# DISCOVERING AND CHARACTERIZING EMBEDDED STELLAR CLUSTERS IN THE NEAR-INFRARED 

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SIGNED: Andrea Lynn Leistra

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## DEDICATION

For my partner, Jane Rigby, who kept me going.

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#### Abstract

I present a near-infrared search for new embedded stellar clusters in the Galaxy, and the results of near-infrared followup observations of a subset of newly discovered stellar clusters. I discuss the initial mass function of these embedded clusters and the implications of the apparent method-dependent systematic error in the IMF.

First, I present near-infrared $J, H$, and $K$ images of six embedded stellar clusters in the Galaxy, and $K$-band spectroscopy for two. I find a significant fraction of pre-main-sequence stars present in at least two of the clusters. For the clusters dominated by main-sequence stars, we determine the initial mass function (IMF) both by using the $K$ luminosity function and a global extinction correction and by deriving individual extinction corrections for each star based on their placement in the $K$ vs. $H-K$ color-magnitude diagram. Based on our IMFs we find a significant discrepancy between the mean IMF derived via the different methods, suggesting that taking individual extinctions into account is necessary to correctly derive the IMF for an embedded cluster. I find that using the KLF alone to derive an IMF is likely to produce an overly steep slope in stellar clusters subject to variable extinction, and examine literature results to see if the same effect exists in the work of other authors. I conduct a two-phase search of the 2MASS Point Source Catalog to discover previously unknown embedded stellar clusters and construct a more complete sample than has previously been available. Based on comparisons with the sample of known embedded stellar clusters we determine the completeness of the total existing sample to be $\sim 75 \%$ within 2 kpc . I discuss the limitations of previously employed algorithms for stellar cluster detection, and suggest possible alternatives for use in areas of high stellar background den-


sity. Finally I present a detailed look at two of the incorrectly identified embedded cluster candidates to better understand the limitations of our algorithm.

## Chapter 1

## INTRODUCTION

Although massive stars are both intrinsically rare and short-lived, thus making up less than $1 \%$ of the stars in our galaxy, they are critically important in understanding star formation and the interstellar medium. Through their stellar winds, strong ionizing radiation, and eventual supernovae they shape not just the formation of lower-mass stars in their immediate vicinity but enrich the interstellar medium to influence the formation of other stars even much later. Since massive stars are so short-lived, with O stars living for less than 10 million years, almost any massive star is a young massive star, and is near its formation site. Understanding the present star formation in the Galaxy thus requires understanding where the massive stars in our Galaxy are forming.

This question breaks down naturally into two questions: 1. Where are stars forming in our galaxy? and 2. Do massive star formation regions directly track lowmass star formation regions, or are they biased - does the initial mass function vary, such that some star-forming regions from proportionally more massive stars than others?

The first question is the easier of the two to answer, and has been addressed many times. Star formation is an extremely complicated process, and different wavelengths are best suited for observing different stages of this process. The earliest stages, cold, dense molecular cores, are best probed by sub-millimeter and radio surveys that track cold dust and dense gas. Slightly older regions are tracked by far-infrared studies of warm dust, sensitive to the infalling envelopes around protostars; only once the star-formation process has advanced to the point that nuclear fusion has begun can newly-formed stars be effectively located in the near-infrared and at shorter wavelengths. Until recently, no all-sky survey com-
parable to the radio and far-IR surveys had been conducted in the near-infrared (NIR). This left a gap in our understanding of star formation in the Galaxy; while star-formation regions could be located, the clusters forming within them remained unknown unless they were old enough (and therefore clear of enough of their natal molecular cloud) or nearby enough to be detectable at optical wavelengths. The Two Micron All-Sky Survey (2MASS) has partially bridged this gap, providing a survey that can provide a primarily distance-, rather than extinction, limited survey of embedded stellar clusters. The extinction limitation of optical surveys translates into a limitation both in distance, due to line-of-sight extinction in the Galactic Plane, and age, due to internal extinction within embedded clusters. Although the relatively bright limiting magnitude of 2MASS ( $J=15.8, H=15.1, K=14.3 \mathrm{mag}$ for a $10 \sigma$ detection) restricts it to probing relatively nearby regions of the Galaxy for embedded clusters (within $\sim 2 \mathrm{kpc}$ ), within this range it can probe to very high extinctions ( $A_{V} \gtrsim 30$ ) and thus very young, embedded clusters. It thus more nearly allows a distance-limited sample than previous, extinction-limited optical surveys have done.

The second question is considerably more difficult, though equally well-studied. The idea that the initial mass function (IMF) is universal, rather than entirely dependent on local conditions, was first put forth by Salpeter (1955), who found that the number of stars formed with a given mass follow a power-law with the form $\frac{d N_{*}}{d(\log M)} \propto M^{\Gamma}$ with a slope of $\Gamma=-1.35$. In the last fifty years this result has held up remarkably well, with intense scrutiny (nearly 1900 citations to Salpeter's paper as of May 2006) - although the errors in most determinations of the initial mass function are large, in the intermediate mass range covered by Salpeter's original data (well-covered between $\sim 0.3$ and $\sim 10 \mathrm{M}_{\odot}$ ) there have been no results convincingly shown to be inconsistent with this form and slope
of the IMF.
An alternative form for the IMF was suggested by Miller \& Scalo (1979), who fit both a log-normal form and a broken power-law to the IMF, proposing a higher slope (i.e. a more rapidly dropping IMF) at large masses, with power-law coefficients of $\Gamma=-0.4$ for $0.1 \leq M \leq 1 M_{\odot}, \Gamma=-1.5$ for $1 \leq M \leq 10 M_{\odot}$, and $\Gamma=-2.3$ for $M>10 M_{\odot}$. Revisiting the issue, Scalo (1998) suggested that the IMF was not truly universal and instead varied from region to region, but offered the caution that an alternative option, of uncertainties large enough that "little can be said about an average IMF or IMF variations", remained a possibility.

Although the "Miller-Scalo" form of the IMF has been frequently employed, especially in studies of extragalactic populations, the claim of IMF variations in the high- and intermediate-mass range has been less widely embraced. While Scalo was analyzing IMF data up to 1998 and finding evidence for variations, Massey et al. (1995,a) found at the same time that the IMF in Local Group stellar clusters was consistent with a Salpeter slope across a wide range of cluster masses and metallicities, while the massive-star IMF in the field had a steeper slope, finding $\Gamma=-4.0$. This has been interpreted to mean the IMF may be different for massive stars forming in near-isolation, though it may be a result of the small reservoirs of gas available to form the isolated, non-cluster stars.

Modern IMF studies frequently focus on clusters, since their coeval nature means the conversion from the measured, present-day mass function to the true IMF is simpler than for field stars. Additionally, since the relationship between stellar mass and lifetime flattens out for O stars, in a coeval cluster the presence of O stars is an indicator that the effects of stellar evolution on the observed mass function will be small, and little correction from the observed mass function to the true initial mass function will be necessary, other than in accounting for mass
loss on the main sequence.
In his 2003 review of massive stars, Massey (2003) cites the case of NGC 6611, where the original study by Hillenbrand et al. (1993) found an IMF slope of $\Gamma=-1.1 \pm 0.1$ and a reanalysis of the same data with a different treatment of extinction (Massey et al., 1995) found a slope of $\Gamma=-0.7 \pm 0.2$. The formal $1 \sigma$ error bars do not overlap, which suggests that the statistical uncertainties quoted in IMF determinations may not reflect the true uncertainties, and that systematic effects not reflected in the quoted errors must be taken into account when making comparisons between IMFs, and especially when making claims about variations or the lack thereof.

The similarity of the upper mass limit in very different environments suggests that the fundamental star formation processes are quite similar, and thus that the overall massive-star IMF may also be insensitive to environment. Several recent studies have suggested that the lack of any observed stars more massive than $\sim 120 \mathrm{M}_{\odot}$ is not just a statistical artifact of the decline of the IMF toward high masses, but represents a fundamental upper limit to the mass of stars. The R136 cluster in the 30 Doradus star-forming region, with a LMC metallicity of $Z \simeq 0.008$, and the Arches cluster near the Galactic Center, with approximately solar metallicity ( $Z \simeq 0.02$ ) (Najarro et al., 2004) both have a statistically significant absence of stars more massive than $\sim 120 \mathrm{M}_{\odot}$ (Figer, 2005; Oey \& Clarke, 2005). If all the clusters are treated together, the lack of observed stars more massive than $\sim 120 M_{\odot}$ is strongly inconsistent with a non-truncated IMF, where stellar masses would be limited only by the total mass available in the molecular cloud rather than by other physical processes, and suggests an upper mass limit in the neighborhood of $150 \mathrm{M}_{\odot}$. Even if the R136 cluster in 30 Doradus is considered in isolation, taking into account the suggestions of Weidner \& Kroupa (2006)
that, for example, ten clusters of $10^{4} \mathrm{M}_{\odot}$ each may not be equivalent to a single cluster of $10^{5} \mathrm{M}_{\odot}$, the results are similar, though the constraint on the upper mass limit becomes weaker. While the actual limiting stellar mass has not been terribly well-constrained, and can be said with confidence only to lie between $\sim 100$ and $200 \mathrm{M}_{\odot}$, it seems clear that such a limit does exist, and that no metallicity dependence has been observed. This suggests that any metallicity dependence in the massive-star IMF, such as has been suggested to produce a very top-heavy IMF for zero-metallicity Pop III stars, takes effect only at metallicities lower than can be observed in the Local Group.

The similarity of the IMF slopes for massive and intermediate-mass stars determined across a range of metallicities and star-formation densities suggests that, if variations in the IMF do exist, they are either random in nature, such that true IMFs exist in a range centered around the Salpeter slope (such a result would be extremely difficult to distinguish from a truly universal Salpeter IMF subject to observational errors and statistical fluctuations due to the finite number of stars in any single cluster), or that variations appear only under conditions more extreme than those found locally, such as the zero-metallicity conditions under which Population III stars formed or the extremely high star formation rates of starburst galaxies.

Although no results for masses larger than $1 \mathrm{M}_{\odot}$ have convincingly shown an IMF with a slope different from the Salpeter value (claims of a flat IMF, with a slope of $\Gamma=-0.7$, for the Arches cluster near the Galactic Center (Figer et al., 1999) have been attributed to mass segregation (Stolte et al., 2002)), the situation is less clear near and below the hydrogen-burning limit, a regime not probed by Salpeter's original study, and which is beyond the scope of this thesis.

It seems most likely that, for massive and intermediate-mass stars:

1. All local results are consistent with a Salpeter slope; deviant values cannot be ruled out, nor can a universal Salpeter slope.
2. Any variations from a universal Salpeter slope do not appear to correlate with obvious characteristics of the star-formation regions such as stellar density, star formation rate, or metallicity.
3. Real, systematic (i.e. non-random) deviations from the local Salpeter slope may exist in extreme conditions that do not exist locally and which cannot be probed by IMF methods that rely on resolving individual stars.

The situation for low-mass stars is more complex and is beyond the scope of this thesis.

Despite fifty years of intense study, the question of the universality of the massive-star IMF has not been put to rest. The multiplicity of methods for determining the IMF may contribute to this controversy; if a re-analysis using the same general method but different parameters (e.g. a different extinction law, as in the case of NGC 6611, different evolutionary tracks or mass/luminosity relations, etc.), then the use of an entirely different method could certainly have a comparable, or even more pronounced, effect. I will consider only methods used for Local Group clusters and field stars, when individual sources can be resolved; the types of methods used for integrated populations in more distant galaxies, such as determining the hardness of the ionizing field to determine the shape of the massive-star IMF, are beyond the scope of this thesis.

