoing accretion by their Hα emission, these objects are classified as transitional objects (TOs). TOs are generally considered quite rare, typically only making up around 10% of discs in most regions of young stars [Megeath et al., 2005]. The rarity of TOs in regions such as Taurus had led to the thinking that TOs are a brief evolutionary phase between class II discs and cleared class III systems, however, the large fraction in η Cha questions this hypothesis.

The longest wavelength observations of the η Cha cluster published to date are from Sicilia-Aguilar et al. [2009], which includes Spitzer MIPS 24 and 70 μm photometry. The star η Cha is seen to have significant excess at these wavelengths, which given the lack of accretion signatures in optical spectra or spectral features from small dust grains in Spitzer IRS spectra, is attributed to a debris disc. Four of the discs around low mass members are detected at 70 μm. Prior to Herschel observations there have been no published observations at wavelengths longer than 70 μm, so very little is know about the properties of the outer regions or total masses of the discs in this cluster. The GASP survey with Herschel is adding deeper 70 μm photometry, and photometry at 160 μm using the PACS instrument. Follow-up with the SPIRE instrument will provide photometry at 250, 350 and 500 μm, however, the sensitivity at the longer wavelengths is poor due to extragalactic confusion (3σ ~ 20 mJy in all three bands [Nguyen et al., 2010]). Thus there is strong motivation for ground-based sub-millimetre photometry to provide constraints on the outer disc properties and dust masses of the discs.

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16The photometry in Sicilia-Aguilar et al. [2009] supersedes previously published values in Gautier et al. [2008].
5.5. LABOCA OBSERVATIONS OF SOUTHERN GASPS TARGETS

LABOCA Observations

The LABOCA data used here comprise three variants of raster spiral maps:

- $4 \times 4 + 3 \times 3$ raster spirals (2.4 h total), evenly covering all but RECX 16, taken by another project on 15 August 2008
- $3 \times 3 + 2 \times 2$ raster spirals (3.3 h total), evenly covering the central $15' \times 15'$ of the cluster, taken on 26/28 December 2008
- six compact raster spirals centred on the outer members, taken on 21 October (RECX 11+12 and 16) and 26 December 2008 (RECX 1, 10, 17, 18)

All of these scans were reduced using the improved weak-source pipeline with final $\sigma$-clipping thresholds of $\pm 3\sigma$. Maps with $\frac{1}{4}$-beam (4.5") pixels were produced from the scans and stacked. The resulting weight map is shown with the member positions overlaid in Fig. 5.6. It is especially unfortunate that the $3 \times 3 + 2 \times 2$ raster spiral observations, which provide the deepest coverage of the central 11 cluster members, were performed in the December run when the instrument sensitivity was impaired.

The map RMS measured around each of the targets ranges from 6 to 12 mJy/beam. None of the cluster members are detected at greater than 3$\sigma$. The fluxes measured for an effective beam at the position of each target, along with the measured RMS, are given in Table 5.3. RECX 17 is marginally detected at $2.9\sigma$ (18.4 ± 6.3 mJy). A map centred on RECX 17, smoothed with the beam, is shown in Fig 5.5. Although there are no significant detections, these observations place the first upper limits on the dust masses of the discs in the cluster.

Masses computed from the 3$\sigma$ upper limits, assuming $T_{\text{dust}} = 30$ K and $\kappa_\nu(870 \mu m) = 0.17 \text{ kg/m}^2$ are also given in Table 5.3. These masses can be approximately scaled to other assumed values of $T_{\text{dust}}$ and $\kappa_\nu(870 \mu m)$ by $M_{\text{dust}} \propto \kappa_\nu(870 \mu m)T_{\text{dust}}$.

If the marginal detection of RECX 17 is real, then given the lack of excess at wavelengths of 8 $\mu m$ and less by Sicilia-Aguilar et al. [2009], this is potentially another transition object, with a substantial outer disc. There is no published MIPS photometry of RECX 17 with which to compare the LABOCA photometry. However, MIPS observations of RECX 17 have been made (AOR: 25961084, PROGID: 50316, PI: Rieke). An inspection of the 70 $\mu m$ post-BCD image of RECX 17 shows that there is no detection. This casts doubt on the reality of the LABOCA detection, although it cannot rule out a significant mass of dust at a temperature lower than $\sim 30$ K.
**Figure 5.5:** Weight map of η Chamaeleontis with the 18 stars over-plotted

**Figure 5.6:** Map centred on RECX 17, smoothed with the beam (effective FWHM = 26.3′′). Left: original map. Right: 18 mJy beam centred at (0, 0). Right: map with 18 mJy beam subtracted.
Stacking Analysis

In spite of obtaining no significant detections of cluster members, it is possible to constrain the mean disc flux, and hence mass, by stacking (averaging) maps centred on all 18 cluster members (with equal weighting). The resulting stacked map, after smoothing with a beam-sized (FWHM = 18.6′′) Gaussian, is shown in Fig. 5.7 (left). The RMS in the stacked map is 2.3 mJy/beam, and there is a peak near the centre. An effective-beam (26.3′′ FWHM) fit to the central peak gives a flux of 5.8 mJy on a background level of −0.6 mJy, and with a centroid 9′′ from the map centre. Given the low S/N and the likelihood of relatively poor pointing accuracy due to being at a very southerly declination, this centroid offset is not considered significant. As the fitted flux is less than 3σ, a 3σ upper mass limit is also computed.

The mass estimates assume κν(870 μm) = 0.17 kg/m², and the cluster distance of 94 pc. Treating the stacked flux as a detection of 5.8 mJy and assuming a typical dust temperature of 60 K yields a dust mass of 0.26 M_⊙. Instead, taking 3σ = 7.0 mJy as an upper limit on the flux, and 30 K as a lower limit on the dust temperature, yields a mass upper limit of 0.72 M_⊙. Given the lack of knowledge of the characteristic dust temperature and the uncertainty on the opacity, the most reliable conclusion that can be made is that the average dust mass is less than about 1 M_⊙. This average mass upper limit is shown in context with sub-millimetre detected discs in Fig. 5.8.

![Stacked map of all 18 members of η Chamaeleontis (left), effective-beam (26.3′′ FWHM) fit (centre), and map with fit subtracted (right).](image)

**Figure 5.7:** Stacked map of all 18 members of η Chamaeleontis (left), effective-beam (26.3′′ FWHM) fit (centre), and map with fit subtracted (right).

**GASPS First Results: RECX 15**

The classical T Tauri star RECX 15 (ET Cha) has been one of the first objects to be studied by the GASPS project using observations from *Herschel* [Woitke et al. 2010, in prep]. By fitting the SED of RECX 15 (Fig. 5.9), including *Herschel* PACS photometry at 100 and 160 μm, and...
the LABOCA 870\(\mu\)m upper limit presented here, Woitke et al. [2011] determine a dust mass of 0.07\(M_\odot\). This corresponds to an 870\(\mu\)m flux of \(\sim 2\) mJy. However, there is still significant uncertainty in this dust mass due to the lack of detections longward of 160\(\mu\)m.

Figure 5.9: SED fit for RECX 15 from Woitke et al. [2011]. The 3\(\sigma\) LABOCA 870\(\mu\)m upper limit is shown at the right.
Table 5.3: η Cha cluster member properties, 870 µm photometry and mass limits. Spectral types, Hα types, SED types and 70 µm photometry are from Sicilia-Aguilar et al. [2009]. Stellar masses are from Lyo et al. [2004] (inferred from evolutionary grids). The LABOCA 870 µm photometry is listed with 1σ uncertainty. The dust mass upper limits are computed from the 3σ uncertainty assuming a dust temperature of 30 K and \( \kappa_\nu(870 \, \mu m) = 0.17 \, kg/m^2 \).

WTTS = Weak-lined T-Tauri Star, CTTS = Classical T-Tauri Star; class II = near-IR SED flatter than \( \lambda F_\lambda \propto \lambda^{-4/3} \) (flared protoplanetary disc), flat = physically flat disc (\( \lambda F_\lambda \propto \lambda^{-4/3} \)), TO (Transition Object) = no near-IR excess but has mid-IR excess, class III = no near- or mid-IR excess, debris = debris disc, spectrally featureless due to having only large dust grains.

<table>
<thead>
<tr>
<th>RECX</th>
<th>Names</th>
<th>Spectral Type</th>
<th>Stellar mass(es) ((M_\odot))</th>
<th>Hα Type</th>
<th>SED Type</th>
<th>( F_\nu(70 , \mu m) ) (mJy)</th>
<th>( F_\nu(870 , \mu m) ) (mJy)</th>
<th>( M_{\text{dust}}(30 , K) ) (( M_\odot ))</th>
</tr>
</thead>
<tbody>
<tr>
<td>1 AB</td>
<td>EG Cha, CPD -78.367</td>
<td>K7.0 + M0.0</td>
<td>1.00 + 0.70</td>
<td>WTTS</td>
<td>class III</td>
<td>&lt; 20</td>
<td>-1 ± 6</td>
<td>&lt; 1.9</td>
</tr>
<tr>
<td>2</td>
<td>η Cha, HD 75416</td>
<td>B8 + M1?</td>
<td>3.40 + 0.50?</td>
<td>WTTS</td>
<td>class III / debris</td>
<td>22 ± 4</td>
<td>14 ± 10</td>
<td>&lt; 3.1</td>
</tr>
<tr>
<td>3</td>
<td>EH Cha</td>
<td>M3.0</td>
<td>0.32</td>
<td>WTTS</td>
<td>class III</td>
<td>&lt; 20</td>
<td>-1 ± 12</td>
<td>&lt; 3.7</td>
</tr>
<tr>
<td>4</td>
<td>EL Cha</td>
<td>M1.3</td>
<td>0.49</td>
<td>WTTS</td>
<td>TO</td>
<td>&lt; 20</td>
<td>4 ± 10</td>
<td>&lt; 3.1</td>
</tr>
<tr>
<td>5</td>
<td>EK Cha</td>
<td>M3.8</td>
<td>0.26</td>
<td>CTTS</td>
<td>TO</td>
<td>87 ± 9</td>
<td>16 ± 11</td>
<td>&lt; 3.4</td>
</tr>
<tr>
<td>6</td>
<td>EL Cha</td>
<td>M3.0</td>
<td>0.35</td>
<td>WTTS</td>
<td>class III</td>
<td>&lt; 20</td>
<td>-4 ± 12</td>
<td>&lt; 3.7</td>
</tr>
<tr>
<td>7</td>
<td>EM Cha</td>
<td>K6.9 + M1.0</td>
<td>1.08 + 0.50</td>
<td>WTTS</td>
<td>class III</td>
<td>&lt; 20</td>
<td>-6 ± 11</td>
<td>&lt; 3.4</td>
</tr>
<tr>
<td>8</td>
<td>RS Cha, HD 75747</td>
<td>A7 + A8 + M1?</td>
<td>1.86 + 1.82 + 0.50?</td>
<td>WTTS</td>
<td>class III</td>
<td>&lt; 20</td>
<td>-9 ± 11</td>
<td>&lt; 3.4</td>
</tr>
<tr>
<td>9</td>
<td>EN Cha</td>
<td>M4.4 + M4.7</td>
<td>0.20 + 0.18</td>
<td>CTTS</td>
<td>TO</td>
<td>42 ± 5</td>
<td>5 ± 7</td>
<td>&lt; 2.2</td>
</tr>
<tr>
<td>10</td>
<td>EM Cha</td>
<td>M0.3</td>
<td>0.60</td>
<td>WTTS</td>
<td>class III</td>
<td>&lt; 20</td>
<td>5 ± 6</td>
<td>&lt; 1.9</td>
</tr>
<tr>
<td>11</td>
<td>EP Cha, CPD -78.388</td>
<td>K6.5</td>
<td>1.04</td>
<td>CTTS</td>
<td>class II</td>
<td>&lt; 20</td>
<td>3 ± 6</td>
<td>&lt; 1.9</td>
</tr>
<tr>
<td>12</td>
<td>EQ Cha</td>
<td>M3.2 + M2.7?</td>
<td>0.35 + 0.35?</td>
<td>WTTS</td>
<td>class III</td>
<td>148 ± 16</td>
<td>4 ± 7</td>
<td>&lt; 2.2</td>
</tr>
<tr>
<td>13</td>
<td>HD 75905</td>
<td>A1</td>
<td>1.85</td>
<td>CTTS</td>
<td>class III</td>
<td>&lt; 20</td>
<td>-14 ± 9</td>
<td>&lt; 2.8</td>
</tr>
<tr>
<td>14</td>
<td>ES Cha, ECHA J0841.5-7853</td>
<td>M4.7</td>
<td>0.19</td>
<td>WTTS</td>
<td>TO / flat</td>
<td>&lt; 20</td>
<td>7 ± 10</td>
<td>&lt; 3.1</td>
</tr>
<tr>
<td>15</td>
<td>ET Cha, ECHA J0843.3-7005</td>
<td>M3.4</td>
<td>0.37</td>
<td>CTTS</td>
<td>class II</td>
<td>136 ± 14</td>
<td>5 ± 10</td>
<td>&lt; 3.1</td>
</tr>
<tr>
<td>16</td>
<td>ECHA J0841.2-7833</td>
<td>M5.5</td>
<td>0.14?</td>
<td>flat</td>
<td>—</td>
<td>2 ± 6</td>
<td>&lt; 1.9</td>
<td></td>
</tr>
<tr>
<td>17</td>
<td>ECHA J0838.9-7916</td>
<td>M5.0 + M5.0?</td>
<td>0.16 + 0.16?</td>
<td>flat</td>
<td>class III</td>
<td>—</td>
<td>18 ± 6</td>
<td>≤ 1.9</td>
</tr>
<tr>
<td>18</td>
<td>ECHA J0839.2-7908</td>
<td>M5.3 + M5.3?</td>
<td>0.15 + 0.15?</td>
<td>flat</td>
<td>class III</td>
<td>—</td>
<td>10 ± 6</td>
<td>&lt; 1.9</td>
</tr>
</tbody>
</table>

5.5. LABOCA OBSERVATIONS OF SOUTHERN GASPS TARGETS
5.5.3 Herbig Ae/Be Stars

Four Herbig Ae/Be (HAeBe) stars were observed as part of this programme. These are all bright sources, but until now they lacked sub-millimetre photometry due to their southerly declinations making them inaccessible to the JCMT and instruments such as SCUBA. The parameters of the stars are summarised in Table 5.4. They are all very close to, or are already on, the ZAMS.