The most accurate and reliable method of determining the initial mass function, which requires the fewest steps between the observation and the inferred stellar mass, is to obtain spectra for a substantial fraction of all stars in the cluster, then convert the spectral type to mass. Perhaps the most ambitious example of this approach is the spectroscopic study of Hillenbrand (1997) covering the Orion

Nebula Cluster (ONC). She obtained optical spectra for $\sim 900$ stars in the cluster, which combined with optical and near-IR photometry gave an extremely wellcharacterized stellar IMF, found to be consistent with the Miller-Scalo formulation. Combined with the later Slesnick et al. (2004) spectroscopic determination of the substellar IMF in the ONC, which found a peak at $\sim 0.2 \mathrm{M}_{\odot}$ followed by a steep decline before the IMF leveled off in the substellar regime, this represents the most complete spectroscopic IMF determination to date, to the extent that it is difficult to compare to other works since nothing comparable exists.

This project was a major undertaking even for the relatively nearby ONC, the nearest region where massive stars are currently forming, and most IMF determinations rely at least in part on photometric data. When spectra are not available, clearly a direct conversion of spectral type to stellar mass is not possible; instead, stellar luminosity must be derived and converted to mass via theoretical evolutionary models. Determining the luminosity requires knowledge of the distance and extinction for the stellar cluster. Since all stars in a cluster lie at a common distance, a distance determined from kinematic data for an associated radio source, or from photometry and a spectrum of a single star, is sufficient for the entire cluster (so long as the distance is much larger than the size of the cluster, as is the case for all embedded clusters in the Galaxy). The situation for handling extinction, which for embedded clusters can vary significantly from star to star, is somewhat more complex, and (even apart from different extinction laws) has produced different methods of determining the IMF for massive and intermediate-mass stars in a stellar cluster.

The simplest method is to use a single extinction, determined from the main sequence locus, and correct the observed magnitudes of each source by the same amount. This method has been used with near-infrared photometry by (e.g.

Figuerêdo et al., 2002; Blum et al., 2000). If the distribution in extinction is narrow, and stars are randomly distributed in extinction around the fiducial value used, this method will have larger uncertainties than individually deriving extinction corrections for each source, but should not produce systematically differing results. An approach using multicolor photometry determines individual extinction corrections for each star, by assuming an isochrone and de-reddening each source until it lies on the isochrone. For the upper main sequence, this method produces a single value for the mass of the star; for lower-mass stars, T Tauri stars can complicate the situation and produce degeneracies (see, e.g., Figure 1 of Meyer, Calvet, \& Hillenbrand (1997)). If, however, stars are not randomly distributed around the average extinction, but instead suffer from a correlation between mass and extinction, this method is more reliable than the assumption of a common extinction.

Two factors might be expected to produce such an effect: the tendency of massive stars to form in the central, densest parts of stellar clusters would suggest the most massive stars in a coeval population lie in regions of high extinction, while the strong winds and radiation from massive stars would suggest that once they form they clear away local gas and dust more rapidly, producing the opposite effect. It is thus not immediately clear whether such an effect would be expected at all, or what the sign would be; nevertheless, it should be looked for.

In this thesis I take advantage of the 2MASS Point Source Catalog to discover previously unknown embedded clusters in the Galaxy, and follow up the discoveries with near-infrared photometry and spectroscopy of a subset of the clusters to determine the IMF of the massive and intermediate-mass stars and compare IMF determination methods when used on the same datasets.

In Chapter 2 I describe the search methodology used for a two-phase search of
the 2MASS Point Source Catalog for embedded clusters and discuss the results, as well as presenting a more detailed examination of two objects which point to weaknesses of the algorithm. In Chapter 3 I present the southern subsample of 2MASS-selected embedded cluster candidates, for which we obtained JHK images and K-band spectroscopy, with the massive- and intermediate-mass stellar IMF and discussion of the embedded clusters. In Chapter 4 I present the northern subsample, for which we obtained only JHK images, and discuss the IMFs and the individual clusters. In Chapter 5 I present conclusions and suggest possible directions for future work in this area.

## CHAPTER 2

## Discovering Infrared Embedded Clusters with 2MASS

### 2.1 Introduction

Optically-selected samples of embedded clusters are limited by extinction, both internally and along the line of sight. Lada \& Lada (2003) estimate that their catalog of known embedded clusters is only $\sim 30 \%$ complete within 2 kpc of the Sun, and less so at greater distances. Since most stars are formed in embedded clusters (Elmegreen et al., 2000) this corresponds to an incomplete knowledge of where stars in the Galaxy are forming away from the solar neighborhood. The near-infrared (NIR) offers the potential to expand the sample of known embedded clusters to greater distances and deeper degrees of embedding (and thus younger ages). The Two Micron All-Sky Survey (2MASS) provides a near-IR point source catalog of the entire sky, ideal for unbiased, IR-selected searches. The ability of 2MASS to detect embedded clusters is limited by its sensitivity and resolution, rather than by extinction; for the most part, it should be capable of providing a complete sample out to some distance. The completeness limits of the 2MASS survey are $J=15.8, H=15.1, K_{s}=14.3$, with the detection limits approximately $0.5-1$ mag fainter. With these limits, an A0 star will be detectable to a distance of $\sim 5500 \mathrm{pc}$ in the absence of extinction; at $A_{V}=30$ this is reduced to $\sim 1400 \mathrm{pc}$. While more massive stars will be detectable across much of the Galaxy, their detection alone is insufficient to establish the existence of a cluster since a mere 4 or 5 stars are unlikely to stand out enough to be recognizable as a cluster, and stars must be detected to low enough masses that the enhancement in stellar density can be observed. Thus, we expect that a 2 MASS -selected sample will extend the range of known embedded clusters in an extinction-dependent
way; those with the highest internal extinction will be detectable only nearby, while less deeply embedded clusters (though still too heavily embedded for the optical at $A_{V} \sim 5-10$ ) can be observed to greater distances. The low resolution of 2MASS presents a further constraint. Embedded clusters are compact, with most of those with a known size in the compilation of Lada \& Lada (2003) having a size of $<2 \mathrm{pc}$. For a cluster of $\sim 100$ members this suggests an average separation of $\sim 0.01 \mathrm{pc}$, suggesting that the resolution of 2 MASS will begin to limit point-source-based detection of embedded clusters beyond $\sim 2 \mathrm{kpc}$. The blending of the most closely spaced stars into a single 2MASS detection, while it would influence derived properties of the cluster, will not affect the detectability of the cluster as a whole. Thus, since stars in a real cluster will not be distributed uniformly throughout the cluster volume, we expect that clusters will remain detectable to larger distances than this simple calculation predicts, since they will be detected by 2MASS as clusters of point sources each made up of several closely spaced stars. Even at greater distances, searches based primarily on visual inspection (e.g. Dutra et al., 2003b) will be able to detect cluster candidates based on the presence of extended emission.

Several searches of portions of the 2MASS database have been conducted. (e.g. Dutra \& Bica, 2000, 2001; Ivanov et al., 2002) These have produced a number of cluster candidates, and many have proven to be genuine embedded clusters (e.g. Dutra et al., 2003; Leistra et al., 2005). However, these surveys remain incomplete, and have produced a number of false detections (Dutra et al., 2003); a systematic approach covering the entire sky, with selection criteria improved by the results of followup from these earlier studies, can provide a more complete and less contaminated sample of embedded clusters in the Galaxy.

Subjective searches based solely on visual inspection (e.g. Dutra et al., 2003b)
have located many embedded cluster candidates, and may include all embedded clusters associated with radio and optical nebulae (though some candidates are difficult to confirm as clusters based on 2MASS data due to their distance and confusion). However, a more quantitative approach will help confirm genuine clusters, as well as uncover potential infrared embedded clusters unassociated with optical nebulosity.

We present the results of a 2-phase search of the 2MASS database for embedded cluster candidates. In Section 2.2 we describe the methods used for the targeted and full-sky searches of the database. In Section 2.3 .1 we discuss the results, both the candidates we recover and the known cluster candidates we fail to recover, and in Section 2.5 we provide a more detailed look at some of the false cluster detections. Finally, in Section 2.4 we discuss the implications of our study for future embedded cluster searches, and possible methods for superior cluster searches.

### 2.2 Searching the 2MASS Catalog

### 2.2.1 Algorithm

We chose to use a color selection in addition to a stellar density criterion; previous automated studies (Dutra \& Bica, 2000) used only a stellar density selection. Visual inspection and deeper, higher-resolution imaging (Dutra et al., 2003b) of the cluster candidates found by these authors suggested that a significant fraction $(\sim 50 \%)$ of the candidates were chance superpositions rather than genuine clusters. In addition, since our primary interest was in young, embedded clusters, color selection provided a mechanism for preferentially selecting reddened clusters, rather than more evolved open clusters. Since the intrinsic near-infrared colors of stars on the upper main sequence vary by only a small amount, at the
depth of 2MASS the $H-K$ color serves as an indicator of extinction for all but the most nearby clusters (where the lower main sequence, and thus stars with differing intrinsic near-IR colors, can be observed). In addition, it is less likely that an overdensity due to chance superposition will have colors that differ from the background, so the rate of false cluster detections should be decreased by such a method. Finally, this different algorithm provides a complementary method of searching for embedded clusters; while the most prominent clusters with significant density enhancements will be found by any automated search, less obvious clusters may be missed by one or the other.

Thus, our criteria for designating a cluster candidate were: stellar density exceeding the locally determined background level by $5 \sigma$ or more, average $H-K$ color in the overdense region redder than the background, and that the color criterion holds when the single reddest source was excluded from the average. Without the final requirement, a single very red source superposed on an overdense region could result in designation of a region as a cluster candidate. In the targeted search phase, bins used for determining an overdensity were 30 "squares; bin size was variable between $15^{\prime \prime}$ and $45^{\prime \prime}$ in the full-sky phase to account for the variable stellar density in different parts of the sky. Once an overdense region was found meeting the criteria, the extent of the cluster candidate was mapped by determining the extent of the $3 \sigma$ overdense region. Only sources with good photometry in all three 2MASS bands (i.e. not saturated or filled in from the other bands based on a blackbody assumption) were included in the determination.

### 2.2.2 Targeted Search

Initially we conducted a targeted search around known star-formation regions in the Galaxy where embedded clusters would be expected. We selected the Sharpless catalog of HII regions and sources from the IRAS point source catalog with
colors consistent with star formation as described in (Carpenter et al., 1995). As an additional check, we added the clusters found by Dutra \& Bica (2001) in their initial search. For this initial study we downloaded the 2MASS Point Source Catalog in windows of $3^{\prime}$ in diameter around each target region and searched for density enhancements coincident with the target. Candidates selected by the algorithm were inspected visually to weed out chance superpositions and detector artifacts. In the course of our followup observations (Leistra et al., 2005) we found that several of our targets toward the Galactic Bulge were extinction features, rather than clusters; regions of lower extinction allowed Bulge stars to be observed that were completely obscured in regions of high extinction, producing a region with a higher stellar density and redder colors on average. While it may initially seem counterintuitive for a low-extinction hole to produce a region of redder stellar colors, this is the result when a region of stars with non-zero extinction can be observed through the "hole", while the extinction outside the hole is large enough that they are not just reddened but completely invisible to the limiting magnitude of 2MASS; in these regions outside the "hole", then, only the nonextincted, blue foreground stars are visible. Upon observation of a larger field than was used for the initial automated search, we determined that these extinction features had stellar density and color consistent with other nearby regions, and that they were not genuine stellar clusters. Thus, for subsequent studies we used a larger search radius to better determine the background stellar density and color and thus reduce the number of false-positive cluster candidates.