As these targets are not closely associated with one another and their observations have required different analysis, they are each discussed separately. Their (sub-)millimetre SED properties are then discussed in the context of previous studies of HAeBes.

Table 5.4: Properties of the four Herbig Ae/Be stars in this work. Distances are from F. van Leeuwen [2007], and the other properties are taken from van Boekel et al. [2005], Collins et al. [2009] and Grady et al. [2004].

<table>
<thead>
<tr>
<th>Star</th>
<th>Association</th>
<th>d</th>
<th>SpT</th>
<th>Teff</th>
<th>L</th>
<th>M</th>
<th>Age</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td></td>
<td>(pc)</td>
<td></td>
<td>(K)</td>
<td>(L⊙)</td>
<td>(M⊙)</td>
<td>(Myr)</td>
</tr>
<tr>
<td>HD 97048</td>
<td>Cha I</td>
<td>158 ±16</td>
<td>B0.5 V</td>
<td>10000</td>
<td>44</td>
<td>2.5</td>
<td>&gt; 2</td>
</tr>
<tr>
<td>HD 100433</td>
<td>LCC</td>
<td>121 ±10</td>
<td>A9 Ve + M4 Ve</td>
<td>7400</td>
<td>8</td>
<td>1.7</td>
<td>10</td>
</tr>
<tr>
<td>HD 100516</td>
<td>LCC</td>
<td>97 ±4</td>
<td>B9 Vae</td>
<td>10500</td>
<td>32</td>
<td>2.4</td>
<td>&gt; 10</td>
</tr>
<tr>
<td>HD 104237</td>
<td>ε Cha</td>
<td>115 ±5</td>
<td>A7.5 Ve</td>
<td>7300</td>
<td>25</td>
<td>2.1</td>
<td>5</td>
</tr>
</tbody>
</table>

HD 97048 and the Chamaeleon I Dark Cloud

HD 97048 is located in the Chamaeleon I dark cloud, and the LABOCA observations cover a significant portion of this region. The full extent of the LABOCA map is shown in Fig. 5.10. The median age of members of Cha I is ~2 Myr [Luhman, 2008]. There is a lot of useful information which can be extracted for objects within the field, other than HD 97048, however, this is beyond the scope of this work.

HD 97048 is the brightest source in this programme. The photometry for this source has been produced by reducing each of the observed scans in the same manner as for flux calibrator observations. There are six usable scans, taken consecutively on 26 December 2008, with very stable conditions (PWV = 0.50–0.53 mm, E1 = 35°). A calibration correction factor of 1.654 has been applied to all observations on that night to compensate for the abnormally low sensitivity. The mean flux and sample standard deviation from the six scans are 2610 and 47 mJy respectively. The photometry is thus calibration limited, and a 1σ uncertainty of 5% is assumed. Due to the instrument problems at the time of observation, this uncertainty may be optimistic. The standard deviation of the calibration factors determined from flux calibrator
observations on the night, as given in Table 5.2, is 4.7% of the mean value. Considering that the random uncertainty on these measurements is significant due to short scan durations, a total calibration uncertainty of 5% is reasonable.

The scans for HD 97048 have also been reduced using the weak-source reduction process described above. Final σ-clipping thresholds of −4σ and +9σ were chosen to preserve the source flux. A stacked map with a 1′ pixel scale was produced. This was used to investigate the spatial extent of the source by fitting and subtracting an elliptical Gaussian to the map (Fig. 5.11, top row). The fitted FWHMs are consistent with the source being unresolved, and thus the photometry produced from the calibrator reduction is considered valid (the calibrator reduction determines the peak flux, which is a lower limit for the total flux of resolved sources). As the LABOCA beam is somewhat variable due to pointing and focus drifts, with stated uncertainty of 1.0′′ [Siringo et al., 2009], it is not meaningful to derive deconvolved sizes from FWHMs less than ∼22″, which would correspond to a deconvolved size of ∼11″. Thus the best constraint that can be placed on the size of the emitting region in HD 97048 is that it is smaller than ∼11″ (∼1800 AU).

The emission seen to the North of HD 97048 in the map in Fig. 5.11 is most likely due to two related lower mass T-Tauri stars separated by ∼35″ from HD 97048 [Habart et al., 2003].

**HD 100453**

HD 100453 is a member of the Lower Centaurus-Crux (LCC) complex, which is one of the three regions constituting the Scorpius-Centaurus OB association [de Zeeuw et al., 1999]. HD 100453 is quite isolated, with only one other object listed by SIMBAD within the LABOCA field (a galaxy, LEDA 2793008). HD 100453 is a binary system, with a low mass companion (∼0.2 $M_\odot$) separated by 1.06″ (∼130 AU projected) [Chen et al., 2006, Collins et al., 2009]. This separation is well below the size scales which can be resolved in the LABOCA observations, so all the results here are for the sum of the two components and their discs.

The LABOCA observations of HD 100453 comprise eight scans taken on 28 December 2008 in good, but variable conditions (PWV = 0.39–0.83 mm, El = 44–51°). A flux calibration correction factor of 1.554 was applied to the data for that night. As the S/N in maps produced from the individual scans is low (∼5), the photometry presented here is obtained from a stacked map. The scans were reduced, and maps produced, in the same way as for flux calibration observations. The measured source flux and RMS in the map are 494 mJy and 35 mJy/beam. The RMS is added in quadrature with 5% calibration uncertainty to give a total photometric uncertainty of 43 mJy. Note that using the improved weak-source pipeline, the resulting RMS is considerably lower (∼15 mJy), however, the calibration is more uncertain.
CHAPTER 5. LABOCA DATA REDUCTION & OBSERVATIONS OF SOUTHERN
CIRCUMSTELLAR DISCS

As for HD 97048 above, the spatial extent of HD 100453 was examined by fitting and subtracting an elliptical Gaussian to a stacked map with 1'' pixels (Fig. 5.11, middle row). The map for this was produced using the improved weak-source pipeline, with final $\sigma$-clipping thresholds of $-3\sigma$ and $+5\sigma$ chosen to approximately preserve the peak flux whilst minimizing noise. The fit is consistent with HD 100453 being unresolved, and no significant emission is seen after subtraction of the elliptical Gaussian fit.

HD 100546

HD 100546 is also a member of the LCC complex [de Zeeuw et al., 1999], and is located on the outskirts of a molecular cloud [DC 296.2-7.9, Vieira et al., 1999]. There are a spattering of other LCC members in the LABOCA field, although none are detected or lie close to HD 100546. There are five stars reported within $\sim 10''$ of HD 100546, however, these are all faint, distant background stars, and HD 100546 is considered to be single [Grady et al., 2001]. Optical/near-IR coronographic imaging has shown that the disc of HD 100546 extends to radii of 2-5'', and there is an extended envelope visible to radii of $\sim 10''$ [Grady et al., 2001].

The LABOCA observations for HD 100546 comprise nine scans taken on 20 October 2008 in excellent and stable conditions (PWV = 0.27-0.32 mm, El = 29-34°). No flux calibration correction factor was required for this night. As for HD 97048, it was possible to perform photometry on maps produced from individual scans, reduced as for flux calibrators. The mean flux and sample standard deviation from the nine scans are 1324 and 18 mJy respectively. Thus the photometric uncertainty is completely calibration dominated, and a value of $1\sigma = 5\%$ is assumed, giving a final value of 1324 $\pm 66$ mJy.

The spatial extent was examined in the same manner as for the previous two sources, by fitting an elliptical Gaussian to a stacked map with 1'' pixels, produced using the improved weak-source pipeline with final $\sigma$-clipping thresholds of $-4\sigma$ and $+9\sigma$ (Fig. 5.11, bottom row). The fit is consistent with the emission being unresolved, and there is no significant residual emission after subtracting the fitted Gaussian. This is of significance, as previous 1.3 mm mapping by Henning et al. [1998] showed this source to have possible extended emission to the South-East with a FWHM $\sim 23''$ beam. Henning et al. [1998] do point out that this extended emission is at their detection limit, and given the much higher S/N in the LABOCA map (> 100 vs. 10) it would appear that this was just noise in the 1.3 mm map. The lack of extended emission suggests that there is relatively little mass of dust outside the $\sim 10''$ envelope detected by Grady et al. [2001] in scattered light optical imaging.
5.5. LABOCA OBSERVATIONS OF SOUTHERN GASPS TARGETS

HD 104237 ABCDE and the ε Chamaeleontis association

HD 104237 is the second most massive star in the ε Chamaeleontis association. The ε Cha association, originally proposed in Feigelson et al. [2003], is thought to have been formed in close proximity to and at the same time as the η Cha cluster which was studied above [Jilinski et al., 2005]. The age of these associations is approximately 6 Myr. Age estimates for HD 104237 have been as low as 2 Myr [van den Ancker et al., 1997, Feigelson et al., 2003], however, Grady et al. [2004] show that the spectral types that were used in the previous works were too early, and they derive an age of 5 Myr using a revised spectral type of A7.5.

HD 104237 is the most massive star in a multiple system of at least 5 stars [Feigelson et al., 2003]. The four other components, B, C, D and E, are K–M type T-Tau type stars [Feigelson et al., 2003, Grady et al., 2004]. There is a close component at a separation of 1.37″, which is potentially responsible for truncation of the disc of HD 104237 [Grady et al., 2004]17. The C component is separated by 5.3″, and the D and E components are separated by 10.7″ and 14.9″ respectively. The D and E components are thus marginally resolved in the LABOCA beam. The E component is considered to be a classical T-Tauri star by Feigelson et al. [2003], and Grady et al. [2004] detect significant excess at 11.9 μm. The D component is classified as a weak-lined T-Tauri star and lacks a detected excess in the observations of Grady et al. [2004].

The LABOCA observations of HD 104237 comprise eight scans taken on 28 December 2008 in good, stable conditions (PWV = 0.42–0.52 mm, El = 33–34°). A flux calibration correction factor of 1.554 was applied to the data for that night. This source is fainter than the previous ones, and indeed the photometric uncertainty is not calibration dominated. To maximise S/N, all analysis has been performed on stacked maps produced from scans reduced using the improved weak-source pipeline, with final σ-clipping thresholds of −3σ and +5σ. The central region of the map is shown in Fig. 5.12. The RMS in the map is 18 mJy/beam. Two other members of the ε Cha association listed in Feigelson et al. [2003] fall within the LABOCA field (ε Cha and CXOU J115008.2-781232), but these are not detected.

It is clear from the map that the emission toward HD 104237 is extended. Based on the mid-IR imaging in Grady et al. [2004], it was hypothesised that the disc of the classical T-Tauri E component was most likely to be responsible for the extra emission. A model of two beam-sized Gaussians with a separation vector equal to the A-E separation was fitted to the map. The fitted parameters were the amplitude of the two Gaussians, a position offset, and the background level. Due to the relatively low S/N, the fit was performed to both an unsmoothed map using FWHM = 18.6″, and to a map smoothed with a half-beam Gaussian using FWHM = 20.8″. The results show little difference, and are summarised in the table in Fig. 5.12. There is no

17 It is not clear whether the 'B' component in Feigelson et al. [2003] is the same as the star designated HD 104237-2 in Grady et al. [2004], although both have separations ≈1″.
significant residual emission after subtracting the fit, and the position offset is well within the typical pointing uncertainty. The total flux from the fit to the unsmoothed map is 245 mJy. Given the uncertainties in the size of the beam, and the fact that the data reduction for this map was different to that of flux calibrators, a calibration uncertainty of \(\sigma = 10\%\) is estimated. When added in quadrature with the map RMS of 18 mJy/beam this results in a photometric measurement for the whole system of 245 ± 30 mJy. The fluxes for the A and E components are subject to a larger uncertainty due to their values being anti-correlated. Assuming the uncertainty of each to be 10\% of the flux of the A component, added in quadrature with 18 mJy, gives 154 ± 24 mJy and 91 ± 24 mJy for the A and E components respectively. If there is any emission from the D component, then this will mostly be included in the photometry of the E component, and likewise any emission from the B and C components will be incorporated in the photometry of the A component.

It is interesting to note that the flux associated with the E component is similar to the flux of the disc around the primary star, despite the widely different spectral types [A7.5 vs. K3; Grady et al., 2004]. The spectral types, however, are a little misleading (see §1.3.1), and the ratio of the fluxes is essentially the same as the ratio of the stellar masses [2.1 vs. 1.3 \(M_\odot\); Grady et al., 2004].

To allow comparison with with millimetre photometry in the literature in the following section, the map has been smoothed to an effective beam size of FWHM = 23″ by smoothing with a Gaussian of FWHM = 13.5″. The value at the position of HD 104237 A after subtracting an average background value is 190 mJy (this is within 1 mJy of the peak value in this map). Assuming a calibration uncertainty of 10\% and the map RMS of 18 mJy/beam makes this value 190 ± 26 mJy.
Figure 5.10: Full LABOCA map of HD 97048 in relation to Chamaeleon I. Top-left: plot of all members of Cha I with contours of extinction, $A_V$, taken from Luhman et al. [2008]. Top-right: optical colour-composite image of Cha I ($1.4 \times 2$') taken from Luhman [2008]. The region of the LABOCA map is highlighted as a green circle. Bottom: full LABOCA map centred on HD 97048, smoothed with half-beam Gaussian. The other very bright object, seen to the lower left, is the Chamaeleon IR nebula. The LABOCA map shown here was reduced without frequency flattening to preserve large scale structure, and with final $\sigma$-clipping thresholds of $\pm 3\sigma$ to minimise noise in the background.
**Figure 5.11**: LABOCA maps of the three bright Herbig Ae/Be stars HD 97048 (top), HD 100453 (middle), and HD 100546 (bottom). The maps are unsmoothed and have 1″ pixels. Left: maps; centre: elliptical Gaussian fit; right: map with fit subtracted. The table gives the FWHMs, position offsets and total flux of the elliptical Gaussian fits.

<table>
<thead>
<tr>
<th>Target</th>
<th>FWHM/″</th>
<th>( \Delta x_{\text{peak}}/″ )</th>
<th>( \Delta y_{\text{peak}}/″ )</th>
<th>( F_v/\text{mJy} )</th>
</tr>
</thead>
<tbody>
<tr>
<td>HD 97048</td>
<td>19.7 × 18.5</td>
<td>−0.9</td>
<td>2.8</td>
<td>2655</td>
</tr>
<tr>
<td>HD 100453</td>
<td>20.4 × 18.3</td>
<td>3.0</td>
<td>3.0</td>
<td>477</td>
</tr>
<tr>
<td>HD 100546</td>
<td>20.1 × 18.0</td>
<td>−0.2</td>
<td>3.2</td>
<td>1317</td>
</tr>
</tbody>
</table>
Figure 5.12: HD 104237. Top row: unsmoothed with 1" pixels; second row: smoothed with half-beam Gaussian (FWHM = 9.3''); bottom: component positions plotted on smoothed map. Left: map; centre: fitted model of 2 effective beam-sized Gaussians (FWHM = 18.6 and 20.8'') respectively with separation of A and E components; right: maps with model subtracted. The table shows the position offsets and component fluxes for the fits (the separation vector and FWHMs are fixed).
Spectral Slopes

The (sub-)millimetre spectral slope of emission from dust is commonly used to infer information about the size distribution of dust grains. This works under the assumption that the emission at such long wavelengths is optically thin, so the presence of small dust grains causes the spectral slope to be steeper than a black-body (e.g. §1.5.2). Under a further assumption that the dominant dust temperatures are high enough that the Rayleigh-Jeans approximation holds at the wavelengths of interest, or that the dominant dust temperatures of discs are all similar, the spectral slopes of discs can be directly compared without considering the finite dust temperatures. Studies of these spectral slopes for HAeBes have previously been performed by Meeus et al. [2001], Natta et al. [2001], Acke et al. [2004]. The LABOCA photometry presented here now allows the sub-mm slopes for these four HAeBe to be determined when combined with previously published ∼1.3 mm photometry.

Table 5.5 presents the LABOCA photometry along with photometry at an effective wavelength of 1.27 mm from the literature, and the spectral power law exponent $\alpha$ with uncertainty computed from these using [Jester et al., 2001],

$$\alpha = \frac{\ln(F_1/F_2)}{\ln(\nu_1/\nu_2)} \quad \text{and} \quad \sigma_\alpha = \frac{1}{\ln(\nu_1/\nu_2)} \sqrt{\frac{\sigma_1^2}{F_1^2} + \frac{\sigma_2^2}{F_2^2}} \quad (5.5)$$

All the 1.27 mm photometry was performed with a single bolometer instrument on the Swedish-ESO Sub-millimeter Telescope (SEST, decommissioned in 2003). A problem arises that significantly different fluxes have been reported for all sources which have been observed multiple times. This is likely due in part to mapping [Henning et al., 1998] versus simple on-off [Henning et al., 1993, 1994] observing modes. For completeness all reported 1.27 mm fluxes, and $\alpha$ values computed from them, are given.

The computed values of $\alpha$ for HD 97048 are consistent with optically thin emission from small dust grains, similar to the ISM (opacity power law index $\beta \gtrsim 1.6$). For the other three discs, however, unusually small values of $\beta$ ($\sim 0$–1) would be required to explain the shallow spectral slopes by dust grain properties alone. Possible other explanations for the shallow slopes of these three discs are:

- The discs are optically thick at $\lambda \sim 1$ mm
- The 1.27 mm fluxes are elevated due to free-free emission in outflows, or gyrosynchrotron emission from stellar magnetic activity [Rodríguez, 2000, and references therein]
- There is a significant dust population with very low temperatures ($T \ll 30$ K)
Table 5.5: LABOCA 870 μm photometry from this work, 1.27 mm photometry from literature, computed spectral power law index $\alpha$ ($F_\nu \propto \nu^\alpha$), and disc masses for the four Herbig Ae/Be stars. The disc masses are computed from $F_\nu(870 \mu m)$ assuming a dust opacity of $\kappa_\nu(870 \mu m) = 0.35 \text{m}^2/\text{kg}$, $T_{dust} = 20 \text{ K}$ and a gas to dust ratio of 100:1 [as used by Andrews and Williams, 2005].

<table>
<thead>
<tr>
<th>Target</th>
<th>mee01 group</th>
<th>$F_\nu(870 \mu m)$ (mJy)</th>
<th>$M_{dust}$ ($M_\odot$)</th>
<th>$M_{disc}$ ($M_\odot$)</th>
<th>$F_\nu(1.27 \text{ mm})$ (mJy)</th>
<th>$F_\nu(1.27 \text{ mm})$ reference</th>
<th>$\alpha$</th>
</tr>
</thead>
<tbody>
<tr>
<td>HD 97048</td>
<td>I</td>
<td>2610 ± 131</td>
<td>647</td>
<td>0.196</td>
<td>700 ± 60</td>
<td>Henning et al. [1998]</td>
<td>3.63 ± 0.27</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td>452 ± 34</td>
<td>Henning et al. [1993, 1994]</td>
<td>5.84 ± 0.25</td>
</tr>
<tr>
<td>HD 100453</td>
<td>I</td>
<td>494 ± 43</td>
<td>72</td>
<td>0.021</td>
<td>265 ± 21</td>
<td>Meeus et al. [2002, 2003]</td>
<td>1.72 ± 0.32</td>
</tr>
<tr>
<td>HD 100546</td>
<td>I</td>
<td>1324 ± 66</td>
<td>124</td>
<td>0.038</td>
<td>660 ± 66</td>
<td>Henning et al. [1998]</td>
<td>1.92 ± 0.31</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td>465 ± 20</td>
<td>Henning et al. [1994]</td>
<td>2.89 ± 0.18</td>
</tr>
<tr>
<td>HD 104237 23&quot;</td>
<td>II</td>
<td>190 ± 26</td>
<td>25</td>
<td>0.008</td>
<td>92 ± 19</td>
<td>Henning et al. [1994]</td>
<td>2.00 ± 0.68</td>
</tr>
<tr>
<td>HD 104237 A</td>
<td>II</td>
<td>154 ± 24</td>
<td>20</td>
<td>0.006</td>
<td>66 ± 13</td>
<td>Henning et al. [1993]</td>
<td>2.92 ± 0.66</td>
</tr>
<tr>
<td>HD 104237 E</td>
<td></td>
<td>91 ± 24</td>
<td>12</td>
<td>0.004</td>
<td></td>
<td></td>
<td></td>
</tr>
</tbody>
</table>
CHAPTER 5. LABOCA DATA REDUCTION & OBSERVATIONS OF SOUTHERN CIRCUMSTELLAR DISCS

For HD 97048 it is also possible that these factors are significant, and the steep spectral slope may be due in part to emission from an envelope of ISM dust grains [Prusti et al., 1994, detected mid-IR emission extending to \( \sim 10\mu m \)].

Large surveys of T-Tauri discs around lower mass stars such as Andrews and Williams [2005] show that \( \alpha \sim 2 \) is typical (in fact they determine a value of \( \langle \alpha \rangle = 1.93 \pm 0.01 \) using 850 \( \mu m \) and 1.3 mm only). Andrews and Williams [2005] fit a disc model to 44 sources in their sample which have well sampled SEDs and from the fits determine disc masses. Their model assumes \( \kappa_v(850\mu m) = 0.35 m^2/kg \), \( \beta = 1 \) and a gas to dust ratio of 100:1. They use these results to determine a characteristic dust temperature of 20 K, which gives a best fit of masses computed assuming isothermal optically thin emission to the masses from their model fits (for the same \( \kappa_v \), \( \beta \) and gas to dust ratio). Disc masses for the four HAeBes in this work computed from the LABOCA photometry using these assumptions are given in Table 5.5. As HD 97048 has a significantly different \( \alpha \), this mass is likely an over-estimate.

With these masses, or the assumed characteristic temperature, the optical depth of the discs can be estimated by, \( \tau_v = \kappa_v M_{\text{dust}}/A = F_v d^2/(AB_v(T_{\text{dust}})) \), where \( A \) is the physical area of the disc perpendicular to to the observer. Assuming the area of a face-on uniform disc of radius \( r_{\text{max}} \), the optical depth corresponding to the total disc masses in Table 5.5 is,

\[
\tau(870\mu m) \approx 10 \left( \frac{M_{\text{disc}}}{M_\odot} \right) \left( \frac{r_{\text{max}}}{100 \text{AU}} \right)^{-2}.
\]  

(5.6)

Assuming \( r_{\text{max}} = 100 \text{AU} \) would imply that HD 97048, HD 100453 and HD 100546 have discs which are likely to be optically thick (\( \tau > 0.2-2 \)) at 870 \( \mu m \), and HD 104237 A and E have discs which are optically thin (\( \tau \lesssim 0.1 \)). These optical depths can explain the spectral slopes of HD 100453, HD 100546 and HD 104237 without requiring any emission from outflows or especially cold grains. The high mass and steep spectral slope of HD 97048 suggest that the (sub-)mm emission is dominated by a region at least 2–3 times larger than 100 AU in radius. This size scale could either correspond to a large disc, or to a remnant envelope.

Comparison with Previous Studies

Mees et al. [2001] (me01) classified HAeBes into two groups based on the spectral slope around 10–20 \( \mu m \)\textsuperscript{18}. me01 group II sources can be fitted by a single power law from near-IR wavelengths to at least 20 \( \mu m \). me01 group I sources show a similar power law short-ward of \( \sim 10\mu m \), but have additional emission at longer wavelengths which can be fit by a black-body in the mid-IR. Both groups are subdivided based on the presence of a silicate emission feature

\textsuperscript{18}The classification of HAeBes in me01 is similar to a previous, more theoretical, classification introduced by Hillenbrand et al. [1992]. It is unclear why Hillenbrand et al. [1992] is not cited by Mees et al. [2001].
at 10\,\mu m, although no group II sources are seen without this feature. Whilst mee01 used spectra covering approximately 2–40\,\mu m to classify sources, it has subsequently been shown that classification can reliably be achieved by using near-IR and mid-IR IRAS photometry [Acke et al., 2004, and references therein]. It has been proposed that the physical distinction between the two groups is in the geometry of the outer disc regions, with group I sources having flared outer discs which contribute the strong mid-IR excess, and group II sources having more slender self-shadowed outer regions which contribute far less mid-IR emission [Meeus et al., 2001, Dullemond and Dominik, 2004]. Example SEDs and a diagram of the physical interpretation are shown in Fig. 5.13. The mee01 classes of the four LABOCA targets, from Acke et al. [2004], are given in Table 5.5.

![Figure 5.13: Example SEDs for mee01 group I and II sources, and explanation by disc geometry. Taken from Dullemond and Dominik [2004].](image)

Acke et al. [2004] investigated how the (sub-)millimetre spectral slope of HAeBeSes compared between mee01 group I and II sources. They found average values of $\alpha$ of 2.60 and 2.06 for groups I and II respectively, with sample sizes of 13 for each (11 and 9 respectively with both sub-mm and mm photometry, with the remainder using mm and IRAS 100\,\mu m photometry only). Adding the four discs observed here does not significantly alter their results. The high $\alpha$ of HD 97048 is not uncommon among the group I sources, with V376 Cas and AB Aur having comparable values. Likewise, the low $\alpha$ of HD 104237 is typical of group II sources. Both HD 100453 and HD 100546 are classified as group I sources and have values of $\alpha$ lower than any of the group I sources in Acke et al. [2004, $\alpha \simeq 2.2$ for Elias 3-1 and HD 139614], however, given the uncertainties on $\alpha$ this is not significant. Further photometry at other wavelengths, e.g. ground-based 350\,\mu m with SABOCA or 250–500\,\mu m photometry with Herschel’s SPIRE instrument, would help to more accurately determine the spectral slopes.
5.5.4 Other Stars

Seven stars from more nearby young stellar associations were also observed in this programme. These are summarised along with their LABOCA photometry in Table 5.6. All of these were observed with the compact raster spiral scan pattern. HD 16978, HD 45627, HD 55279 and GSC 08056-00482 were observed during the October 2008 run, and HD 44627, HD 53842 and TWA 21 were observed in the December 2008 run. Flux calibration correction factors have been applied to the December data as discussed previously. The scans were reduced using the improved weak-source pipeline with final σ-clipping thresholds of ±3σ. The fluxes given in Table 5.6 are from a fit of a beam-sized Gaussian at the position of each star, and the quoted uncertainty is the map RMS around each star.

None of the sources are detected, however, as the distances to these stars are relatively small, useful dust mass upper limits can be determined. Dust mass upper limits computed from three times the map RMS, assuming a dust temperature of 30 K and $\kappa_\nu(870\mu m) = 0.17\,m^2/kg$, are listed in Table 5.6.

Table 5.6: Summary of stars from nearby associations, with LABOCA photometry and dust mass upper limits. Association membership and ages are from Torres et al. [2008]. Distances are from F. van Leeuwen [2007], Mamajek [2005] and Riaz et al. [2006].