### 2.2.3 Full-Sky Search

Following the targeted search we conducted a search of the entire 2MASS Point Source Catalog, to perform an unbiased search for infrared-selected embedded stellar clusters. Due to the format in which the full catalog is available, slight ad-
justments to the algorithm were necessary. We divided the Point Source Catalog into rectangular chunks of 1000 sources (square regions provide more efficient tiling of the entire sky than circular regions) and searched each region for overdensities as determined in that region. Since the number of sources in each search region was constant, the actual size of the regions selected varied considerably between Galactic Plane and out-of-plane fields. The bin size was thus determined on a field-by-field basis. The actual criteria for selecting a cluster candidate were the same as for the targeted search, and the catalog was searched twice with the boundaries between sections offset in order to avoid missing clusters that fell on the edge of a search region.

Once a list of cluster candidates was obtained, the 2MASS images in all three bands were obtained for each candidate, and a color composite was inspected by eye to determine whether a genuine cluster was present. In many cases, especially in the case of large clusters, the same cluster would be detected multiple times. Since distinguishing between a single cluster with complex morphology or substructure and several related clusters is subjective and difficult, we count each detection as a "candidate" even if multiple detections correspond to the same portion of the sky. Out of 1376 candidates found by the algorithm, we determined by visual inspection and cross-reference with SIMBAD that 342 were genuine, previously known clusters, some found previously in our targeted search (including embedded and non-embedded clusters; 71 candidates corresponded to 19 globular clusters, since many clusters were detected in multiple chunks), two were previously unknown clusters not found by the targeted search, and 1025 were false detections (extinction features, chance superpositions lacking any indication of extended emission or indications of star formation at other wavelengths, detector artifacts, or non-cluster features such as galaxies). Seven candi-
dates remained ambiguous, and need deeper and higher-resolution imaging to resolve. The visually confirmed clusters (with multiple detections of the same complex or cluster combined into a single listing) are listed in Table 2.1.

Table 2.1. Coordinates and associated sources (previous identifications of the cluster, or any associated radio source) of cluster candidates recovered via our algorithm and confirmed by visual inspection. $a$ : Bica et al. (2003) b: Dutra et al. (2003b)

| RA | DEC | Associated with |
| :---: | :---: | :---: |
| 01912.5 | 655030.6 | IRAS 00165+6534 |
| 4568.5 | 472139.6 | Sh 219/IRAS 04523+4718 |
| 45828.3 | 475759.5 | Sh 217/IRAS 04547+4753 |
| 51059.3 | 37575.4 | IRAS 05075+3755 |
| 52243.6 | 332510.2 | NGC 1893 |
| 53124.6 | 341420.7 | NGC 1931 |
| 5355.7 | -5 279.8 | OMC 1 Cluster |
| 5399.9 | 354510.3 | Sh 2-233 |
| 5416.6 | 354924.0 | IRAS 05377+3548 / BDB2003 G173.63+02.81 |
| 5523.7 | 272413.2 | IRAS 05489+2723/ BDB2003 G182.05+00.42 |
| 6082.3 | 312231.4 | IRAS 06048+3123 |
| 6842.2 | 213347.3 | BDB G189.02+00.81\% IRAS 06056+2131 |
| 6842.2 | 214027.3 | BDB G188.94+00.88 / IRAS 06054+2141 |
| 6921.6 | 212321.7 | BDB G189.23+00.90/ IRAS 06065+2124 |
| 6927.9 | 245531.0 | IRAS 06063+2456/ BDB G186.13+02.59 |
| 6945.3 | 203352.8 | NGC 2175 |
| 6946.6 | 205444.0 | IRAS 06067+2055 / BDB G189.69+00.72 |
| 61127.8 | 172615.5 | Sh 2-259 / IRAS 06084+1727 |
| 61252.8 | 175916.3 | Sh2-255 / BDB2003 G192.59-00.05 |
| 61321.2 | 152355.8 | IRAS 06104+1524 / BDB2003 G194.93-01.20 |
| 61328.3 | 175533.1 | Sh2-258 / IRAS 06105+1756 / BDB2003 G192.72+00.03 |
| 61429.4 | 135050.3 | BDB2003 G196.45-01.67 / Sh 269 |
| 61445.9 | 19032.3 | IRAS 06120+1903 / BDB2003 G191.92+00.82 |
| 61550.9 | 14161.7 | IRAS 06127+1418 / DBS2003 81 ${ }^{\text {b }}$ |
| 61641.8 | 223334.9 | IC 443 |

Table 2.1—Continued

| RA | DEC | Associated with |
| :---: | :---: | :---: |
| 61845.1 | 151633.2 | IRAS 06159+1514 / BDB 2003 G195.65-00.10 |
| 65934.3 | -4468.2 | Sh 2-287 / IRAS 06571-0441 / BDB2003 G218.02-00.32 |
| 73157.2 | -165655.4 | DBS2003 4/ IRAS 07298-1648 |
| 73544.4 | -18494.3 | DBS 2003 8/ IRAS 07334-1842 |
| 7512.5 | -121819.5 | SH 2-297 / DBS2003 96 |
| 7532.6 | -111350.9 | IRAS 07032-1105 |
| 7838.0 | -41920.4 | BDB G218.74+01.85 / SH 2-288 |
| 84020.7 | -455113.0 | IRAS 08389-4545 |
| 84417.6 | -454755.0 | IRAS 08389-4533 |
| 104546.5 | -60013.2 | Trumpler 16 / IRAS 10441-5949 |
| 121221.6 | -625730.4 | IRAS 12100-6242 |
| 13133.4 | -625959.0 | IRAS 13098-6244 |
| 131339.3 | -621920.0 | IRAS 13101-6200 / DBS 2003 133 |
| 132959.5 | -613543.1 | IRAS 13268-6114 |
| 133218.8 | -624021.0 | Trumpler 21 / IRAS 13286-6225 |
| 13540.7 | -615337.0 | NGC 5316 |
| 135828.0 | -61429.0 | Loden 1101/IRAS 13552-6129 |
| 13749.3 | -62405.0 | IRAS 13050-6218 |
| 144840.2 | -595525.1 | IRAS 14450-5944 |
| 153956.4 | -555559.0 | IRAS 15360-5543 |
| 15401.1 | -535825.0 | IRAS 15362-5351 |
| 155326.2 | -544510.0 | IRAS 15496-5434 |
| 155436.5 | -535117.0 | IRAS 16078-5131 2003 148/ IRAS 15507-5341 |
| 161147.7 | -514044.0 | IRAS 16082-5150 |
| 161211.7 | -515729.0 |  |

Table 2.1-Continued

| RA | DEC | Associated with |
| :---: | :---: | :---: |
| 161340.0 | -5120 41.6 | IRAS 16097-5109 |
| 161940.8 | -505648.0 | IRAS 16159-5049 |
| 162053.2 | -5059 46.0 | IRAS 16170-5043 |
| 164650.5 | -114629.9 | RAFGL 2683 |
| 165921.1 | -42 3451.1 | DBS2003 176 / IRAS 16558-4228 |
| 16595.8 | -42 4225.8 | IRAS 16556-4235 |
| 170416 | -4133 33 | IRAS 17010-4129 |
| 172017.7 | -35 5416.2 | BDB 2003 G351.23+00.67 |
| 172334.6 | -35 5525.1 | BDB 2003 G351.61+00.17/ IRAS 17200-3550 |
| 17235.0 | -35 5521.3 | IRAS 17197-3552 |
| 172532 | -34 2351 | NGC 6357a |
| 172603 | -341639 | GRS 353.3+00.60 |
| 172607 | -341853 | IRAS 17229-3418 |
| 172927.6 | -343751.1 | IRAS 17262-3435 |
| 174458.3 | -29 4432.0 | IRAS 17417-2940 / BDB 2003 G359.28-00.25 |
| 174540.8 | -29 045.9 | Galactic Center |
| 174614.2 | -2850 27.3 | Quintuplet / BDB2003 |
| 181845.9 | $-162532.0$ | IRAS 18161-1626 |
| 183922 | -5 5330 | IRAS 18367-0556 |
| 184413.9 | -4 1812.4 | DBS 2003 123/ IRAS 18416-0421 |
| 18463.7 | -2 3923.1 | IRAS 18434-0242 |
| 18858.2 | -2055.8 | DBS 2003 113/ IRAS 18060-2005 |
| 191327.5 | 105317.7 | DBS 2003135 |
| 192214.3 | 14348.9 | IRAS 19202+1359 |
| 192342.2 | 143039.5 | W 51/ DBS 150 |

Table 2.1—Continued

| RA | DEC | Associated with |
| :---: | :---: | :---: |
| 192429.8 | 20478.3 | DBS 13 / Sh 2-83 |
| 195843.1 | 312025.3 | Sh 2-98 |
| 20142.1 | 333447.0 | IRAS 19597+3327a/ BDB G070.30+01.59 / W58 |
| 201756.4 | 364528.9 | IRAS 20160+3636/ Sh 2-104 |
| 202141.5 | 37266.1 | DB 2001 CL 5 / IRAS 20198+3716 |
| 202227.3 | 401930.1 | Sh 2-108 |
| 202234.0 | 401150.1 | DWB 63 |
| 202723.6 | 37229.1 | Sh 106/ BDB G076.37-00.61 |
| 202726.1 | 372252.2 | Sh 106/ LK 2002 Cl 02 |
| 20278.5 | 39283.9 | BDB 2003 G078.04+00.62 |
| 202935.7 | 39246.8 | BDB2003 G077.96-00.01 |
| 203032.7 | 411523.2 | BDB2003 G079.87+01.18/ IRAS 20286+4105 |
| 203143.8 | 385621.0 | IRAS 20300+3847/ BDB 2003 G078.16-00.37 |
| 203221.4 | 43417.0 | BDB2003 G082.04+02.33 |
| 203227.4 | 385123.7 | BDB2003/ IRAS 20306+3841 |
| 203228.1 | 401714.9 | IRAS 20306+4005/ BDB2003 G079.30+00.29 |
| 203522.4 | 422134.6 | BDB 2003 G081.31+1.10/ W72 |
| 203832.4 | 42810.9 | BDB 2003 G081.44+00.48 |
| 203837.5 | 423925.9 | BDB 2003 G082.00+00.80 / W75n |
| 203856.8 | 422248.3 | BDB 2003 G081.71+00.58 / W75s |
| 203925.2 | 411945.4 | IRAS 20375+4109/ DB2001 Cl 14 |
| 204188.9 | 415159.1 | BDB 2003 G081.57-00.07 / BDB 2003 G081.66-00.02 |
| 204237.6 | 425930.6 | BDB 2003 |
| 20453.3 | 291129.1 | BDB 2003 G066.96-01.28 / IRAS 20028+2903 |
| 204536.5 | 441521.7 | BDB 2003 G083.94+00.78 |
| 20 |  |  |

Table 2.1—Continued

| RA | DEC | Associated with |
| :---: | :---: | :---: |
| 20455.0 | 291357.0 | IRAS 20027+2905 / Roslund 4 |
| 223244.2 | 58287.9 | Sh2-138/ BDB 2003 G105.62+00.34 |
| 224929.7 | 595445.8 | DBS2003 34 / IRAS 22475+5939 |
| 225651.9 | 623958.5 | DBS2003 40 |
| 22573.9 | 623818.5 | DBS2003 41 |
| 231344.8 | 61285.1 | NGC 753 |

The false positive rate of this non-targeted search was quite high, which was unsurprising; the threshold was set fairly low to avoid missing genuine clusters, but including the entire sky (including regions far out of the Galactic Plane where we would expect few real clusters) increases the background without substantially increasing the number of genuine clusters.