<table>
<thead>
<tr>
<th>Object</th>
<th>Alt. Name</th>
<th>Assoc</th>
<th>d (pc)</th>
<th>Age (Myr)</th>
<th>SpT</th>
<th>$F_\nu(870\mu m)$ (mJy)</th>
<th>$M_d(30,K)$ ($M_{\odot}$)</th>
</tr>
</thead>
<tbody>
<tr>
<td>HD 16978</td>
<td>ε Hyi, HR 806</td>
<td>Tuc-Hor</td>
<td>46</td>
<td>30</td>
<td>B9 V</td>
<td>$4 \pm 4$</td>
<td>$&lt; 0.3$</td>
</tr>
<tr>
<td>HD 45627</td>
<td>AB Pic</td>
<td>Carina</td>
<td>46</td>
<td>30</td>
<td>K2 V</td>
<td>$3 \pm 7$</td>
<td>$&lt; 0.5$</td>
</tr>
<tr>
<td>HD 45681</td>
<td>AO Men</td>
<td>β Pic</td>
<td>39</td>
<td>10</td>
<td>K3.5 V ke</td>
<td>$0 \pm 5$</td>
<td>$&lt; 0.3$</td>
</tr>
<tr>
<td>HD 53842</td>
<td>HIP 32435</td>
<td>Tuc-Hor</td>
<td>56</td>
<td>30</td>
<td>F5 V</td>
<td>$4 \pm 14$</td>
<td>$&lt; 1.5$</td>
</tr>
<tr>
<td>HD 55279</td>
<td>HIP 33737</td>
<td>Carina</td>
<td>59</td>
<td>30</td>
<td>F5 V</td>
<td>$1 \pm 4$</td>
<td>$&lt; 0.5$</td>
</tr>
<tr>
<td>GSC 08056-00482</td>
<td></td>
<td></td>
<td>28.5</td>
<td>30</td>
<td>M2 Ve</td>
<td>$6 \pm 4$</td>
<td>$&lt; 0.1$</td>
</tr>
<tr>
<td>TWA 21</td>
<td>HD 298936</td>
<td>TW Hya</td>
<td>45</td>
<td>8</td>
<td>K3 Ve</td>
<td>$-7 \pm 11$</td>
<td>$&lt; 0.8$</td>
</tr>
</tbody>
</table>

These dust mass limits place any discs around these stars well below typical protoplanetary disc dust masses, and would be consistent with typical masses of sub-millimetre detected debris discs around more nearby stars (see Fig. 1.14). If there were protoplanetary discs around these stars, assuming a typical gas to dust ratio of 100 would give gas mass upper limits of $0.03-0.5M_{\text{Jup}}$, indicating that these discs are no longer able to form gas giant planets.

208
5.6 Summary & Potential Follow-Up

Observations at 870 μm have been obtained for 29 stellar systems with ages of approximately 5–30 Myr, including the 18 confirmed members of the η Cha open cluster. The median 3σ detection limit is approximately 20 mJy, however, there is considerable variation (3σ ~ 12–40 mJy) due to reduced instrument sensitivity during the December run.

Four bright Herbig Ae/Be systems have been detected, and the 870 μm photometry has allowed their (sub-)millimetre spectral slope to be determined for the first time. For the three brightest sources (HD 97048, HD 100453 and HD 100546) it is argued based on their 870 μm flux that their circumstellar discs will have significant optical depth. This is corroborated for HD 100453 and HD 100546, which have spectral slopes consistent with optically thick (black body) radiation. HD 97048, which has the largest dust mass of the four sources, however, has a steep spectral slope consistent with optically thin emission from primordial grains. This suggests that HD 97048 still retains a significant envelope. The fourth system, HD 104237, contains five stars, two of which (the A and E components), have previously been shown to have significant mid-IR excess. By fitting a two source model to the LABOCA map of the HD 104237 system it has been possible to measure separate masses for the A and E components. The ratio of the disc masses of HD 104237 A and E is approximately equal to the ratio of the stellar masses.

None of the T Tauri stars (with the exception of HD 104237 E), or the early-type stars in the η Cha cluster, were significantly detected. For the members of nearby associations, with ages of ≥10 Myr and distances of ~30–50 pc, the dust masses are constrained to be significantly below 1 M⊙. For the η Cha members, with ages of ~6 Myr and a distance of 94 pc, the 3σ dust mass upper limits are ~2–3 M⊙, with an upper limit being placed on the average disc mass of 0.7 M⊙.

These results indicate that these systems lack sufficient remaining solid material to form rocky planets or the cores of gas giant planets, adding to the prevailing view that planet formation is largely completed by an age of 10 Myr. Assuming a typical gas to dust ratio of 100 suggests total gas masses well below 1 M Jup, indicating that there is insufficient gas remaining in these discs to form the atmospheres of gas giant planets.

Such low disc masses for all stars in η Cha were not expected, given the variety of disc types and accretion rates indicated by optical to mid-IR wavelength observations (several Class II / flat spectrum / transitional object SED classifications, and several classic T Tauri stars). The low masses found here are, however, being confirmed by observations with Herschel at 70–160 μm, which show the SEDs to drop rapidly longward of ~100 μm, indicating small disc outer radii [Woitke et al., 2011].

There is a case for obtaining much deeper mapping of the η and ε Chamaeleontis associ-
ations with LABOCA or future facilities. The upper limits from the current observations are particularly high in the central region of η Cha. These related clusters are of particular interest due to their evolutionary state, which results in a wide variety of circumstellar discs from the massive protoplanetary discs around HD 104237 A and E, to debris discs around η Cha and other members. As already noted, the dust masses of several of these discs seem unusually low for the mid-IR SED properties and accretion rates, and sub-millimetre detections will be required to properly constrain the regions beyond $\gtrsim 10$ AU.

The four Herbig Ae/Be stars observed here are all excellent candidates for observing with ALMA due to their large fluxes and angular extents of a few arcseconds. As well as allowing discs to be spatially resolved, this would allow emission from discs, envelopes and outflows to be disentangled and the emission from multiple component stars to be distinguished.
Chapter 6

Sub-millimetre Study of $\epsilon$ Indi and $\alpha$ Centauri with LABOCA

6.1 Introduction

This chapter presents 870 $\mu$m mapping observations with LABOCA of two southern nearby systems of sun-like stars: $\alpha$ Centauri AB and $\epsilon$ Indi A+BC. The goal of this work was to test for the presence of cold debris discs by performing the first sensitive observations of these systems at wavelengths longer than 70 $\mu$m. This was motivated by the sub-millimetre detections with SCUBA of discs around the similar stars $\epsilon$ Eridani and $\tau$ Ceti at more northerly declinations [Greaves et al., 1998, 2004b].

The $\alpha$ Centauri system, at a distance of 1.34 pc, is the closest stellar system to the Sun. The system consists of a binary of G2 and K2 spectral type components (the AB pair) with an orbital semi-major axis of 24 AU, and an M5.5 spectral type companion (Proxima) with a projected separation of 10500 AU. $\alpha$ Centauri AB is one of the brightest systems in the southern sky, with a combined $V$ band magnitude of $\sim$0.29$^m$.

$\epsilon$ Indi is a K4 spectral type star at a distance of 3.62 pc from the Sun. $\epsilon$ Indi has a companion brown dwarf binary with a projected separation of 1456 AU from $\epsilon$ Indi A. The brown dwarf binary, referred to as the B and C components here, is the closest known brown dwarf binary to the Sun. The B and C components have spectral types of T1 and T6 respectively, and an orbit semi-major axis of 2.4 AU [Cardoso et al., 2009].

The author has been responsible for all data reduction and analysis, and authoring the accepted proposal to obtain second epoch observations of $\alpha$ Cen AB.
6.2 Observations and Data Reduction

A description of the LABOCA instrument, its relevant observing modes, and calibration and data reduction procedures was given in the previous chapter. All observations used here were performed in the compact raster spiral scan mapping mode as shown in Fig. 5.1. These observations provide reasonably uniform coverage, and very dense spatial sampling, over an approximately 10' diameter field. For α Cen AB the field was centred between the A and B components, although the separation of ∼8'' at these epochs is negligible compared to the size of the field. For ε Indi A+BC the field was centred between the A and BC components, which are separated by 6.7''.

The initial observations were performed on 10–16 November 2007 (excluding the 15th), and follow-up second epoch observations of the α Cen field were performed on 19 September 2009. The observations are summarised by night in Table 6.1. Before 15 November 2007 there was an obstacle in front of LABOCA in the receiver cabin which shadowed 32 bolometers on the Eastern edge of the array\(^1\). These bolometers were flagged as bad during the reduction of scans performed on affected nights. The effects of this are relatively minor – primarily a slight reduction in the eastern extent of the field of view.

Photometric calibration was checked by reducing scans of flux calibrators (including the primary calibrator Uranus), and was found to be within 10% of nominal during all nights.

6.3 ε Indi A+BC

The reduced map of the ε Ind field is shown in Fig. 6.1, with cut-outs centred on the A and BC components shown at the bottom. The rms noise, measured using the 25-point grid technique described in §5.4.3, for grids with two beam (37.2'') spacing centred on the A component, BC component and the map centre are 3.2, 3.5 and 3.6 mJy/beam respectively. Neither component is detected, and there is no extended emission visible (1' corresponds to 217 AU). Dust mass upper limits were computed using Eqn. (1.17) assuming an opacity of \(\kappa(870 \mu m) = 0.17 \text{ m}^2/\text{kg}\). As the mass limit is lower for higher dust temperatures, a conservatively cool temperature of 30 K was assumed. For unresolved emission the dust mass limit for a 3σ flux upper limit of 10.5 mJy is \(1.6 \times 10^{-3} M_\odot\). If it is assumed that emission is uniformly spread out over a circle of radius 200 AU, then this would be an area approximately 25 times the beam area \((\pi \text{FWHM}^2 / 4 \ln 2)\). In this case the uncertainty on the total flux would be elevated by a factor of \(\sqrt{25}\), giving a more realistic 3σ dust mass upper limit of \(8 \times 10^{-3} M_\odot\).

The non-detection of dust emission is consistent with Spitzer/MIPS observations of ε Ind,\(^1\)

Table 6.1: Observation summary. ‘Shadow’ flags nights before 2007-11-15 in which 32 bolometers were shadowed by an obstacle in the receiver cabin. $N_{scan}$ gives the number of usable scans performed each night, each of which includes three repetitions of the compact raster spiral mode (4 $\times$ 35 s spirals). $t_{int}$ gives the approximate total usable integration time (an overhead factor of two is typical, so a total of $\sim$ 22 hours of telescope time was used). PWV and EL give the mean precipitable water vapour depth and elevation angle respectively, and $\tau_z \sec \tau$ gives an estimate of the mean optical depth assuming $\tau_z = 0.06 + 0.32 \times$ PWV and $\tau = 90^\circ - EL$.

<table>
<thead>
<tr>
<th>Field</th>
<th>Date</th>
<th>Shadow</th>
<th>$N_{scan}$</th>
<th>$t_{int}$ (hours)</th>
<th>PWV (mm)</th>
<th>EL (°)</th>
<th>$\tau_z \sec \tau$</th>
</tr>
</thead>
<tbody>
<tr>
<td>$\epsilon$ Ind A+BC</td>
<td>2007-11-10</td>
<td>S</td>
<td>0</td>
<td>0.00</td>
<td>—</td>
<td>—</td>
<td>—</td>
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<tr>
<td></td>
<td>2007-11-11</td>
<td>S</td>
<td>11</td>
<td>1.28</td>
<td>1.08</td>
<td>43</td>
<td>0.59</td>
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<tr>
<td></td>
<td>2007-11-12</td>
<td>S</td>
<td>0</td>
<td>0.00</td>
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<td>S</td>
<td>9</td>
<td>1.05</td>
<td>0.99</td>
<td>41</td>
<td>0.57</td>
</tr>
<tr>
<td></td>
<td>2007-11-14</td>
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<td>5</td>
<td>0.58</td>
<td>0.61</td>
<td>27</td>
<td>0.56</td>
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<tr>
<td></td>
<td>2007-11-16</td>
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<td>14</td>
<td>1.63</td>
<td>0.74</td>
<td>40</td>
<td>0.46</td>
</tr>
<tr>
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<td></td>
<td>($)</td>
<td>39</td>
<td>4.55</td>
<td>0.88</td>
<td>39</td>
<td>0.54</td>
</tr>
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<td>$\alpha$ Cen AB</td>
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<td>S</td>
<td>1</td>
<td>0.12</td>
<td>1.36</td>
<td>44</td>
<td>0.71</td>
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<tr>
<td></td>
<td>2007-11-11</td>
<td>S</td>
<td>10</td>
<td>1.17</td>
<td>0.97</td>
<td>43</td>
<td>0.54</td>
</tr>
<tr>
<td></td>
<td>2007-11-12</td>
<td>S</td>
<td>7</td>
<td>0.82</td>
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<tr>
<td></td>
<td>2007-11-13</td>
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</tr>
<tr>
<td>total</td>
<td></td>
<td>S</td>
<td>39</td>
<td>4.55</td>
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<tr>
<td>$\alpha$ Cen AB</td>
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<td>1.87</td>
<td>0.25</td>
<td>34</td>
<td>0.25</td>
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</tr>
</tbody>
</table>

which show no significant photometric excess at 24 or 70 $\mu$m [Trilling et al., 2008]. The MIPS observations essentially rule out detectable emission from dust at temperatures above $\sim$ 30 K, however, the low luminosity of this star ($\sim$ 0.1$L_\odot$) means that the presence of dust at colder temperatures is quite conceivable. For example, the orbital distance from a star of this luminosity at which typical debris disc grains would have a temperature of 30 K is only $\sim$ 30 AU [Eqn. (1.14), with $L_\star = 0.1L_\odot$, $\lambda_0 = 100$ $\mu$m]. The debris disc detection rate for F–K type stars at 70 $\mu$m with Spitzer/MIPS is approximately 16% [Trilling et al., 2008], which makes $\epsilon$ Ind typical. Based on debris detection statistics for binary star systems of different separations [e.g. chapter 4 of this thesis and Trilling et al., 2007], the presence of the companion brown dwarfs at a distance of $\sim$ 1500 AU is unlikely to affect the probability of detecting a debris disc around the primary star.
Figure 6.1: LABOCA map of $\epsilon$ Indi A+BC, smoothed with the beam to give an effective beam of FWHM = 26.3". Top: $\sqrt{\text{weight}}$ map (proportional to $1/\text{rms}$). Middle: full map. Bottom: cut-outs centred on A (right) and BC (left) components. The cut-out fields and locations of the stars are indicated on the full field maps (top, middle).
6.4 α Centauri AB

The original reduced map of the α Cen field is shown in Fig. 6.2 (left column), with a cut-out centred between the A and B components at the epoch of observation (J2007.86; 14 39 31.88 −60 50 02.4) shown at the bottom. This field contains a significant amount of background galactic cirrus emission. This was expected as α Cen is located near the galactic plane, with a galactic latitude of only $l = -0.67\degree$.