### 2.3 Analysis

### 2.3.1 False Cluster Detections

We expected that the rate of false-positive cluster detection in the full-sky search would be relatively high. The bins used to determine stellar density vary in size, averaging $30^{\prime \prime}$ on a side, so precisely calculating the number of false-positive detections we expect is difficult; with bins of $1^{\prime}$ square, we would expect approximately $1305 \sigma$ detections from statistical fluctuations alone. Since our bins vary in size, averaging $30^{\prime \prime}$ on a side, we would expect roughly four times this number if a purely density-based threshold were employed. The color selection should cut the number in half (since we require only that the overdensity be redder than the
surroundings). Despite the large number of false cluster detections this threshold will produce, we chose this threshold to reduce the number of missed cluster detections, based on the significance of detections in the targeted search and at the location of known embedded clusters. False detections due to statistical fluctuations alone can be easily distinguished by visual inspection of the 2MASS images. Statistical fluctuations, generally in the form of an overdensity in a generally sparse field (such that two or three extra stars in the bin represent a $5 \sigma$ overdensity) account for 943 out of the 1025 total false detections, a large majority and consistent with expectations. This type of false detection will be an issue for any full-sky automated search, and emphasizes the need for visual inspection of the results rather than relying purely on the automated method. Many of the detections claimed by (Dutra \& Bica, 2000) could have been excluded based purely on a visual inspection of this sort. Some of these detections may be genuine stellar multiplets or sparse groupings of $N<10$ stars; such groupings cannot be confidently identified based solely on the 2MASS imaging data, so we do not consider them. We show a typical example of the sparse-field statistical fluctuation detection in Figure 2.1.

The rest of the false cluster detections are of two types: extinction features, as described in Section 2.2.2, which can again generally be recognized on visual inspection of a sufficiently large image, and galaxies or (in a few cases) groups of galaxies. These account for 27 and 55 detections respectively. Figures 2.2 and 2.3 show typical examples of each of these three failure modes of the algorithm.

### 2.3.2 Missed Clusters \& Completeness of the Sample

False negatives, when the algorithm fails to detect known clusters, are at least as important as false positives. False positives can be weeded out by visual inspection, but false negatives reflect on the completeness of the final sample. To


Figure 2.1: 2MASS $K_{s}$ image of an "extinction hole" detected as a cluster candidate by the automated search.


Figure 2.2: 2MASS $K_{s}$ image of a sparse-field overdensity detected as a cluster candidate by the automated search, where a small number of extra stars triggered a detection.


Figure 2.3: 2MASS $K_{s}$ image of a galaxy detected as a cluster candidate by the automated search.
test the reliability of our algorithm, we looked at the sample of known embedded clusters compiled by Lada \& Lada (2003), containing 76 clusters within $\sim$ 2.5 kpc of the Sun. We expected that both the most nearby and the most distant clusters would be missed by our algorithm; the most nearby because they would fill an entire search region, so that the "background" density would reflect the cluster, and the most distant because the stars would be sufficiently faint and crowded as to be indistinguishable as point sources at the resolution and sensitivity of the 2MASS data. Of the 76 embedded clusters tabulated, we recovered 29 and missed 57. Based on our visual inspection of the 2MASS images, nineteen of those were missed for the reasons we anticipated; either they were spread out enough that they entirely filled the search region, too faint for 2MASS to detect enough point sources, or compact enough that they were unresolved by 2MASS; additionally, three had enough nebulosity that the number of resolved point sources was small. We would expect the nearby and highly nebulous clusters to be well-recovered by an entirely visually-based search. Six were in regions of high background stellar density and one was near a bright open cluster, which affected the background density calculations; four were in a region with red background stars. We found no clear correlation of our success rate with the number of cluster members, and a correlation with distance; we failed to recover clusters nearer than 400 pc (with the exception of Rho Oph). Since the most distant cluster in the Lada \& Lada (2003) sample were at a distance comparable to our outer limit we found no decline in the success rate at the larger distances in this sample. The nearby clusters were missed because of field-filling issues, while the more distant were missed because of crowding and faintness. We could find no obvious reason for the remaining eleven non-detections. These final objects (the eleven with no obvious reason for the non-detection, and the eleven with failure
due to the environment rather than the characteristics of the cluster itself) represent our incompleteness; we thus expect that our sample is approximately $50 \%$ complete in the regime accessible by 2MASS. The total completeness from recent 2MASS searches in this region is somewhat higher, as we recover only approximately $50 \%$ of the clusters found by Dutra et al. (2003b). Of the clusters we did not recover we find that $\sim 50 \%$ would pass our visual inspection; considering all recent searches of the 2 MASS data together, we expect the census of clusters within 2 kpc is $\sim 75 \%$ complete. This is an improvement over the Lada \& Lada (2003) compilation, which predates most of the 2MASS searches.

Based on the results for known clusters, we conclude that our algorithm is reliable for detecting embedded clusters only at distances $\lesssim 1.5 \mathrm{kpc}$, and not in regions of high stellar density (e.g. toward the Galactic Bulge); while some clusters can reliably be detected in such regions (e.g. the G353-0.4 cluster (Leistra et al., 2005), detectable in large part because its natal molecular cloud obscures the background stars in the immediate vicinity), the completeness in such regions is low.

### 2.3.3 Embedded Cluster Candidates

The searches (targeted and full-sky) detected a total of 339 visually confirmed embedded cluster candidates. 206 are well-known, well-studied clusters, 78 had been found by previous 2MASS searches (but had not been previously known), and 55 were new. 118 cluster candidates from Dutra \& Bica (2000) were not recovered by the automated search. We performed a similar analysis on these clusters, after excluding objects which they classified as groups (containing sufficiently few stars that we would have excluded them based on visual inspection) or as "cluster candidates", where their identification was less confident. Out of those which were rejected by our algorithm, $57 \%$ would have passed our visual inspec-


Figure 2.4: Distribution of embedded cluster candidates in galactic coordinates.
tion, but either lacked color contrast or consisted primarily of nebular emission and lacked a significant overdensity of point sources. The purely visual method appears to be somewhat better at finding sparse or very young clusters with few point sources, but the degree of non-overlap suggests that the two approaches are complementary and neither is clearly superior.

We plot the distribution of all of our embedded cluster candidates in Galactic coordinates in Figure 2.4. The majority of the clusters recovered are confined to the Galactic Plane, with a widening in the distribution apparent toward the Bulge and a significant overdensity of clusters at a Galactic longitude of $\sim 80$, toward Cygnus. Two clusters were found at Galactic latutides with $|b|>10$. These are the OMC-1 cluster $(b=-19.45)$ and a cluster associated with IRAS 16170-5043 ( $b=20.95$ ).

We expected that clusters recovered in regions of high field star density would
tend to be more compact, and that large, more spread-out clusters would be found primarily in regions of lower field-star density; in regions of high fieldstar density, a large and diffuse cluster would be spread out over multiple search regions, while in regions of very low field-star density the 1000-star search region can span tens of arcminutes. We plot the major axis of our cluster candidates (as determined by visual examination of the 2MASS images) against galactic longitude in Figure 2.5, broken down by galactic latitude. Such a correlation may be present, with slightly larger clusters found preferentially further from the Galactic Plane or away from the Galactic Bulge, but it is by no means complete. Smaller-diameter clusters are found everywhere, and occasional large clusters are found in regions of higher average field star density. The complex morphology of clusters and the presence of dust can explain these findings; dust can obscure local background stars entirely (as in the case of the G353-0.2 cluster discussed in Chapter 3) such that large-diameter clusters can be found even in regions where, considered on larger scales, the field star density is high.

### 2.4 Limitations of Automated Searches and Implications for Future Studies

While careful selection of criteria for an automated cluster search can improve the success rate, any set of simple color and density parameters that will find all genuine clusters will include numerous false positives as well. Some of the problems we discuss can be addressed by a more sophisticated algorithm (e.g. dealing with "extinction features" by requiring the cluster candidate to be denser than all regions of comparable size in the surrounding field, rather than just denser than the average, which would also exclude some "sparse-field" statistical fluctuations, or restricting searches to the Galactic Plane, ruling out most galaxies and statistical fluctuations and excluding only the most nearby clusters, which are unlikely to


Figure 2.5: Major axis (in arcmin) of cluster candidates as determined from 2MASS images. Circles: $|b|<0.5$ Crosses: $0.5<|b|<1.0$ Squares: $|b|>1$
have gone undetected in the optical). However, the presence of diffuse emission from nebular gas is a key indicator in visual determinations and is not reflected in the source catalogs used for automated searches, so that it cannot be easily included in an automated search.

Unfortunately, embedded clusters and the most likely sources of contamination and confusion both lie in the Galactic Plane. Even when candidates are reliably detected, determination of membership is difficult, requiring spectroscopic data.

An intriguing possibility for improving the success rate of automated searches is to use color information more fully than we have done, by searching for overdensities in CMD space rather than in the sky; that is, searching for enhancements around an isochrone placed at some distance. If such an enhancement is found with spatial correlation among the members, it would be a good candidate for a genuine cluster.

### 2.5 Followup Observations of Cluster Candidates: False Detections

Since our initial targeted search was based on a $3^{\prime} \times 3^{\prime}$ window, some of our candidates for followup imaging proved upon examination of the larger $8^{\prime} \times 8^{\prime}$ FOV of the IRIS2 camera used for followup imaging (see Leistra et al. (2005) for the results of the successful cluster candidate observations) to be the extinction-hole false detections described above; their stellar density was not noticeably different from that of the denser portions of the field, and in one case there was no nebulosity coincident with the cluster candidate. We obtained JHK images of these cluster candidates to further investigate their natures.

After reducing the data as described in Leistra et al. (2005), I defined a candidate region and a field region by eye. The candidate region was chosen to
correspond to the dense portion of the cluster candidate, as for the confirmed clusters, while the background region was chosen to be as large as possible while avoiding dust lanes. Selected regions for both candidates are shown in Figures 2.6 and 2.7. As described in Leistra et al. (2005) we performed aperture photometry using IRAF on both images. Despite the crowded fields, aperture photometry was found to give smaller photometric errors than PSF-fitting photometry due to the undersampled PSF. Systematic photometric errors, which would result in an overall miscalibration of the source magnitudes, will not affect the relative number of source counts or measured magnitudes between the candidate and field regions, and crowding characteristics of the two regions are similar.

### 2.5.1 DB11 Cluster Candidate

The $K$ vs. $H-K$ CMD for the full "field region" corresponding to the DB11 cluster candidate is shown in Figure 2.8. Two sequences can be seen at different $H-K$ colors, with a clear separation between the two. These correspond to upper main-sequence stars in the foreground and to upper main-sequence stars at a higher extinction, presumably lying in the Galactic Bulge. In the region of still higher extinction immediately surrounding the DB11 candidate, only the bluer foreground sources are visible, while the more extincted Bulge sources are visible in the surroundings. It appears that the cluster candidate is a region of low extinction where the Bulge stars are again visible; the CMD within the cluster candidate region, shown in Figure 2.9, show a similar distribution.

1516 sources were detected in $K$ to a limiting magnitude of $K=16.5$ in the field region, and 158 in the cluster region. Based purely on the ratio of the areas we would expect 147 stars in the cluster region. The "cluster candidate" thus represents an overdensity of only 11 stars out of 158 , or approximately $1 \sigma$. By contrast the overdensity toward the cluster associated with G305-0.2, with the


Figure 2.6: $8^{\prime} \times 8^{\prime}$ K-band image of the DB11 cluster candidate from IRIS2/AAT.
The regions used for "candidate" and "field" photometry are indicated.


Figure 2.7: $8^{\prime} \times 8^{\prime}$ K-band image of the DB41 cluster candidate from IRIS2/AAT.
The regions used for "candidate" and "field" photometry are indicated.


Figure 2.8: All sources detected in the "field region" for the DB11 cluster candidate. Sources with $K>10.5$, which suffer from saturation, are plotted but were excluded from all analysis.


Figure 2.9: All sources detected in the "cluster region" for the DB11 cluster candidate. Sources with $K>10.5$, which suffer from saturation, are plotted but were excluded from all analysis.


Figure 2.10: Statistically-corrected "cluster region" sources for the DB11 cluster candidate.
highest field star density of any of our confirmed clusters, was 115 stars out of 139 (approximately 24 field star interlopers).

We performed the same statistical correction procedure on the DB11 cluster candidate as on the confirmed clusters described in Leistra et al. (2005), where stars were randomly selected for removal from the cluster region based on the distribution of field stars in the $K$ vs $H-K$ CMD. The resulting CMD is shown in Figure 2.10. No trend is apparent in the remaining sources. In the genuine clusters, the sources remaining after statistical correction for field contamination tended to be redder and brighter than the sources that were removed.

We conclude that the cluster candidate designated as Cluster 11 by Dutra \&

Bica (2000) is not an actual embedded stellar cluster, but is most likely an extinction feature or a statistical overdensity with no physical connection between the stars. The automated search ,and the inspection of the small surrounding field, picked it out because it is locally surrounded by a region of high extinction, making the "field" density in its immediate surroundings artificially low. However, the overdensity relative to the less heavily extincted field a few arcminutes away is statistically insignificant, and no color difference is discernible.

We note that there is a small, but apparently genuine embedded cluster (designated Cluster 10 by Dutra \& Bica (2000)) in the field, near the edge of our IRIS2 images. There is significant nebulosity associated with this cluster, but too few stars to calculate a meaningful IMF.

### 2.5.2 DB41 Cluster Candidate

In this case the "cluster region" is less clearly well-defined than in the case of the DB11 cluster candidate; while there is a sharp edge coinciding with some nebulosity to the southwest at the edge of a dust lane, the separation is less clear on the other edge of the region. We chose to define the candidate region to coincide with the extent of the nebulosity. We repeated the procedures described above for the DB11 cluster candidate. 56 sources were detected in $K$ in in the candidate region to a limiting magnitude of 16.5 , compared with 49 expected in the same area of the field region. This represents an overdensity of 7 sources, about $1 \sigma$. With the total image area covering approximately 150 times the area of the cluster candidate, we would expect such an overdensity purely from statistical fluctuations. However, the coinciding nebular emission suggests that there may be a genuine physical association between the stars in this situation.

The CMDs for cluster candidate and field are shown in Figures 2.11 and 2.12. The two CMDs span the same range of $H-K$ color and $K$ magnitude. The cluster


Figure 2.11: All sources detected in the "field region" for the DB41 cluster candidate.
candidate region may show an enhancement of relatively bright, blue stars (four stars with $H-K \sim 0, K \sim 12$; the color of these sources suggests that, while they may be part of a small open cluster, they are not in an embedded cluster. This could, however, simply be due to the random selection of stars from the field used for this comparison, which were not weighted by the abundance in the full-field CMD.

To better quantify this apparent discrepancy, we performed the statistical correction on the DB41 cluster candidate as described above, removing stars from the cluster candidate region based on their frequency in the field. The statisti-


Figure 2.12: All sources detected in the "field region" for the DB41 cluster candidate.


Figure 2.13: Statistically-corrected "cluster region" sources for the DB41 cluster candidate.
cally corrected cluster-candidate CMD is shown in Figure 2.13. This leaves more stars than the actual cluster overdensity, but the pattern of removal appears to be mostly random, in that the remaining sources do not disproportionally inhabit a particular region in the CMD.

We cannot determine from our photometric data whether candidate DB41 is a genuine cluster. The field star contamination is in any case too high to allow analysis of this object as a cluster. Portegies Zwart et al. (2001) discuss the issue of unrecognizable clusters toward the Galactic Center, where the high field star
density can render even genuine clusters unfindable. In this case the presence of nebulosity suggests that there may be star formation somewhere nearby, but does not unambiguously identify DB41 as a cluster. We conclude that this object remains ambiguous, as either a small cluster or a statistical fluctuation.

### 2.6 Finding Embedded Clusters: Conclusions

We found that, for a targeted search of known star-formation regions, the success rate of a near-IR search for embedded clusters is improved by the use of color selection, but that human intervention (visual inspection of cluster candidates) is still required to eliminate chance superpositions and extinction features. This suggests that ongoing searches for embedded clusters in Spitzer data (e.g. Mercer et al., 2005) will also benefit from visual inspection of the images surrounding their candidates.

Our full-sky survey using the same algorithm we used for the targeted search produced no major surprises; the star-forming regions that produce embedded clusters (especially those nearby enough to be recognized in 2MASS) are also detected in the far-IR, so the targeted search of IRAS sources recovered the majority of clusters. The full-sky survey confirmed this with only 2 embedded clusters detected that had not been found by the targeted search.

Most embedded-cluster surveys in the Galactic Plane report a list of newly discovered clusters; in the absence of considering both false detections and wellknown clusters that were missed by their method, these are less useful than they could be; while the individual targets may be interesting for studies of star formation, analysis of the statistical properties of the embedded cluster population can only be done if the completeness of the sample is understood and contaminants weeded out.

## CHAPTER 3

Southern 2MASS-Selected Young Stellar Clusters: Photometry, Spectroscopy, and the Initial Mass Function

The contents of this chapter were previously published in Leistra et al. (2005).

### 3.1 INTRODUCTION

Despite their intrinsic rarity and short lifetimes, massive stars are extremely important in the evolution of galaxies. They play an important role in determining the course of the formation of less massive stars, though the nature of this role is still uncertain, and their stellar winds and eventual supernovae shape the interstellar medium. They produce most of the heavy elements in the universe, as well as much of the UV radiation in galaxies. Their rarity, combined with the effects of large Galactic extinctions, often results in the availability of more comprehensive studies of massive stars in external galaxies, where the entire stellar population can be observed at once, than within our own where massive stars must be studied individually and the census of massive stars is still very incomplete. High optical extinction within the galactic plane $\left(A_{V} \gtrsim 20\right)$ has limited optical studies of massive stars to relatively nearby regions ( $\mathrm{R}_{\text {solar }} \lesssim 3.0 \mathrm{kpc}$, Massey, 2003). Even within that radius, optically selected catalogs of $O$ stars have been found to be incomplete, especially in star-forming regions and young clusters (e.g. Hanson \& Conti, 1995). This incompleteness necessitates the use of infrared, radio and X-ray observations, particularly in the inner regions of the Galaxy and in star formation regions. The near-infrared (NIR, 1-5 $\mu \mathrm{m}$ ) is an especially useful regime for the study of massive stars; the stellar atmosphere is still observed directly, but since for example $A_{K} \simeq 0.11 A_{V}$, we can observe these stars in regions where
dust, either along the line of sight or local to the star-forming region, makes them inaccessible at optical wavelengths. The discovery and characterization of stellar clusters observable only in the infrared can significantly enhance our understanding of obscured Galactic regions which harbor embedded massive stars or massive protostars.

Recent studies indicate that clusters may account for 70-90\% of star formation and that embedded clusters (those still partially or fully enshrouded in their natal molecular cloud) may exceed the number of more traditional open clusters by a factor of $\sim 20$ (Elmegreen et al., 2000; Lada \& Lada, 2003). In the last decade, advancements in NIR observational capabilities resulted in the discovery and classification of some of the most massive young stellar clusters in the Galaxy, each containing dozens of O and WR stars (e.g. Nagata et al., 1995; Cotera et al., 1996; Figer, Morris, \& McLean, 1996). Recent studies (Figer et al., 1999) have suggested that within these clusters, the initial mass function (IMF) does not follow the canonical Salpeter form with a slope $\Gamma=-1.35$, but instead is more heavily weighted toward massive stars; mass segregation has been proposed as a solution (Stolte et al., 2002). In the last several years a number of studies of well-known star formation regions have also been carried out in the NIR, (e.g Okumura et al., 2000; Blum, Damineli, \& Conti, 2001; Conti \& Blum, 2002; Figuerêdo et al., 2002). These studies have in most cases found an IMF consistent with the Salpeter value, and have uncovered candidate massive YSOs. In addition, within the past ten years, massive YSOs within molecular clouds have been studied in the NIR, (e.g. Chakraborty et al., 2000; Ishii et al., 2001) and in young stellar clusters (e.g. Hanson, Hayworth, \& Conti, 1997). Massive YSOs, however, remain significantly less studied and are poorly understood in comparison with their lower-mass counterparts; many more must be identified and studied before we can adequately
address how the formation of massive stars differs from that of low-mass stars.
The final release of the Two Micron All Sky Survey (2MASS) has fostered studies which can probe the entire Galaxy for previously unknown stellar clusters. Initial attempts were made which searched for stellar density enhancements, (e.g. Dutra \& Bica, 2000, 2001; Dutra et al., 2003), but the identification of previously unknown clusters has met with limited success. For example, Dutra \& Bica (2000) identified 52 candidate clusters, which subsequent observations (Dutra et al., 2003) indicated were in fact 10 confirmed clusters, 3 "probable" clusters, and 11 "dissolving cluster candidates"; the remainder were not clusters. Our observations of at least one of the Dutra et al. (2003) "confirmed clusters", however, indicates that the "cluster" is most likely a region of low extinction rather than a true cluster (Cotera \& Leistra, 2005). We have performed an independent search of the 2MASS archive, using color criteria in addition to stellar density enhancements. We have searched in the vicinity of regions identified as likely sites of star formation based on radio and IRAS far-infrared flux ratios, and are currently conducting a search of the entire 2MASS Point Source Catalog. We search the Point Source Catalog for regions of higher stellar density than the background (determined locally within a $5^{\prime}$ radius) which are redder in $H-K$ than the local field. This selects for embedded clusters, with the color criteria helping to eliminate chance superpositions and regions of low extinction. In contrast, Dutra \& Bica $(2000,2001)$ use only stellar density to select clusters. Our method has been relatively successful to date; correctly selecting 7 clusters out of 9 potential targets, including 4 candidates toward the inner Galaxy. We present NIR imaging and spectroscopy of the two confirmed clusters in the inner Galaxy in this paper, and discuss the two unconfirmed targets in detail in Cotera \& Leistra (2005). The cluster near G305.3+0.2 was independently discovered by Dutra et al. (2003b).

The additional 5 outer-galaxy targets are described in Paper II.
NIR imaging and spectroscopy of both young stellar clusters and nascent stellar clusters enables us to expand the study of the IMF in objects where there has been little to no stellar evolution off the main sequence or cluster evaporation, and where the cluster age can be constrained to within $\sim 2$ Myr. Spectral typing of the most massive stars in the cluster allows their masses to be determined relatively precisely, and when combined with photometry it facilitates a reliable determination of the masses of stars throughout the entire cluster (Massey, Johnson, \& DeGioia-Eastwood 1995; Massey 2002), allowing the initial mass function of the cluster to be determined more accurately than photometry alone would permit. In this paper we present the results of NIR observations of two clusters found toward the inner Galaxy, which we designate by the Galactic coordinates of their centers, G353.4-0.36 (17:30:28-34:41:36 J2000) and G305.3+0.2 (13:11:39.6 -62:33:13 J2000). In Paper II we will present the results of similar observations of five clusters in the outer Galaxy.