A bright resolved feature, with a peak flux density of $\sim 120$ mJy, was detected approximately $50''$ from the position of the stars, as can be seen in the maps in Fig. 6.2. The stars themselves (with separation of approximately half the beam FWHM) were also detected at the centre of the field. The bright feature was potentially very exciting, as the projected separation of $\geq 70$ AU would correspond to a stable circum-binary orbit [Wiegert and Holman, 1997]. If this feature were associated with the system then we would potentially be seeing the aftermath of a recent collision between large planetesimals in a circum-binary disc. The dust mass of the clump, if at the distance of the α Cen system, would be approximately $1 \times 10^{-3} M_\odot$ (assuming $\kappa(870 \mu m) = 0.17 m^2/kg, T = 50 K$ and $F_\nu(870 \mu m) = 200$ mJy). This mass is approximately half the mass of Pluto, so the progenitors would have to be almost planet size bodies.

A search for known objects in the region was performed using the SIMBAD² database and inspection of various sky survey images. No relevant objects were found. All images of wavelengths shorter than 70 μm which were inspected showed strong contamination from the extremely bright α Cen AB within several arcminutes of the system. The Southerly declination of $-61\degree$ means that this region was not observed by SCUBA on the JCMT [Di Francesco et al., 2008].

To test whether the bright feature was associated with the α Cen system, or was an unrelated background object, it was proposed to repeat the LABOCA observations at a later epoch in order to compare the proper motion of the feature with the $\sim 3.7''$/yr proper motion of α Cen AB. This proposal was accepted (ESO programme 384.C-1025; PI: Phillips) and the observations were performed on a single night in September 2009. The reduced map of these observations is shown in Fig. 6.2 (right column), centred on the same position as the 2007 map in the figure. The analysis of these images is presented in the following subsections. The background cirrus structure has been used to align the maps to remove any pointing offset, and as a result it has been possible to perform high precision relative astrometry, allowing the proper motion of the bright feature to be determined to an accuracy of $1\sigma \sim 1''$/yr.

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²http://simbad.u-strasbg.fr/
Figure 6.2: LABOCA maps of α Cen AB produced from observations made in November 2007 (left) and September 2009 (right), smoothed with a half-beam FWHM Gaussian to give an effective beam of FWHM = 20.8′. The maps are centred half way between the A and B components at the epoch of the 2007 observations (14 39 31.88 −60 50 02.4) using only the astrometry from the scan data. Top: $\sqrt{\text{weight}}$ map (proportional to $1/\text{rms}$). Middle: full map. Bottom: central region.
6.4.1 Astrometry of the α Cen AB System

The motion of α Cen AB across the sky is complicated, and several factors needed to be accounted for in order to determine the positions and motions of the stars and their centre of mass at the observation epochs. The proximity of the α Cen AB system to the Solar System means that the binary is resolved, with typical separations of $\sim 10''$. The proximity also means that the system has a large proper motion, with magnitude of $\sim 3.7''/yr$, as well as significant heliocentric parallax of 0.747''. The astrometry was modelled by combining the orbit, the proper motion of the components, and the parallax of the system.

Orbital and Proper Motion

Orbital elements from Pourbaix et al. [2002] were used to determine the separation $\rho$ and position angle $\theta$ between the stars for any epoch using an algorithm from Meeus [1982]. The orbit of the B component relative to the A component is shown in Fig. 6.3. The most accurate measurements of the proper motion are those from Hipparcos [F. van Leeuwen, 2007], however, these proper motions are for the individual components and include the motions due to the orbit of the stars about the system centre of mass. The Hipparcos measurements for α Cen B have high uncertainties ($\sim 20\ mas/yr$ compared to $\sim 3\ mas/yr$ for the A component) due to the brighter A contaminating the observations. It was thus desirable to be able to use the Hipparcos astrometry for the A component for the system as a whole. To determine the positions of the stars at a given epoch, it was necessary to use the orbit and the masses of the stars to determine the proper motion of the centre of mass, and the position of the centre of mass at the Hipparcos epoch of J1991.25.

The derivative of the separation vector, $\textbf{r}(t) = [\rho(t) \sin \theta(t), \rho(t) \cos \theta(t)]$, determined using the orbital elements, relates the proper motion vectors of the two stars at any instant in time:

$$\mu_B(t) = \mu_A(t) + \frac{d\textbf{r}(t)}{dt}. \quad (6.1)$$

Introducing the centre of mass (CoM), which lies along the separation vector,

$$\mu_A(t) = \mu_{CoM} - \frac{m_B}{m_A + m_B} \frac{d\textbf{r}(t)}{dt} \quad \text{and} \quad \mu_B(t) = \mu_{CoM} + \frac{m_A}{m_A + m_B} \frac{d\textbf{r}(t)}{dt}, \quad (6.2)$$

where the centre of mass proper motion, $\mu_{CoM}$, is considered constant in time. $\mu_{CoM}$ can be determined from either component proper motion at a given epoch using the derivative of the separation vector computed from the orbital elements for that epoch:

$$\mu_{CoM} = \mu_A(t) + \frac{m_B}{m_A + m_B} \frac{d\textbf{r}(t)}{dt} = \mu_B(t) - \frac{m_A}{m_A + m_B} \frac{d\textbf{r}(t)}{dt}. \quad (6.3)$$
Figure 6.3: Orbit of $\alpha$ Cen B relative to $\alpha$ Cen A projected on the sky. The relative positions at the two epochs of LABOCA observations, and the Hipparcos epoch of 1991.25, are shown.

The masses of $\alpha$ Cen A and B are accurately known. Masses of $m_A = 1.105 \pm 0.0070 \, M_\odot$ and $m_B = 0.934 \pm 0.0061 \, M_\odot$ from Pourbaix et al. [2002] were used here. As stated above, the most accurate and reliable proper motion measurement for the system is that for the A component in F. van Leeuwen [2007]:

$$(\mu_A)_*\, (1991.25) = -3679.27 \pm 3.88 \, \text{mas/yr}, \quad (\mu_A)_\delta\, (1991.25) = 473.67 \pm 3.23 \, \text{mas/yr}. \quad (6.4)$$

The separation vector, computed using the orbital elements from Pourbaix et al. [2002], and its derivative at the same epoch are,

$$r_\alpha\, (1991.25) = -11.082068'' \quad \text{and} \quad r_\delta\, (1991.25) = -15.670225'' \quad (6.5)$$

and

$$\frac{dr_\alpha}{dt}\, (1991.25) = 91.916 \, \text{mas/yr} \quad \text{and} \quad \frac{dr_\delta}{dt}\, (1991.25) = 481.060 \, \text{mas/yr}. \quad (6.6)$$

Using Eqn. (6.3) the centre of mass proper motion (assumed constant in time) is then,

$$(\mu_{\text{CoM}})_\alpha = -3637.17 \, \text{mas/yr}, \quad (\mu_{\text{CoM}})_\delta = 694.03 \, \text{mas/yr}. \quad (6.7)$$
In the centre of mass frame the position vectors of the components are,

\[ r_A(t) = -\frac{m_B}{m_A + m_B} r(t) \quad \text{and} \quad r_B(t) = \frac{m_A}{m_A + m_B} r(t), \quad (6.8) \]

and the absolute positions are then,

\[ \alpha_A(t) = \alpha_{\text{CoM}}(t_0) + \left( t - t_0 \right) (\mu_{\text{CoM}})_\alpha \gamma - \frac{m_B}{m_A + m_B} r_\alpha(t) \cos \delta \quad (6.9) \]
\[ \delta_A(t) = \delta_{\text{CoM}}(t_0) + (t - t_0)(\mu_{\text{CoM}})_\delta - \frac{m_B}{m_A + m_B} r_\delta(t) \quad (6.10) \]
\[ \alpha_B(t) = \alpha_{\text{CoM}}(t_0) + \left( t - t_0 \right) (\mu_{\text{CoM}})_\alpha \gamma + \frac{m_A}{m_A + m_B} r_\alpha(t) \cos \delta \quad (6.11) \]
\[ \delta_B(t) = \delta_{\text{CoM}}(t_0) + (t - t_0)(\mu_{\text{CoM}})_\delta + \frac{m_A}{m_A + m_B} r_\delta(t) \quad (6.12) \]

**Parallax**

The astrometry from *Hipparcos*, which has been used here, is in the ICRS frame, which has the Solar System barycentre as its origin. To compute the astrometric position of an object for an observer on Earth, the Cartesian vector from the barycentre to the observer must be subtracted from the vector from the barycentre to the object. The Earth–barycentre vector was obtained for each epoch using a routine from the SLALIB astrometry library\(^3\), and the barycentre–object vector was computed from the ICRS position of each star, computed from Eqs.(6.9),(6.10),(6.11) and (6.12), and the system distance.

**Full Sky Motion**

The computed separation vectors and absolute ICRS positions for an observer on Earth are given in Table 6.2 and shown graphically in Fig. 6.4. These have all been generated from the *Hipparcos* position and proper motion for the A component, and the orbital elements from Pourbaix et al. [2002].

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\(^3\)http://www.starlink.rl.ac.uk/star/docs/sun67.hx/node96.html
Table 6.2: Positions of the centre of mass and A and B components of α Cen AB. The A position at epoch J1991.25 comes from F. van Leeuwen [2007], all other values are computed.

<table>
<thead>
<tr>
<th></th>
<th>CoM position</th>
<th>A position</th>
<th>B position</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>14 39 40.252</td>
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<td>14 39 30.945</td>
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<td>r_A</td>
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<td>3.21''</td>
</tr>
<tr>
<td>r_B</td>
<td>10.40''</td>
<td>4.55''</td>
<td>3.80''</td>
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</table>

Figure 6.4: Full sky motion as seen from Earth of α Cen AB, including proper motion and parallax. The origin corresponds to the ICRS (barycentric) position of the centre of mass at epoch J1991.25.
6.4.2 Determining Pointing Offset and Bright Feature Motion

Before the motion of any objects could be determined from the maps it was necessary to check for any offset between the pointings of the two maps. The maps have identical astrometry metadata, as the same centre position and dimensions were used in their production. However, any difference between the telescope pointing calibrations at the two epochs will cause the maps to be offset from one another. Due to the bright cirrus structure throughout the maps it was possible to measure the pointing offset by cross-correlating the maps⁴. All maps used below have a pixel scale of 1.0''/pix.

Normalised Cross-Correlation

The algorithm used for this analysis is the normalised cross-correlation. The justification for this comes from thinking of the overlapping regions of the images for a given offset \((\Delta x, \Delta y)\) as vectors of pixel values, and using the angle between them as a measure of their similarity, with zero angle for a perfect match. For two vectors \(a\) and \(b\) we have,

\[
\cos \alpha = \frac{a \cdot b}{|a||b|}.
\]  

(6.13)

This is unaffected by the lengths of the vectors (image contrast/gain), however, it is dependant on the mean component values of the vectors (image brightness). To remove the dependence on the means, they can be subtracted giving,

\[
\cos \alpha = \frac{(a - \overline{a}) \cdot (b - \overline{b})}{|a - \overline{a}||b - \overline{b}|}
\]  

(6.14)

Defining this as the correlation coefficient, \(C\), and replacing \(b\) by a version of \(b\) translated by \((\Delta x, \Delta y)\), in component form we have,

\[
C(\Delta x, \Delta y) = \frac{\sum_{x,y} (a[x,y] - \overline{a}) (b[x + \Delta x, y + \Delta y] - \overline{b})}{\sqrt{\sum_{x,y} (a[x,y] - \overline{a})^2 \sum_{x,y} (b[x,y] - \overline{b})^2}}
\]  

(6.15)

The sums, including those implicit in \(\overline{a}\) and \(\overline{b}\), are computed over all the pixels where both \(a[x,y]\) and \(b[x + \Delta x, y + \Delta y]\) are defined (i.e. the overlapping regions). In general \(\overline{a}\) and \(\overline{b}\) vary with \((\Delta x, \Delta y)\), so they are computed prior to the explicit sums in (6.15) for every \((\Delta x, \Delta y)\).

\(C\) can have values between \(-1\) and \(1\), with \(1\) corresponding to complete correlation, where the overlapping image regions are identical except potentially for mean and scaling.

⁴A potential improvement which has not been carried out is to cross-correlate maps of individual scans, or groups of scans, with the full maps, to remove pointing drift between scans (e.g. fitting a linear drift for groups of scans between pointing calibration observations). This would, however, require considerable effort, and the only effect would be to produce slightly higher resolution maps.
CHAPTER 6. SUB-MILLIMETRE STUDY OF ε INDI AND α CENTAURI WITH LABOCA

The offset between images was determined by varying \((\Delta x, \Delta y)\) over a suitable range of values, and determining the position \((\Delta x_0, \Delta y_0)\) of the peak of \(C(\Delta x, \Delta y)\). To determine \((\Delta x_0, \Delta y_0)\) to sub-pixel accuracy, a general two dimensional Gaussian of the form

\[
g(x, y) = A \exp\left(-[a(x-x_0)^2 + 2b(x-x_0)(y-y_0) + c(y-y_0)^2]\right),
\]

where

\[
a = \frac{\cos^2 \theta}{2\sigma_1^2} + \frac{\sin^2 \theta}{2\sigma_2^2},
\]

\[
b = -\frac{\sin 2\theta}{4\sigma_1^2} + \frac{\sin 2\theta}{4\sigma_2^2},
\]

\[
c = \frac{\sin^2 \theta}{2\sigma_1^2} + \frac{\cos^2 \theta}{2\sigma_2^2},
\]

was fitted to an \(11 \times 11\) grid of integer \((\Delta x, \Delta y)\) points centred on the maximum value. A two dimensional quadratic was also tried as the fitting function, however, the Gaussian produced significantly better fits both within the \(11 \times 11\) pixel central region and at larger distances from the peak.

**Cross-Correlation Masks**

It would have been possible to cross-correlate the full maps for the two epochs, however, the large range of noise (weight) across the field, and the inclusion of the moving stars and potentially moving bright feature, meant that the offset produced would have been suspect. To achieve meaningful offsets, masks were applied to the first epoch map prior to the cross-correlation. As only overlapping regions are used in the cross-correlation there was no need to mask the second image for small offsets as used here. Three masks were used, which are shown in Fig. 6.5. One mask removed all but a \(30'\) radius circle centred on the bright feature, which was used to determine the offset for the feature between the two maps. The other masks remove pixels with weights \((\propto 1/\text{Noise}^2)\) lower than 25% of the peak value, so that only high S/N areas of the map remain. One of these masks has a \(60'\) radius circle which covers the stars and bright feature removed, and the other has only a \(25'\) radius circle centred on the stars removed.

The resulting \(C(\Delta x, \Delta y)\) surfaces for \(-20 \leq \Delta x, \Delta y \leq 20\), obtained by cross-correlating the masked half-beam smoothed first epoch map with the corresponding half-beam smoothed second epoch map, are shown in Fig. 6.6. The fits of Eqn. (6.16) to the \(C(\Delta x, \Delta y)\) values within an \(11 \times 11\) pixel square centred on the \(\Delta x, \Delta y\) of the peak \(C(\Delta x, \Delta y)\) value are also shown. The fitted results using maps of the same pixel scale which have not been smoothed are similar, but there is significantly more noise and the cross-correlation values are correspondingly much
Figure 6.5: Cross-correlation masks applied to the 2007 map for determining the pointing offset between the two epochs. Maps have 1.0″ pixels. (a) full 2007 map; (b) retaining only 30 pixel radius circle centred on bright feature; (c) clipped to region with weight greater than 25% of the peak weight (equivalent to clipping to twice the minimum noise), with 60 pixel radius circle covering bright feature and α Cen AB removed; (d) clipped as in (c), but only 25 pixel radius circle centred on α Cen AB removed.

lower.

The estimated uncertainty on the fitted peak positions is approximately 0.5″, based on having tried a variety of slightly different masks and map smoothings. The maximum distance between the fitted peaks for the three masks is 1.8″. This difference appears between the cross-correlation for the bright feature alone and for the background with bright feature and stars removed. The mask with only the stars removed has its peak position between the other two,
Figure 6.6: Cross-correlation surfaces from cross-correlating the half-beam smoothed 2007 and 2009 maps after applying the three masks in Fig. 6.3. Fits of a 2D Gaussian to the central 11 \times 11 points are shown. Fixed scales are 1.0''/pix.
as may be expected. The offset between the feature and background of 1.8″ is significantly smaller than the 7.4″ motion of the stars’ centre of mass between the two epochs (Table 6.2), so it is concluded that the feature is not co-moving with the stars. It is possible that the 1.8″ offset is due to differing geometric distortion in the maps. Such distortion could be caused by uncertainties in the measured relative positions of the bolometers in the array, which are different for the two epochs.

The offset for the mask including both the background and the bright feature of (−0.17″, 3.20″) is taken here as the offset between the maps, and the bright feature is considered to be stationary relative to the background.

6.4.3 Stacking in the Frame of the Background
With the pointing offset between the maps known, a stacked map was produced by translating the 2009 map to remove the offset. The weight maps produced by BoA were utilised to produce a weighted stacked map.

Analysis of the Bright Feature
The weighted stacked map was used to fit a 2D Gaussian to the bright feature in order to estimate its angular extent and total flux, as shown in Fig. 6.7. A total flux of $F_\nu(870 \mu m) = 220 \pm 22\, mJy$ and deconvolved FWHM size of $24" \times 19"$ (major $\times$ minor) were determined from the Gaussian fits. This assumes a 10% uncertainty on the flux which allows for calibration uncertainty and some error due to deviation of the true profile of the feature from the fitted 2D Gaussian.

The feature is not seen in *Spitzer*/MIPS images of α Cen, so its emission must peak at a wavelength significantly longer than 70 μm. Taking a conservative upper limit 70 μm flux of 100 mJy implies a temperature upper limit of 15 K assuming a modified black body with $\beta = 2.0$ (18 K for $\beta = 1.0$). It is hypothesised here that this feature is likely to be a prestellar core.

Using a typical opacity for dust in prestellar cores of $\kappa(870 \mu m) = 0.10 m^2/kg$ [Launhardt et al., 2010]$^6$, and assuming a temperature of 15 K and a dust to total mass ratio of 1:150 [Launhardt et al., 2010], yields a distance dependent dust mass of,

$$M_{\text{tot}} = 5.3 M_\odot \times \left( \frac{d}{1 \text{kpc}} \right)^2.$$  \hspace{1cm} (6.20)

$^6$Launhardt et al. [2010] state an opacity of $\kappa_\nu = 0.5 \text{cm}^2\text{g}^{-1} (\lambda/1.3 \text{mm})^{-1.8}$, which they say is “a value that is typical for dense and cold molecular cloud cores and intermediate between the opacities usually adopted for unprocessed interstellar grains in the diffuse gas and for processed and condensed grains in protoplanetary disks.”
Launhardt et al. [2010] measure hydrogen masses for prestellar cores of 1–20 $M_\odot$, which corresponds to total masses of 1.3–34 $M_\odot$. For these maximum and minimum masses, the corresponding minimum and maximum distances for the bright feature are 0.50 and 2.53 kpc. At these distances the physical FWHMs of the feature are 9500 × 11800 AU = 0.05 × 0.06 pc and 48000 × 59000 AU = 0.23 × 0.29 pc respectively. Typical FWHMs for prestellar cores measured by Ward-Thompson et al. [1999], Kirk et al. [2005] and Launhardt et al. [2010] are ~ 0.01–0.1 pc. This suggests that the feature is likely to be at a distance of ~1 kpc and have a mass ~5 $M_\odot$ if it is indeed a prestellar core.

The $\alpha$ Cen field is close to the Scorpius-Centaurus-Lupus-Crux complex (see Fig. 6.8), which is a region of recent star formation at a distance of approximately 120 pc. If the bright feature
were a prestellar core at this distance then it would have a total mass of only \( \sim 0.07 M_{\odot} \) and physical FWHM size of \( \sim 0.01 \) pc. This mass is an order of magnitude lower than typical prestellar cores measured in the same manner, so it is likely that the object is more distant (the mass would be a factor of two larger if the dust temperature was assumed to be 10 K instead of the 15 K, but this does not alter the conclusion).

**Figure 6.8:** Member stars of the Scorpius-Centaurus-Lupus-Crux complex determined from *Hipparcos* proper motions. Figure taken from de Zeeuw et al. [1999]. The position of \( \alpha \) Cen AB is indicated by the circle+cross symbol at \((l,b) = (315.73^\circ, -0.68^\circ)\)

### 6.4.4 Stacking in the Frame of the Stars

By adding the difference between the two centre of mass positions for the two epochs \( (\Delta \alpha^* = -7.34'', \; \Delta \delta = +1.27''; \; \text{Table 6.2}) \) to the pointing offset, it was possible to create a weighted stacked map in the centre of mass reference frame. As the motion between the two epochs is considerably less than the beam size, the shifting and stacking has relatively little effect on suppressing the background structure, and the only significant improvement is an improvement in S/N. The 2D Gaussian fit to the background bright feature (Fig. 6.7) was subtracted from the maps prior to stacking in the centre of mass frame.

**Stellar Flux Prediction**

The photometry of the stars was predicted from shorter wavelength photometry in order to allow comparison with the photometry measured from the LABOCA maps. There is very little
resolved photometry for α Cen AB due to the combination of being too bright for any modern surveys, and the binary separation being too small for legacy photomultiplier photometry. To avoid the complexities of modelling the stellar photospheres using the optical and near-IR photometry that is available, 24 μm photometry has been used alone. α Cen AB has been observed in the MIPS-24 photometry mode used in chapter 4, so the reduction methods developed there were used to obtain MIPS-24 photometry. The centre of the PSF for both stars is saturated, so pixels near the centre were ignored in the PSF fitting. The MIPS-24 image before and after PSF subtraction is shown in Fig. 6.9. The resulting PSF fit magnitudes in the Vega system (see chapter 4 for details on calibration) are,

$$[24]_A = -1.568^m \pm 0.006^m \quad \text{and} \quad [24]_B = -0.704^m \pm 0.012^m. \quad (6.21)$$

Within the uncertainties of the LABOCA photometry, it was acceptable to assume zero [24] - [870] colour (i.e., assuming the spectrum from 24 to 870 μm is the same as Vega’s photosphere). The zero-point flux density at 870 μm was determined from the MIPS-70 zero-point flux of 778 mJy [Gordon et al., 2007], assuming a Rayleigh-Jeans ($F_\nu \propto \lambda^{-2}$) spectrum,

$$ZP(870 \, \mu m) = 778 \times (71.42/870)^2 = 5.24 \, \text{mJy}. \quad (6.22)$$

The resulting predicted 870 μm fluxes for the stars were then,

$$F_A(870 \, \mu m) = 22.2 \pm 1.1 \, \text{mJy} \quad \text{and} \quad F_B(870 \, \mu m) = 10.0 \pm 0.5 \, \text{mJy}. \quad (6.23)$$

An uncertainty of 5% has been adopted to account for systematic uncertainties in the MIPS-24 photometry and the simplistic method of transferring the MIPS-24 photometry to 870 μm.

**Stellar Flux Measurement**

The flux from the stars was measured from the LABOCA maps. As the stars were only marginally resolved (separation of 7-8″ at the LABOCA epochs, compared to the 18.6″ FWHM LABOCA beam), only combined photometry has been obtained. The maps were smoothed with a full beam Gaussian (18.6″ FWHM), and an effective beam sized Gaussian (26.3″ FWHM) with variable amplitude, background level and position was fitted to determine the flux. This was performed on the 2007, 2009 and weighted stacked maps with the bright feature subtracted (all in the same astrometry relative to the centre of mass, with the centre of mass at the centre according to the astrometry in the 2007 map) as shown in Fig. 6.10. The fitted flux of 28±3 mJy (5% calibration uncertainty for LABOCA assumed) is consistent with the predicted combined
Figure 6.9: MIPS-24 images (Epoch J2005.27). Left: original post-BCD image, shown with square root scale with cuts of 0 and 30%. Right: star subtracted version of left image after PSF fitting excluding the saturated regions at the centre (these are not subtracted either, and hence remain white in the subtracted image). North is up. Note that the bright feature, which is approximately 50″ due South of the stars at this epoch, is not visible at this wavelength.

flux of $32 \pm 1.6 \text{ mJy}$ from above. As expected from shorter wavelength photometry, there is no sign of unresolved excess flux. Visual inspection of the maps with the stars and bright feature subtracted (Fig. 6.10, right) shows no sign of any resolved ring-like circum-binary structure.
Figure 6.10: Photometry of the stars, performed by fitting an effective-beam-sized symmetrical Gaussian (FWHM = 26.3") to maps smoothed with a beam-sized (FWHM = 18.6") Gaussian. The pointing offset between the maps has been removed, and the bright feature has been subtracted from the original maps before smoothing. Top: 2007, middle: 2009, bottom: weighted stacked map with 2009 map shifted by (–7, –3) pixels to account for the motion of the centre of mass between the epochs.
6.5 Summary

This chapter has presented 870 $\mu$m continuum mapping observations of the nearby G/K spectral type stellar systems $\epsilon$ Indi A + BC and $\alpha$ Centauri AB. These observations were performed using the LABOCA bolometer array on the APEX telescope in Chile. The instrument and data reduction procedures were covered in the previous chapter of this thesis.

The map of $\epsilon$ Indi A + BC (Fig. 6.1) showed no detections, near either the primary star or its brown dwarf binary companion. The noise in the map was measured to be 3.5 mJy/beam, and a 3$\sigma$ dust mass upper limit of $8 \times 10^{-3} M_\odot$ was determined assuming a dust temperature of 30 K and an emitting area equal to a face-on 200 AU radius circle. Whilst $Spitzer$/MIPS observations also show a non-detection of debris [Trilling et al., 2008], the LABOCA non-detection rules out the presence of a massive cold disc ($T \leq 30$ K, $r \geq 30$ AU), which would have been undetectable by MIPS at 24 or 70 $\mu$m.

The map of $\alpha$ Centauri AB (Fig. 6.2) was considerably more complicated to interpret due to the presence of strong contaminating galactic cirrus emission. A bright resolved feature with deconvolved FWHM size of 24$''$ $\times$ 19$''$ and integrated flux density of 220 mJy was detected approximately 50$''$ from the stars. If the feature were associated with the $\alpha$ Cen system, then its projected separation of $\geq 70$ AU would have corresponded to a stable circum-binary orbit, and we could have been witnessing the aftermath of a collision between Pluto-mass bodies. An extensive search of catalogues and archival images yielded no previous detection of the bright feature.