In $\S 3.2$ we present the observations and data reduction, in $\S 3.3$ we present the spectra and classifications of the spectroscopically observed cluster members as well as the color-magnitude diagrams, and in $\S 3.4$ we describe the luminosity function and the initial mass function.

### 3.2 OBSERVATIONS \& DATA REDUCTION

We observed candidate young stellar clusters with the facility instrument IRIS2 on the 3.9 m Anglo-Australian Telescope (AAT) on July 12-15, 2003. IRIS2 is an imaging spectrometer which uses a $1024 \times 1024$ Rockwell HAWAII- 1 HgCdTe array with a platescale of $0^{\prime} .45$ / pixel, resulting in a $7^{\prime} .7 \times 7^{\prime} .7$ field of view. Images were obtained in $J(1.25 \mu \mathrm{~m}), H(1.63 \mu \mathrm{~m})$, and $K_{s}(2.14 \mu \mathrm{~m})$ filters. $R \simeq 2300$
spectra of selected stars in each cluster candidate were obtained in $K$ for all candidates.

We selected a total of four cluster candidates in the southern hemisphere using the 2MASS Point Source Catalog based on color and density criteria. Two of the candidates observed appear to be regions of low extinction and are discussed elsewhere (Cotera \& Leistra, 2005). The two confirmed clusters are near radio H II regions designated G305.3+00.2 and G353.4-0.4. We present three-color composites of the $8^{\prime} \times 8^{\prime}$ images of the G305.3+00.2 and G353.4-0.36 clusters in Figures 3.1 and 3.2 respectively. G305.3+00.2 is an H II region which has been previously observed using radio recombination lines (Wilson \& Mezger, 1970), C I emission in the submillimeter (Huang et al., 1999), and in the mid-infrared (MIR) by the Midcourse Space Experiment (MSX). The kinematic distance of $3.5 \pm 1.1 \mathrm{kpc}$ obtained for this H II region (Wilson \& Mezger, 1970) agrees well with the distance of 3.3 kpc for masers several arcminutes away (Caswell et al., 1995), suggesting they may be part of a single star-formation complex. A distance of 4 kpc is adopted as an upper limit to the radio kinematic distance by Clark \& Porter (2004) in a study of the star clusters Danks 1 and 2 in this region. The situation is more complex for the G353.4-0.36 cluster, which is in a region known to be a site of massive star formation. There are numerous radio sources located within $1^{\prime}$ of the NIR cluster, which we discuss in detail in §3.3.2.2.

All photometric observations were done in excellent seeing conditions: $0^{\prime \prime} .7$ $0^{\prime \prime} 9$. The images were reduced and combined automatically at the telescope using the ORAC-DR pipeline. ORAC-DR is a generic data reduction pipeline created at the Joint Astronomy Centre in Hawaii, originally for use with various UKIRT and JCMT instruments. Subsequent reprocessing did not noticeably improve the images, therefore the pipeline processed data has been used throughout. Source


Figure 3.1: Color composite ( $J=$ blue, $H=$ green, $K=$ red) of the region around the G305+00.2 cluster. Image is approximately $8^{\prime}$ on a side. The cluster is clearly apparent as a concentration of stars with similar colors; no nebular emission is apparent in the immediate vicinity of the cluster though a ridge of nebulosity is present to the northwest.


Figure 3.2: Color composite ( $\mathrm{J}=\mathrm{blue}, \mathrm{H}=$ green, $\mathrm{K}=\mathrm{red}$ ) of the region around the G353.4-0.36 cluster. Image is approximately $8^{\prime}$ on a side. The cluster is surrounded by intense nebular emission and is contained in a larger dark molecular cloud.
detection, PSF fitting, and photometry was carried out using IRAF-DAOPHOT, and is discussed in detail in §3.3.2.

All spectra were obtained with a $1^{\prime \prime} \times 7^{\prime} 7$ slit. The long-slit format combined with the high stellar density within the FOV resulted in the simultaneous observation of multiple stars. Total integration times ranged from 10 minutes to 30 minutes, and were chosen to provide adequate $\mathrm{S} / \mathrm{N}$ for NIR spectral classification as described in Hanson, Conti, \& Rieke (1996). After the data was flat-fielded, grism curvature was removed using the FIGARO ${ }^{1}$ tasks $c d i s t$ and sdist. Wavelength calibration was performed using the the $\mathrm{OH}^{-}$night sky lines and the FIGARO task arc. The uncertainty in the wavelength calibration fit was determined to be $2.18 \AA$. The FIGARO task irflux was used both to flux-calibrate the spectra and remove the telluric absorption using the G2V standards HD157017 and HD115496. Both of the standards had intrinsic $\mathrm{Br} \gamma$ in absorption, with equivalent widths of $5.7 \AA$ for HD157017 and of $5.6 \AA$ for HD115496; in each case, the absorption line was removed by fitting a line to the continuum in the region of the line in the standard star spectrum prior to flux calibration. The individual spectra were obtained by extracting apertures 4-5 pixels wide from the full spectral array, then performing background subtraction using apertures of the same width on either side of the source, separated by 2 pixels ( $0^{\prime \prime} 9$ ). We also extracted off-source spectra in each cluster to characterize any nebular emission.

### 3.3 Analysis

### 3.3.1 Spectroscopy

The development of NIR spectral atlases of nearby massive stars of known spectral type (Hanson et al. 1996; Morris \& Serabyn 1996; Blum et al. 1997), provides a

[^0]valuable classification scheme for stars too heavily obscured by dust to permit optical spectroscopy. In the $K$ band, in addition to the $\operatorname{Br} \gamma(2.165 \mu \mathrm{~m})$ line, massive O stars have helium (He I $2.058 \mu \mathrm{~m}$, He I $2.112 \mu \mathrm{~m}$, He II $2.189 \mu \mathrm{~m}$ ), carbon (C IV $2.078 \mu \mathrm{~m}$ ), and nitrogen ( N III $2.116 \mu \mathrm{~m}$ ) lines in their spectra which allow for the determination of the spectral type to within a subtype if there is adequate ( $\gtrsim 70$ ) line signal to noise. Table 6 of Hanson et al. (1996) indicates that in many cases the mere presence of these lines in emission or absorption (without considering equivalent width) is sufficient to determine spectral type to within two subtypes for O stars. The situation is more complicated for B stars, which have fewer features in this part of the spectrum; however, they are still classifiable using only $K$-band spectra.

We obtained $K$-band spectra of five stars in the G305.3+0.2 cluster field and three stars in the G353.4-0.36 cluster. In order to reduce the level of foreground contamination, we imposed a color cut of $H-K>0.5$ based on the 2MASS magnitudes and selected the brightest stars meeting this requirement. Despite this cutoff, two of the five stars observed in the G305.3+0.2 cluster proved to be foreground contaminants with sufficient line-of-sight extinction to push them over our threshold. The cluster sequence was much narrower and more wellseparated from the foreground in the G353.4-0.36 cluster, and no obvious foreground contaminants were present in our spectroscopic sample. The G353.4-0.36 cluster was sufficiently red ( $H-K_{\text {cluster }} \gtrsim 1.3$ ), that the time required to obtain a useful signal-to-noise in $H$-band spectra would have been prohibitively large, so only $K$-band were obtained.

### 3.3.1.1 G305.3+0.2 Cluster

We present spectra for the three cluster members, which we label A1-A3, in Figure 3.3. In Figure 3.4 we present a $106^{\prime \prime} \times 120^{\prime \prime}$ image of the cluster and label

Table 3.1. Magnitudes and spectral line identifications of sources in G305.3+0.2 for which we obtained spectra. Numbers in parentheses in the photometry are the errors in the least significant digit.

| Star | $\begin{gathered} \text { RA } \\ 13^{\mathrm{h}} 11^{\mathrm{m}} \end{gathered}$ | $\begin{aligned} & \text { DEC } \\ & -62^{\circ} \end{aligned}$ | Photometry |  |  | Spectral Properties |  |  | Spectral Type |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
|  |  |  | J | H | K | Species | $\lambda(\mu \mathrm{m})$ | EW ( $\AA$ ) |  |
| A1 | $41^{\text {s }} 04$ | $32^{\prime} 56^{\prime \prime} .8$ | 11.75(1) | 10.39(3) ${ }^{\text {a }}$ | $9.58(3)^{\text {a }}$ | $\mathrm{Br} \gamma$ | 2.166 | $-5.7 \pm 0.6$ | O5V-O6V |
|  |  |  |  |  |  | N III | 2.116 | $-2.7 \pm 0.7$ |  |
|  |  |  |  |  |  | C IV | 2.078 | $\chi^{-0.8}$ |  |
| A2 | $33^{\text {s }} .88$ | $33^{\prime} 27^{\prime \prime} 1$ | 12.310(1) | 11.02(3) ${ }^{\text {a }}$ | $10.34(2)^{\text {a }}$ | $\mathrm{Br} \gamma$ | 2.166 | $6.2 \pm 1.2$ | B0V-B1V |
|  |  |  |  |  |  | He I | 2.112 | $0.7 \pm 0.2$ |  |
| A3 | 39s.50 | $33^{\prime} 28^{\prime \prime}$ 2 | 14.063(4) | 12.646(4) | 11.97(2) | $\mathrm{Br} \gamma$ | 2.166 | $5.9 \pm 1.3$ | B2V-B3V |

${ }^{\text {a }}$ 2MASS magnitude
the positions of sources A1-A3. The measured magnitudes (see §3.3.2) and observed spectral lines for A1-A3 are presented in Table 3.1. The other two stars for which we obtained high S/N spectra have late-type spectra, as indicated by strong CO absorption at 2.29 and $2.32 \mu \mathrm{~m}$, suggesting they are either foreground objects or YSOs. The lack of nebular emission in the cluster and the presence of weak (nearly the same as in the G2V spectral standard) $\operatorname{Br} \gamma$ absorption in one of the spectra suggest that these are foreground objects rather than YSOs. In addition, the $K$ magnitudes of these objects ( $K=9.43$ and $K=10.114$ ) make them too bright to be low-mass YSOs at the cluster distance, and the presence of mainsequence $O$ and $B$ stars argues against identifying these objects as massive YSOs. We thus conclude that these two stars are most likely late-type foreground stars, and excluded them from further analysis.

Although nebular emission can be seen in the full image (Figure 3.1), it is significantly removed ( $\gtrsim 1^{\prime}$ ) from the cluster. Nevertheless, in order to ensure that


Figure 3.3: Spectra for the cluster stars in the G305.3+0.2 cluster. Top panel: Source A1, identified as O5-6V. Middle panel: Source A2, identified as B0-1V. Bottom panel: Source A3, identified as B2V-B3V.


Figure 3.4: Region immediately surrounding the G305.3+0.2 cluster. Spectroscopically classified sources are marked as A1, A2, and A3 and sources showing CO absorption are labeled with "CO". This image is approximately 120 " $\times 106$ "and the "cluster" area is marked.
any measured $\operatorname{Br} \gamma(2.166 \mu \mathrm{~m})$ is stellar in origin and not contaminated by nebular emission within the cluster, we extracted a local background spectrum for the cluster. There were no features apparent in the resulting spectrum; we thus conclude that nebular emission within the cluster is negligible. This conclusion is supported by an apparent bubble of MIR emission seen in the MSX Band A image (see Figure 3.5); the MIR emission avoids the cluster itself.

Figure 3.3 shows that source A1 has emission lines with equivalent widths stronger than $-2 \AA$ at $2.116 \mu \mathrm{~m}$ and $2.166 \mu \mathrm{~m}$ (see Table 3.1). The line at $2.166 \mu \mathrm{~m}$ is immediately identifiable as $\operatorname{Br} \gamma$. We identify the line at $2.116 \mu \mathrm{~m}$ as N III, which is consistent with the lines used in the the classification system presented in Hanson et al. (1996); the broad nature of this line is due to the multiplet nature of the transition responsible rather than broadening by stellar winds. The presence of $\operatorname{Br} \gamma$ and N III $2.116 \mu \mathrm{~m}$ in emission, without further information and without equivalent widths, is sufficient to identify the star as being an early to middle O supergiant; the broad $\mathrm{Br} \gamma$, produced in the stellar winds, is not observed in mainsequence O stars (Hanson et al., 1996). There is a possible weak detection ( $\sim 2 \sigma$ ) of C IV in emission at $2.078 \mu \mathrm{~m}$. This line only appears in O stars ranging from O5 to O6.5 (Hanson et al., 1996), and if real, significantly constrains the stellar type. Helium lines are often observed both in emission and absorption in the spectra of massive stars: He I $(2.058 \mu \mathrm{~m})$, He I $(2.112 \mu \mathrm{~m})$, and He II $(2.189 \mu \mathrm{~m})$, are all absent from the spectrum of A1. Poor removal of the telluric features near the $2.058 \mu$ mfeature prevents us from drawing any conclusions based on our nondetection. If real, the absence of the $\mathrm{He} \mathrm{I}(2.112 \mu \mathrm{~m})$ line restricts the spectral type to O6 or earlier. A He II line is expected in an O star; by estimating the strength of possible features dominated by the noise (as described in detail in § 3.3.1.2) we can place an upper limit of $0.5 \AA$ on the equivalent width of any potential


Figure 3.5: $K$-band image of the G305.3+0.2 cluster region with $8 \mu \mathrm{~m}$ contours from the MSX mission. The $K$ image has been stretched to emphasize the nebular emission. Note the close correspondence between the mid-IR emission and the nebular $K$-band emission.

He II $(2.188 \mu \mathrm{~m})$ feature. This is consistent with the width of the feature in the stars observed by Hanson et al. (1996), so the non-detection does not rule out an O star identification for this source. Taken together, these spectral characteristics suggest a spectral type of O5Ib-O6Ib for Source A1. If the weak detection of C IV is discounted, the presence of the N III line and the limit on an He II line at 2.188 $\mu \mathrm{m}$ allows an O7-O8 identification as well. Even when present, however, the C IV line is weak, with an equivalent width weaker than $-2 \AA$; thus, while a positive detection of this line would allow for definitive classification of this source as an O5Ib-O6Ib star, a non-detection at the given $\mathrm{S} / \mathrm{N}$ does not preclude the same classification.

The intrinsic NIR colors of O and B stars range from -0.08 to -0.01 (Wegner, 1994); this small range allows an extinction to be derived even without knowing the precise spectral type of a massive star. For source A1, the extinction thus derived based on the observed $H-K$ color is $A_{V}=12$ assuming the extinction law of Rieke \& Lebofsky (1985). However, the large range in absolute $M_{K}$ for O supergiants prevents us from making a distance determination based on Source A1. We can only say the distance is greater than $\sim 3.3 \mathrm{kpc}$, which would be the distance for a main-sequence O5-O6 star. Clark \& Porter (2004) adopt a distance of 4 kpc to the Danks 1 and 2 clusters in the same star formation complex, calling it an upper limit to the values allowed by the radio and $\mathrm{H} \alpha$ observations, and we will follow suit, acknowledging that the uncertainties in this value are $\sim 0.5 \mathrm{kpc}$.

Source A2 shows a strong $\operatorname{Br} \gamma(2.166 \mu \mathrm{~m})$ line in absorption with an equivalent width of $6.2 \pm 1.2 \AA$ and a probable weak He I $(2.112 \mu \mathrm{~m})$ line in absorption with $E W=0.7 \pm 0.2 \AA$. This combination of features occurs only in B stars; a comparison of the equivalent width of the lines with the B stars of Hanson et al. (1996) suggests an spectral type in the range of B2-B4. If the He I line is considered only
as an upper limit, the classification becomes more problematic, and the star could range from $\mathrm{B} 2-\mathrm{A} 2$. The star has $H-K=0.68$, which for any star in this range of spectral type excludes a foreground object. Unlike for A1, the luminosity class of these sources cannot be determined from these spectral features; as Hanson et al. (1996) points out, the $K$-band spectra of early B supergiants are indistinguishable from those of early B main-sequence stars about half the time, and those of late-B supergiants cannot be distinguished from early-B dwarfs.

If we assume that A 2 is a cluster star, we can constrain the absolute magnitude, and thus the spectral type, by requiring the distance to be the same as for the O star. Since the intrinsic near-infrared colors vary by less than 0.1 magnitude for stars in the range of spectral types allowed by the spectrum (Wegner, 1994), we can derive a extinction for this source rather than use that derived from the O star, thus reducing the effects of differential extinction. This gives an extinction to source A2 of $A_{V}=11.6$, or $A_{K}=1.3$ using the reddening law of Rieke \& Lebofsky (1985). At the distance of 4 kpc , we obtain an absolute magnitude for Source A2 of $M_{K}=-4.0$, roughly that expected for an O 8 V star. This identification is not consistent with the spectral features of A2; a smaller distance, or an identification of A2 as an early-B supergiant, could explain the spectrum of A2. If the radio distance of $3.3 \pm 0.3 \mathrm{kpc}$ is used instead, we obtain an absolute magnitude of $M_{K}=-3.3$ for Source A2, making it a B0V-B1V.

Source A3 shows only $\mathrm{Br} \gamma$ in absorption with an EW of $5.9 \pm 1.3 \AA$. We place an upper limit on an He I absorption line at $2.112 \mu \mathrm{~m}$ of $0.6 \AA$. As discussed above, this width for $\mathrm{Br} \gamma$ only constrain the classification of the star as main sequence B or early A. The observed $K$ magnitude is 11.96 , which corresponds to an absolute $M_{K} \simeq-2.6$ assuming the extinction and distance of an $\mathrm{O} 5 \mathrm{Ib}-\mathrm{O} 6 \mathrm{Ib}$ star for source A1; this is consistent with an identification of A3 as a main-sequence B1V star.

The radio distance would imply a B 2 V identification, also consistent with the spectral features of A3. Source A3 is not among the brightest stars in the cluster region; it happened to fall in the same long slit as one of the foreground contaminants we had targeted for observation. This suggests that the other cluster members brighter than $A 3$ are also late $O$ or early $B$ stars.

### 3.3.1.2 G353.4-0.36 Cluster

Spectra for the three sources observed in this cluster are presented in Fig. 3.6. An enlarged version of the relevant portion of Fig. 3.2 is presented in Fig. 3.7, with the positions of the spectroscopic targets indicated with arrows and labels. The only non-nebular feature which we detect is CO absorption in Source B1; the $\mathrm{Br} \gamma$ emission observed in all three spectra is contaminated by nebular emission to such a degree that we cannot disentangle any stellar component that may be present. While this line is much stronger in B1 than in the other two sources, the nebular emission is highly spatially variable in the cluster region and this does not demonstrate a stellar origin for the line. Additionally, the line width is significantly narrower than that of Source A1 and is similar to that observed in the offsource nebular spectrum (Fig. 3.8). The CO absorption in Source B1 in combination with the red colors (Table 3.2) are similar to those associated with solar-mass young stellar objects (YSOs) (Greene \& Lada, 1996), or a cool giant or supergiant. If B 1 is a YSO , the CO absorption is from the circumstellar material; otherwise it is photospheric in nature. Using the radio kinematic distance (Forster \& Caswell, 2000) of 3.6 kpc to the cluster, we derive an $M_{K}$ for Source B1 of -0.8 without correcting for extinction. Correcting for extinction is difficult to do accurately in this region of highly variable extinction, especially when the intrinsic colors are not known since the nature of the object is uncertain. Nevertheless, limits can be placed on the amount of extinction present, and thus the absolute magnitude of

Source B1. The lower limit is given by the uncorrected value of $M_{K}=-0.8$, which assumes the color observed is the intrinsic color, while the bright limit can be derived assuming an intrinsic $H-K=0.3$, characteristic of late-type stars; this gives an extinction to source B 1 of $A_{V}=16.6$ magnitudes and an extinction-corrected absolute $M_{K}$ of -2.6 . This is several magnitudes brighter than the expected magnitude of YSOs of approximately a solar mass at the distance and extinction of this cluster, $M_{K} \sim 1-3$ (Oasa, Tamura, \& Sugitani, 1999), and somewhat lower than the $M_{K}$ for massive YSOs, $M_{K} \sim-1$ to -5 (Ishii et al., 2001). Finally, we note that this $M_{K}$ is consistent with that for a $7 M_{\odot}$ YSO (Chakraborty et al., 2000). We conclude that if Source B1 is a YSO, it has a mass greater than a few solar masses based on its absolute magnitude in $K$, but observations of more massive YSOs are still sufficiently few that a more accurate mass determination based solely on the absolute magnitude is not possible. Given the nebular emission, seen as He I $(2.058 \mu \mathrm{~m}), \mathrm{H}_{2}(2.12 \mu \mathrm{~m})$, and $\operatorname{Br} \gamma(2.166 \mu \mathrm{~m})$ emission off the stellar sources (see Figure 3.8), G353.4-0.36 is obviously a region of current star formation; therefore, the identification as a massive YSO is more probable than a late type cool giant or supergiant located in the cluster itself.

Since Source B1 was not detected in $J$, it cannot be placed on a color-color diagram to determine whether a NIR excess is present, which could help to discriminate between the YSO and cool field star possibilities. For B1 to be a cool giant, it would need to be a foreground star with the appropriate color and magnitude, which falls by chance in the cluster region. Rather than use the entire $8^{\prime} \times$ $8^{\prime}$ field to determine the field star density, as we did for the G305.3+0.2 cluster (§3.3.2), we used only the heavily extincted region surrounding the cluster. This is because the molecular cloud in which the cluster is embedded extinguishes the background stars to such a degree that using the entire field would significantly


Figure 3.6: Spectra for sources in the G353.4-0.36 cluster. All three are identified as massive YSO candidates. The $\mathrm{Br} \gamma$ emission line seen in B 1 is contaminated by nebular emission (see Fig. 3.8).


Figure 3.7: Maser positions from the literature (Caswell et al. (2000); Argon et al. (2000); Val'tts et al. (2000)) overlaid on the G353.4-0.36 cluster K-band image. Note that they appear in regions which are dark in the near-IR, suggesting a more deeply embedded origin. Sources B1-B3 are indicated, and all cluster sources detected in $H$ and $K$ are marked with crosses.


Figure 3.8: Nebular spectrum from the G353.4-0.36 cluster region. Emission lines present are He I $2.058 \mu \mathrm{~m}, \mathrm{H}_{2} 2.12 \mu \mathrm{~m}$, and $\operatorname{Br} \gamma 2.166 \mu \mathrm{~m}$.