The observations of $\alpha$ Cen were repeated two years after the original observations in order to ascertain whether the bright feature had a proper motion consistent with the motion of the stars, or was a stationary background object. The pointing offset between the two epoch maps, and the motion of the bright feature, were measured using cross-correlation (§6.4.2). The motion of the feature between the two epochs was measured to be 1.8$''$ with an estimated uncertainty of 1$\sigma \sim 0.5''$. Comparison with the 7.4$''$ motion of the stellar centre of mass between the two epochs led to the conclusion that the bright feature is a background object.

It is hypothesised based on the angular size, temperature upper limit and integrated flux that the bright feature is likely to be a prestellar core at a distance of 0.5–2 kpc. The galactic latitude of only $-0.67^\circ$ means that even at a distance of 2 kpc the object would be only 23 pc from the galactic plane.

Using the pointing offset determined by cross-correlation, and the astrometry of the stars, weighted stacked maps were produced from the two epoch maps in both the frame of the background and the frame of the centre of mass of the stars. These were used to measure the size and integrated flux of the bright feature, and measure the combined flux of the stars.
measured flux from the stars of $28 \pm 3\text{ mJy}$ was entirely consistent with a prediction made using
Spitzer MIPS-24 photometry. With both the bright feature and stars subtracted from the maps
(Fig. 6.10) there was no sign of any emission consistent with a debris disc around the $\alpha$ Cen AB
system.

It appears that $\alpha$ Cen AB is a further example of a binary system with a separation of
order 10 AU which lacks a debris disc. This is consistent with Spitzer/MIPS surveys such as in
chapter 4 of this thesis and Trilling et al. [2007] which show a lack of detectable debris around
binary systems with separations of approximately 3–150 AU.
Chapter 7

Conclusions & Future of the Field

7.1 Summary & Conclusions

This thesis presents observational studies of thermal emission from debris discs, and protoplanetary discs in the late stages of their evolution. Observations of debris discs are a window into the outer regions of planetary systems, which are currently inaccessible to most planet detection techniques. The presence and properties of debris discs have important implications for the development of life, for instance in the delivery of water to terrestrial planets and in the hazards posed by frequent bombardment. The debris discs which are detectable with current facilities are significantly more massive than the Solar System's debris disc. Debris disc surveys such as presented here aim to identify system properties which give rise to such massive debris discs. Ultimately this will lead to a statistical picture of how debris disc properties vary among systems, placing the the Solar System in context, and allowing the prospects for life around other stars to be evaluated.

7.1.1 Debris Disc Surveys

Chapter two provided a critical overview of previous surveys of debris discs performed with space-based mid/far-IR observatories such as IRAS, ISO and Spitzer, and ground-based submillimetre observatories. The scientific motivation for surveying large numbers of stars in search of debris discs is to determine what system properties (if any) give rise to the dusty discs which we can detect, and to build a picture of the range of properties of debris discs. Combined with the properties of detected exoplanets, the properties of debris discs set out the scenarios in mature and maturing planetary systems which planet formation theories and models must explain. For example, the correlation between the detection rate of gas giant planets with host
star metallicity, and the lack of correlation between debris disc detection rate with metallicity, can both be explained by the core accretion model of planet formation.

It was shown that the dust mass sensitivity of the previous mid/far-IR surveys is a strong function of spectral type for main sequence stars due to the luminosity dependence of dust temperatures. Such surveys have shown debris incidence rates to be much higher for B–F type stars than for later type stars, however, it is demonstrated that the number of stars as a function of spectral type for which a given mass of dust would be detectable varies in a very similar way to the debris detection rates (Figs. 2.2, 2.4). Observations at sub-millimetre wavelengths are shown to be essentially free from such spectral type bias (Fig. 2.3), however, the sensitivity of previous and current sub-mm instruments has only allowed small surveys (∼ 10 stars) to be performed.

**Seeing the Whole Picture: Longer Wavelengths & Unbiased Surveys**

The *Herschel* space observatory and the SCUBA-2 sub-mm instrument allow efficient surveying of hundreds of nearby stars at 70–850 μm. Two surveys in particular – the DEBRIS *Herschel* Key Programme, and the SUNS legacy survey with SCUBA-2 – which are designed to obtain unbiased debris disc statistics, were described. When these surveys are completed, and the results are combined with previous works, we will have for the first time a view of debris disc properties using the full spectrum of thermal emission from near-IR to sub-mm wavelengths.

In addition to improving SED coverage, *Herschel* provides the greatest ever dust mass sensitivity for debris discs around Sun-like and cooler stars (G–M type; Figs. 2.8, 2.7), and a factor of three improvement in spatial resolution compared to *Spitzer* (*Spitzer* 70 μm FWHM ∼ 21″ vs. *Herschel* 100 μm FWHM ∼ 7″). The increased resolution allows some discs to be spatially resolved [Matthews et al., 2010, Fig. 2.9], and allows *Spitzer* 70 μm detections to be verified by resolving contaminating background sources.

The DEBRIS and SUNS surveys aim to have minimal biases in order to provide maximum legacy value by allowing debris properties to be explored as a function of many system properties. In order to achieve this, the targets are drawn from five volume limited samples of systems with primary stars of spectral types A, F, G, K and M – the selection of which was presented in chapter three of this thesis. Multiple spectral type samples were used due to the large variation in the number density of systems with different spectral types. Each spectral type sample contains approximately 130 systems distributed over the whole sky with maximum distances of 8.6, 15.6, 21.3, 24.1 and 45.5 pc for M, K, G, F and A type systems respectively. For each system the best estimate of the distance, membership of resolved companions, and primary spectral type have been compiled, in addition to an extensive database of cross-identifications.
7.1. SUMMARY & CONCLUSIONS

of components in many common catalogues. This work was published in Phillips et al. [2010]. By-products of this work include an empirical luminosity – effective temperature relationship for the main sequence in the solar neighbourhood (Fig. 3.10, Eqn. (3.8)), binary separation distributions for systems with primaries of each spectral type (Fig. 3.12), and demonstrating that we still have an incomplete census of M type systems within 10 pc of the Sun (Fig. 3.6). These spectral type samples will also be applicable to other future studies of the statistical properties of nearby stars. The overall properties of the systems in these samples, such as the number densities of stars of each spectral type, have already been used in other works e.g. Bonnor and Wyatt [2010].

It has been unfortunate that the timing of these surveys has not allowed the incorporation of any results in this thesis (with the exception of summarising early DEBRIS science demonstration observations). At the outset of this work the intention was to primarily work on the SUNS survey, performing the target selection for the survey prior to taking observations and analysing a selection of detected discs and performing initial statistical analyses. The involvement in the target selection for SUNS naturally led into the development of the highly complementary DEBRIS survey with Herschel – initially in developing the proposal, and in preparation and support from proposal acceptance through to the current time.

Unbiased Survey Example: A-type stars with Spitzer/MIPS

A survey of the volume limited sample of A type systems developed for the SUNS and DEBRIS surveys was performed with Spitzer using the MIPS instrument observing at 24 and 70 μm. The motivation for this survey was many-fold:

- To obtain definitive, unbiased, debris disc incidence rates for A type stars by observing a large volume limited sample in the most homogeneous manner possible. For debris discs around A type stars the dust mass sensitivity of Spitzer/MIPS is roughly equal to that of Herschel, so the detection statistics should not change significantly in the near future.

- To test the results of previous surveys which may be affected by selection effects and inhomogeneities in the way subsamples have been analysed.

- To discover new nearby debris discs around A type stars which were missed by previous surveys due to selection biases.

- To provide complementary data for the SUNS and DEBRIS surveys at longer wavelengths by observing as many targets as possible with Spitzer/MIPS and producing a set of homogeneous and well calibrated photometry.
• To develop stellar photosphere flux prediction techniques and highlight limitations to the achievable accuracy.

A combination of previous observations of stars within the sample, and observations proposed specifically for this work [Phillips et al., 2008], were used. As far as possible all observations were reduced, and photometry was performed, in a homogeneous manner. The photometry of the stellar photospheres at the MIPS wavelengths was predicted using optical and near-IR photometry fitted by model photosphere flux distributions, with typical accuracies of 0.02–0.05\(\mu m\) (2–5\% in flux) achieved. Significant photometric excess was detected for \(\geq 29 \pm 5\%\) and \(\geq 32 \pm 5\%\) at 24 and 70 \(\mu m\) respectively, which is in excellent agreement with a previous work by Su et al. [2006] which studied a sample of A type stars in clusters and kinematic groups. This suggests that the likelihood of a system possessing a detectable debris disc is not strongly affected by age. The results from the DEBRIS and SUNS surveys will place these detection rates into context with Sun-like and cooler stars, as the longer observing wavelengths will allow the same dust mass limits to be reached for these.

The homogeneous nature of the photometry and definition of significant excess allowed the excess rates for various subsamples to be reliably explored. The most pertinent result was that the incidence rates for binaries with separations less than a few hundred AU are significantly lower than for single stars and wider binaries (which have indistinguishable rates). This result disagrees with a study of A and F type binary stars published by Trilling et al. [2007], which stated that excesses were more common around binary stars. This discrepancy, however, is explained here as being an artifact of the different (and far more relaxed) excess detection criteria used by Trilling et al. [2007] for their binary stars at 70 \(\mu m\). Statistical tests applied to the binary separation distributions of multiple systems with and without excess showed that these differ significantly, and no systems at all with separations between \(\sim 3–150\) AU were seen to have significant excess. The results from the DEBRIS and SUNS surveys at longer wavelengths will be required to determine whether the systems with this range of separations do indeed lack any detectable dust or if they possess cold circum-binary discs not detectable at the MIPS wavelengths.

Another important result is that no correlation is seen between metallicity and debris disc incidence for these A type systems. This is in agreement with results of other works [Greaves et al., 2006, Saffe et al., 2008] and contrary to the correlation between the presence of radial velocity detected planets and high stellar metallicity. This result is as expected from the core accretion model of planet formation, which predicts that all systems will build planetesimals, but only protoplanetary discs with large dust masses (i.e. high metal content) are able to form sufficiently massive cores to accrete gas before the gas is dispersed from the disc.
This survey discovered a dozen new debris discs, including an especially massive and potentially cold disc around HD 37594, and a second disc in the triple system CCDM 23489-2808 ABC. The detection of the disc around HD 37594 neatly demonstrated the advantage of the blind nature of the survey, as this system possesses no indication of youth and it did not show up in IRAS surveys due to confusion from nearby nebulosity. For all systems with significant excess (32 with excess at both bands plus 15 with excess in only one band), characteristic dust temperatures, orbital radii and masses have been determined from the MIPS photometry.

A Small Sub-mm Debris Disc Survey: α Centauri & ε Indi

Chapter six examined two of the Sun’s closest neighbours in the southern sky, α Centauri AB and ε Indi ABC, which were observed at 870 μm in search of cold debris discs. This is the first time these systems have been observed at wavelengths longer than 70 μm. No emission was detected around ε Indi A or its brown dwarf binary companion, with 3σ dust mass upper limits of $8 \times 10^{-3} M_\odot$ being determined for both the A and BC components.

The field of α Centauri AB contained significant contaminating galactic cirrus emission. A bright resolved feature (220 mJy integrated flux, $24 \times 19''$ deconvolved FWHM size) was detected 50'' from the stars, which if it were associated with the system could have been the aftermath of a collision between Pluto mass bodies in a circum-binary disc. Observations of α Cen AB were obtained at a second epoch, which showed that the bright feature is not co-moving with the centre of mass of the stars, and is most likely a background object. It is proposed that the bright feature is likely to be a presettar core at a distance of 0.5–2 kpc.

The photosphere emission from α Centauri AB was detected in the LABOCA observations, and the total flux was as predicted from mid-IR photometry. This is one of the first times that the photosphere of a main sequence star has been detected at sub-mm wavelengths. In the future sub-mm detections of stellar photospheres will become more routine with instruments such as Herschel/SPIRE, SCUBA-2 and ALMA, allowing the ability of stellar photosphere models to predict sub-mm fluxes of stars to be tested.

No evidence was seen for a debris disc around α Centauri AB. Given the separation between the stars of $a = 24$ AU, the lack of debris is consistent with the general lack of debris in systems with separations of $\sim 3 – 150$ AU, as found in chapter 4.

7.1.2 Sub-mm Observations of Young Southern Stars

Chapter five presented a project undertaken to obtain sub-millimetre photometry, using the LABOCA 870 μm bolometer array instrument on the 12 m APEX telescope, of southern stars with ages of 5–30 Myr which are part of a large Herschel survey (the GASPS Key Programme).
CHAPTER 7. CONCLUSIONS & FUTURE OF THE FIELD

The sample included four Herbig Ae/Be stars, several T Tauri stars, and all confirmed members of the ~6 Myr old η Chamaeleontis open cluster. Although only the Herbig Ae/Be stars were detected, useful upper limits were placed on the dust masses for the other stars. The non-detection of the known discs in some of the T Tauri systems in η Cha was unexpected at the time the observations were taken. Herschel observations have subsequently shown that the spectral energy distributions of these systems drop off rapidly longward of ~100 μm due to small disc outer radii [e.g. Woitke et al., 2011]. This is consistent with a trend of decreasing disc outer radius with age inferred from sub-mm CO observations of discs [Dent et al., 2005].

The dust mass limits of \( \ll 1M_\odot \) for the 10–30 Myr T Tauri stars, and \( < 2M_\odot \) for the ~6 Myr old stars in η Cham, add to the picture that the formation of rocky planets and the cores of gas giant planets must largely be completed by ages of 5–10 Myr. The implied total gas masses of \( \ll 1M_{\text{Jup}} \) indicate that there is insufficient remaining gas in these discs to form the atmospheres of gas giant planets.

The 870 μm photometry obtained for the four Herbig Ae/Be stars was combined with 1.3 mm photometry from the literature to estimate the spectral power law index and \( \beta \). For HD 100453 and HD 100546 the spectral index suggests the discs are optically thick, which agrees with the optical depth estimated from the fluxes, opacities and emitting region size. For HD 97048 the spectral index is consistent with optically thin emission from ISM-like dust grains (\( \beta = 2 \)), however, the fluxes strongly suggest that the disc should be optically thick, so the majority of the sub-mm emission is likely to arise from a remnant envelope. For the fourth Herbig Ae/Be star, HD 104237, it was possible to resolve a companion T Tauri star with a comparable dust mass to the disc around HD 104237. The relative dust mass of the secondary disc in comparison to the primary disc is larger than has been observed for other binary class I–II sources of similar separations, where typically the secondary is undetectably faint at (sub-)mm wavelengths [Patience et al., 2008, Jensen and Akeson, 2003]. The ratio of the dust masses in the primary and secondary discs is very close to the ratio of the stellar masses.

7.2 Future

7.2.1 Science

Two of the key motivations of surveys of circumstellar discs and planetary systems are to put the Solar System and its formation into context, and to discover the diversity of planetary systems. Observations of debris discs are currently by far the most efficient way of gaining insight into the outer regions of planetary systems, where planet detection techniques generally do not reach. To truly put the Solar System in context, and determine how common ‘clean’ systems like the
Solar System are, debris disc surveys will need to be able to detect dust masses equal to those in the Solar System around significant numbers of stars. Surveys with Herschel are currently pushing the dust mass detection limits for cool discs lower than ever before, but it will take more sensitive future instruments to allow true Kuiper Belt and Asteroid Belt analogues to be detected.

Whilst we have been discovering and characterising debris discs for over 25 years, still little is robustly known about how observable debris and parent planetesimal belts are affected by their environment. Widening the wavelength coverage of large surveys to include the far-IR and sub-millimetre will remove the observational bias against cold debris and low luminosity stars, allowing debris properties to usefully be investigated as a function of stellar mass for the first time. With the detection of hundreds of exoplanets in the last decade (approximately 500 planets in 400 systems at the time of writing\(^1\)), determining if and how the presence of such detected planets affects the observable properties of debris is a key area for investigation.

As well as exploring the architectures of mature systems, understanding the processes and timescales involved in the transition from protoplanetary discs to debris discs is a key area to be addressed by circumstellar disc surveys. The prevailing view is that protoplanetary discs dissipate within approximately 10 Myr, however, models of icy planet formation show that hundreds of Myr can be required to build planets at distances of tens of AU from stars [Kenyon and Bromley, 2008]. These models predict a peak in dust mass at ages when such icy planets grow to a size large enough to stir the planetesimals near their orbit into collisions. As we cannot observe the evolution of individual systems, large samples of systems with accurately known ages will be required to test the validity of models of protoplanetary discs dispersal, rocky and icy planet formation, and debris disc evolution.

\subsection{7.2.2 Observations}

\textit{Spitzer's Legacy}

The \textit{Spitzer} cryogenic mission ended in May 2009. During the five year mission the MIPS instrument performed targeted observations of well over a thousand stars of interest to debris disc studies, and thousands of other stars have been observed in the fields of other observations. These observations will be exploited for years to come, both in the analysis of the MIPS detections alone, and in the analysis of systems observed with current and future facilities. Members of the MIPS instrument team are currently in the process of building a comprehensive database of homogeneously produced MIPS photometry for stars [Su et al., 2010], which

\footnote{\url{http://www.esoplanet.eu/}}
combined with easy access to reduced MIPS images through the new Spitzer Heritage Archive\(^2\) will be a formidable contribution to future debris disc studies.

**Herschel**

The Herschel space observatory, described in §2.4, was launched on 14 May 2009 and is currently performing routine science observations. The debris disc Key Programmes, DEBRIS, DUNES and a Guaranteed Time project, are well underway. First results papers, based on Science Demonstration observations have been published [Matthews et al., 2010, Eiroa et al., 2010, Sibthorpe et al., 2010]. In addition to targeted surveys of stars there are searches underway for serendipitous detections of massive debris discs in wide area surveys [e.g. Thompson et al., 2010]. Herschel offers an impressive combination of the highest dust mass sensitivity of any current or previous instrument to cool dust in debris discs such as the Kuiper Belt, and spatial resolution sufficient to resolve larger debris discs around nearby stars (100 AU FWHM beam-size for \(d = 13\) pc at 100 \(\mu\)m). Herschel is ideally suited to discovering cold dust at large distances (\(\geq 100\) AU) from Sun-like and early type stars, and discovering discs around low luminosity K and M type stars. As suggested in Chapter 4, Herschel observations of binary stars will determine whether the dearth of discs detected by Spitzer around A type binary systems with separations of \(\sim 3–150\) AU is real or if some of these systems possess cold circum-binary discs.

**SCUBA-2 and Other Sub-mm Facilities**

The SCUBA-2 sub-mm instrument on the JCMT is currently expected to commence science operations in mid 2011. As described in §2.3, SCUBA-2 will offer the ability to survey hundreds of stars at sub-millimetre wavelengths, with sufficient sensitivity to detect cold debris discs. SCUBA-2 observations of discs detected at shorter wavelengths, e.g. by Spitzer and Herschel, will significantly improve modelling of discs by constraining the temperatures, dust masses, grain size distributions and radial extents. For stars located too far South to be observed efficiently with SCUBA-2 (\(\delta < -40^\circ\)), the LABOCA bolometer array on APEX is being used to follow up potentially cold debris discs detected by Herschel. In addition to improving the modelling of discs by increasing cold debris disc coverage, SCUBA-2 and LABOCA observations are crucial for planning observations with the ALMA sub-mm interferometer.

**ALMA**

The Atacama Large Millimetre Array [ALMA; Hills et al., 2010, and references therein] is a sub-millimetre interferometer located at an altitude of 5000 m in the Atacama region of northern

\(^2\)http://sha.ipac.caltech.edu/applications/Spitzer/SHA/
Chile. ALMA is of truly epic proportions compared to current sub-millimetre facilities. ALMA will consist of 64 12 m antennas and 12 7 m antennas, with baselines from tens of metres up to 15 km. An artist’s impression of the array is shown in Fig. 7.1. ALMA and its associated infrastructure are currently at an advanced stage of construction, with successful commissioning interferometry observations with three antennas at the site having been performed. The current schedule for ALMA is to start science observations in mid/late 2011 (16 antennas), and have the full array completed in 2013.

![Image of ALMA array](image_url)

Figure 7.1: Artist’s impression of ALMA. Image from [http://www almaobservatory.org/](http://www almaobservatory.org/).

ALMA will spatially resolve the sub-millimetre emission from circumstellar discs on AU spatial scales, allowing unprecedented insights into the architecture of the outer regions of planetary systems – a realm not probed by other planet detection techniques. Previous resolved sub-millimetre images of especially nearby discs [e.g. Holland et al., 1998, Greaves et al., 2005] have shown clumpy ring structures, however, repeat observations have shown the significance of the clumpy azimuthal structure to be questionable [e.g. Backman et al., 2009]. ALMA offers a hundred-fold increase in sensitivity and resolution, which will allow clumpy structure due to resonances with planets or recent collisions [e.g. Wyatt and Dent, 2002, Wyatt, 2006] to be reliably detected.

Surveys such as SUNS, DEBRIS and chapter 4 of this thesis, act as excellent pathfinders for ALMA. Due to the expected high demand for ALMA observing time, ALMA is unlikely to be usable for surveys where the majority of targets will not be detected. The observations of previous unbiased surveys will thus be required to evaluate the suitability of potential ALMA
targets by predicting the angular sizes and surface brightnesses.

**CCAT**

The Cornell Caltech Atacama Telescope [CCAT; Sebring, 2010] is a project to construct a 25 m single dish sub-millimetre telescope at an altitude of 5600 m near the ALMA site in Chile. CCAT is being designed as a sensitive wide field survey facility, with an intention to eventually implement $10^6$ element detector arrays (fully sampled 1° diameter field). The current schedule for CCAT envisages first-light in 2017 [Sebring, 2010], potentially using SCUBA-2 as a first-light instrument. Over the following years the implementation of larger field instruments and wavelength coverage from 200 μm to 2 mm is foreseen. CCAT offers the potential to perform circumstellar disc surveys of hundreds or thousands of targets, and the prospect of sensitive very wide area surveys may lead to the serendipitous discoveries of many interesting systems. The exceptionally dry site chosen for CCAT means that efficient surveying at 200–450 μm will be possible, with spatial resolution of 2–5'. The results of SUNS and the surveys with *Herschel* will be key to designing and evaluating proposals for circumstellar disc surveys with CCAT.

**JWST / MIRI**

The Mid IR Instrument [MIRI; Wright et al., 2004] on the 6.5 m *James Webb Space Telescope* (*JWST*) will provide diffraction limited imaging and spectroscopy in the 5–28 μm wavelength range. This offers incredible potential for spatially resolving warm dust ($\gtrsim 100$ K) in circumstellar discs. Smith and Wyatt [2010] show that the essentially all 24 μm detected debris discs around A type stars will have resolved emission detectable by MIRI. The spectroscopic capabilities will allow the mineralogy of small dust grains in discs to be studied, and the detection of tiny amounts of gas generated by sublimation from dust grains and comets.

**SPICA**

*SPICA* is a Japanese led space telescope design with a 3 m primary mirror cooled to 6 K, currently scheduled to be launched in 2018. SPICA will have instruments operating in the 5–210 μm wavelength range. SPICA’s cooled mirror and improvements in detector technology will offer far superior sensitivity to *Herschel*, allowing efficient surveys of hundreds of objects to the extragalactic confusion limit longward of $\sim 100$ μm. This capability may be used to perform a more sensitive DEBRIS like survey of nearby stars. Such a survey would have the potential to obtain calibration limited photometry of hundreds of stars in a similar manner to Spitzer’s achievements at 24 μm. This may be the first opportunity to truly be sensitive to Kuiper Belt masses of cool dust around a significant number of stars.
7.2. FUTURE

FIRI

Looking further into the future, the quest for increased spatial resolution and sensitivity to dust lead to the desire for a space-based Far-IR Interferometer (the far-IR is generally not observable from the ground, and payload size limitations rule out a sufficiently large single dish facility). Such a mission presents many challenges, however, the strong scientific motivation is driving studies of the FIRI concept [e.g., Helmich and Ivison, 2009]. It is possible that Antarctica could host a ground-based FIRI pathfinder instrument which would be used to develop the technologies and techniques required for a future space-based mission [e.g. Griffin et al., 2008].

GAIA

Throughout this thesis astrometric measurements including positions, proper motions, parallaxes and binary orbital elements have been used. The majority of this data, which is crucial in the planning and analysis of observations of nearby stars, came from the Hipparcos mission which was carried out in 1991. The GAIA mission\(^3\), to be launched in 2012, will perform a whole-sky astrometric survey producing sub-milliarcsecond astrometry for all stars brighter than \( V \sim 20 \) (c.f. \( V \sim 9 \) for Hipparcos). For stars brighter than \( V \sim 15 \) astrometric accuracies of \( \sim 10 \mu \text{as} \) will be achieved. Among many other things, this will lead to a massive leap forward in the inventory of known nearby M dwarf stars and brown dwarfs, and the detection of many planets and brown dwarfs in AU sized orbits around nearby stars. The GAIA astrometry will also lead to the discovery of many new members of stellar kinematic groups, and will also likely reveal new such groups. In addition to the astrometry, GAIA will obtain photometry and spectroscopy of all the stars it measures, yielding a wealth of stellar parameters. GAIA’s data products will revolutionise our knowledge of the structure and dynamics of the Milky Way, and will be a cornerstone of almost all areas of astronomy in the next decade and beyond.

7.2.3 Techniques & Ancillary data

Stellar Photosphere Flux Distribution Modelling

As the sensitivity and calibration of mid/far-IR observations improves, the limiting factor on the dust mass detection limits becomes the uncertainty of the stellar photosphere contribution to the observed flux (for spatially resolved emission the effect is less, however, the inner few AU will generally be unresolved). Chapter 4 of this thesis showed that photosphere prediction accuracies of \( \sim 2\text{–}5\% \) in flux are currently achievable, although this required considerable effort. Accuracies of \( \lesssim 0.5\% \) would be necessary to make the photosphere prediction contribution to

\(^3\text{http://www.rssd.esa.int/index.php?project=GAIA}\)
CHAPTER 7. CONCLUSIONS & FUTURE OF THE FIELD

MIPS-24 dust mass detection limits negligible. Future missions such as *SPICA* offer the potential of similar precision photometry at wavelengths up to \( \sim 200 \mu m \).

In Chapter 4 several areas of improvement applicable to at least A type systems were identified: modelling binary stars using both resolved and unresolved photometry and model flux distributions; developing model flux distributions including common chemical peculiarities and effects of oblateness caused by rotation; and obtaining high accuracy and high spatial resolution near-IR photometry of bright stars.

**Stellar Ages**

Currently one of the most significant limitations for the statistical analysis of nearby star debris disc surveys is knowledge of the ages of the target systems. This is an active area of work within the debris disc survey teams. It currently looks like it will be necessary to utilise many different approaches including kinematic group membership, X-ray luminosities, rotation periods, and evolutionary tracks in H–R diagrams. The ability to reliably determine the ages of systems is required to study the evolution of debris discs and planetary systems during the main sequence lifetime of stars.

**7.2.4 Summary**

Over the coming decade we can look forward to obtaining an unbiased census of debris in stellar systems, including low dust mass Solar System analogues. High angular resolution observations will resolve discs, allowing for the detection of planets through their interactions with the dust. Combined with the ever growing inventory of planets detected by other methods, and observations of protoplanetary discs, we will be able to build a detailed picture of the evolution and diversity of planetary systems and the environmental factors which influence these.
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