Table 3.2. Photometric data for the spectroscopic targets in the G353.4-0.36 cluster.

|  |  |  |  |  |  |
| :---: | :---: | :---: | :---: | :---: | :---: |
| ID | RA (2000) | DEC (2000) | J | H | K |
| B1 | $17: 30: 27.8$ | $-34: 41: 28.1$ | $\ldots$ | 14.14 | 12.85 |
| B2 | $17: 30: 27.9$ | $-34: 41: 34.7$ | $\ldots$ | 14.59 | 13.47 |
| B3 | $17: 30: 27.8$ | $-34: 41: 40$ | $\ldots$ | $\ldots$ | 14.38 |

overestimate the level of field star contamination in the immediate region of the cluster. We estimate the probability of a field source as bright as or brighter than Source B1 and red enough to satisfy the color cut falling within the cluster region to be approximately $18 \%$. This is a conservative estimate, since at the edges of the cloud reddened sources become visible and increase the field star density, especially of red objects, over what it would be at the location of the cluster. Nevertheless, we cannot rule out either a foreground giant or a YSO explanation for Source B1.

As with Source B1, the non-detection of Sources B2 and B3 in $J$ prevents us from using a color-color diagram to measure NIR excess. No photospheric features are detected in the spectra of either Source B2 or B3; Source B2 shows a rising spectrum in $K$ suggesting a strong NIR excess, while the spectrum of B3 is essentially flat in this region. In order to determine whether the spectra were truly featureless or merely had a signal-to-noise too low to see expected features, we fit a continuum to the spectra and examined all excursions above and below the fit. $90 \%$ of these deviations had an equivalent width less than $1.7 \AA$. For comparison, the detected absorption lines tabulated by Greene \& Lada (1996) for low-mass YSOs range in equivalent width from 0.3-5.6 $\AA$ for Na I and Ca I, with CO usually exceeding $2 \AA$ when present. Ishii et al. (2001) conducted a similar survey of massive YSOs; the only emission lines other than $\operatorname{Br} \gamma$ detected in a significant number of sources are CO (with an equivalent width exceeding $4 \AA$ ) and $\mathrm{H}_{2}$ (with an EW $>3 \AA$ in all cases, and $>5 \AA$ in most cases). We thus conclude that Source B2 is genuinely featureless, but cannot classify it. The final source, Source B3, had no reliably detected features but the signal-to-noise was low enough that we cannot reliably call it featureless.

The observed $K$ magnitudes are consistent with a B star identification for
sources B2 and B3; however, the extincted but distance-corrected $M_{K}$ magnitudes of $\simeq-0.2$ to -0.6 are also similar to those observed for the massive YSO $\left(M \simeq 7 M_{\odot}\right) 05361+3539$ (Chakraborty et al., 2000). Thus, although these sources are massive, we cannot distinguish based on their NIR spectra or magnitudes between shrouded B stars and less-evolved YSOs. Mid-IR observations with sufficient resolution to resolve the individual sources (separated by $\sim 5^{\prime \prime}$ ) would aid in this determination; deeper J-band photometry, detecting more of the cluster stars, would also be useful. We note that although we see ionized gas suggesting the presence of O stars, we have not detected any O stars which would be the source of the ionizing radiation in this cluster.

Due to the young age of the sources observed in this cluster and the lack of photospheric features in their spectra, the spectra were unsuitable for determining a reliable distance. Thus, the kinematic distance to the associated maser and UCHII (Forster \& Caswell, 2000) was used instead, adjusted to a distance to the Galactic Center of 8 kpc from the original 10 kpc . This gave a distance to the cluster of 3.6 kpc . Assuming an intrinsic $H-K=0$, we estimate the reddening to the cluster at $A_{V}=22$ based on the narrow cluster sequence at $H-K \simeq 1.3$ and assuming the extinction law of Rieke \& Lebofsky (1985). This estimate is highly uncertain due to the young age of the sources; many are likely to have a near-infrared excess leading to an overestimate of the line-of-sight extinction to the cluster.

### 3.3.2 Photometry

We obtained images in $J, H$, and $K_{s}$ of both clusters to a limiting magnitude of approximately $J=16, H=18, K_{s}=18.5$, with total integration times of 12 minutes in each band. The limiting magnitude was brighter than expected due to confusion, which is most noticeable in $J$ due to the slightly larger PSF and
the greater sensitivity of the instrument at shorter wavelengths. Seeing was $0^{\prime \prime} .7-$ $0^{\prime} .8$, which, since the IRIS2 platescale is $0^{\prime \prime} 45$ / pixel, resulted in a slight undersampling of the point spread function (PSF), thus making PSF fitting more uncertain. Our individual images were taken using a random dither pattern with sub-pixel dithers employed to improve the PSF. In an effort to better understand our errors we performed both PSF fitting and aperture photometry for each source. There was no systematic offset between the two methods, but the errors were $\sim 2$ times larger for the aperture photometry due to the crowded fields.

Photometric calibration was performed using the 2MASS magnitudes of field stars, after correcting from the IRIS2 filter system to the 2MASS filter system as described in Carpenter (2003). The calibrated magnitudes for the stars in the cluster area are presented in Table 3.3. The large field of view and location in the Galactic Plane provided over 100 stars in each pointing which were bright enough to have good photometry with 2MASS, but faint enough to be unsaturated in our IRIS2 images $\left(11.5<K_{s}<14\right)$. Those stars which were relatively isolated in the IRIS2 images were used as the photometric calibration set. We chose to use a relatively large number of calibration stars rather than selecting the few most isolated stars to reduce effects of potential variability and photometric outliers among the calibration stars. The scatter in the photometric calibration derived from comparison to 2MASS is the dominant source of photometric error, contributing two to three times the measurement errors as reported by DAOPHOT. DAOPHOT errors were $\simeq 0.03 \mathrm{mag}$ while the calibration uncertainties were ( $\Delta J= \pm 0.05, \Delta H= \pm 0.06$, and $\Delta K= \pm 0.06 \mathrm{mag}$ ). Quoted errors in the 2MASS photometry were negligible, with most stars having an error of $\pm 0.003$ mag or less in all bands. Thus, the quoted error should be considered an overestimate when considering the relative photometry of stars within either cluster; the
calibration errors from comparison to the 2MASS photometry will shift all our measurements by the same amount. No trend in the photometric errors, either internally or relative to the 2MASS data, was observed with location. Finally, the positions of the stars were also adjusted to agree with 2MASS by minimizing the offsets between the 2MASS and IRIS2 positions allowing for pointing offset and rotation.

Table 3.3. Photometry for all stars in G305.3+0.2 cluster region. Magnitudes $<11.5$ are taken from the 2MASS Point Source Catalog, since the IRIS2 images are saturated. Sources A1, A2, and A3 are listed first followed by the remaining sources. Field stars that were removed before deriving the luminosity and mass functions are included.

| RA (J2000) | DEC (J2000) | J | $\Delta \mathrm{J}$ | H | $\Delta \mathrm{H}$ | K | $\Delta \mathrm{K}$ |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| 13:11:41.040 | -62:32:56.77 | 11.751 | 0.010 | 10.394 | 0.027 | 9.575 | 0.029 |
| 13:11:33.877 | -62:33:27.12 | 12.310 | 0.001 | 11.020 | 0.027 | 10.342 | 0.023 |
| 13:11:39.503 | -62:33:28.17 | 14.063 | 0.004 | 12.646 | 0.004 | 11.969 | 0.018 |
| 13:11:37.680 | -62:33:09.60 | 11.949 | 0.041 | 10.776 | 0.047 | 10.185 | 0.037 |
| 13:11:36.286 | -62:33:13.30 | 12.488 | 0.010 | 11.519 | 0.001 | 10.655 | 0.037 |
| 13:11:41.620 | -62:33:17.40 | 12.936 | 0.029 | 11.650 | 0.038 | 10.953 | 0.034 |
| 13:11:41.111 | -62:33:18.39 | 16.309 | 0.036 | 14.514 | 0.018 | 11.142 | 0.002 |
| 13:11:39.268 | -62:33:24.85 | 13.467 | 0.003 | 12.095 | 0.003 | 11.494 | 0.018 |
| 13:11:39.439 | -62:33:03.63 | 13.218 | 0.003 | 12.018 | 0.002 | 11.524 | 0.018 |
| 13:11:43.767 | -62:33:26.39 | 16.065 | 0.104 | 13.312 | 0.003 | 11.594 | 0.018 |
| 13:11:40.045 | -62:33:18.89 | 13.549 | 0.006 | 12.236 | 0.006 | 11.596 | 0.018 |
| 13:11:38.141 | -62:33:13.66 | 13.428 | 0.003 | 12.270 | 0.003 | 11.745 | 0.018 |
| 13:11:39.493 | -62:33:10.25 | 16.021 | 0.129 | 14.851 | 0.116 | 11.860 | 0.004 |
| 13:11:40.458 | -62:33:03.65 | 15.989 | 0.019 | 14.766 | 0.029 | 11.920 | 0.003 |
| 13:11:40.021 | -62:33:07.26 | 13.884 | 0.004 | 12.591 | 0.003 | 11.970 | 0.018 |
| 13:11:40.992 | -62:33:07.86 | 14.258 | 0.005 | 12.845 | 0.005 | 12.123 | 0.018 |
| 13:11:36.748 | -62:33:11.14 | 14.115 | 0.010 | 12.898 | 0.005 | 12.153 | 0.053 |
| 13:11:34.747 | -62:33:24.02 | 14.067 | 0.010 | 12.742 | 0.002 | 12.311 | 0.029 |
| 13:11:34.525 | -62:33:11.13 | 14.362 | 0.010 | 12.969 | 0.004 | 12.334 | 0.044 |
| 13:11:40.031 | -62:33:11.38 | 14.315 | 0.010 | 13.098 | 0.014 | 12.518 | 0.018 |
| 13:11:40.433 | -62:33:23.29 | 17.628 | 0.108 | 15.615 | 0.049 | 12.698 | 0.005 |
| 13:11:39.217 | -62:33:08.44 | 14.799 | 0.010 | 13.578 | 0.007 | 12.789 | 0.164 |

Table 3.3-Continued

| RA (J2000) | DEC (J2000) | J | $\Delta \mathrm{J}$ | H | $\Delta \mathrm{H}$ | K | $\Delta \mathrm{K}$ |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| 13:11:38.080 | -62:32:59.21 | 16.267 | 0.017 | 13.972 | 0.005 | 12.872 | 0.018 |
| 13:11:40.927 | -62:33:21.57 | 15.859 | 0.156 | 13.841 | 0.005 | 12.890 | 0.018 |
| 13:11:40.921 | -62:33:12.54 | 15.083 | 0.010 | 13.691 | 0.005 | 12.924 | 0.088 |
| 13:11:35.589 | -62:33:00.04 | 14.940 | 0.006 | 13.605 | 0.004 | 12.943 | 0.018 |
| 13:11:35.589 | -62:33:00.04 | 14.940 | 0.006 | 13.605 | 0.004 | 12.943 | 0.018 |
| 13:11:37.319 | -62:33:27.31 | 17.507 | 0.031 | 15.832 | 0.015 | 13.025 | 0.007 |
| 13:11:36.852 | -62:33:07.66 | 14.888 | 0.007 | 13.527 | 0.111 | 13.057 | 0.064 |
| 13:11:38.496 | -62:33:26.35 | 15.185 | 0.010 | 13.835 | 0.006 | 13.075 | 0.065 |
| 13:11:38.934 | -62:32:56.60 | 15.989 | 0.114 | 14.069 | 0.006 | 13.202 | 0.018 |
| 13:11:42.785 | -62:33:27.36 | 14.640 | 0.004 | 13.651 | 0.003 | 13.236 | 0.018 |
| 13:11:36.468 | -62:33:04.13 | 14.966 | 0.006 | 13.826 | 0.004 | 13.296 | 0.018 |
| 13:11:35.686 | -62:33:20.90 | 15.665 | 0.010 | 14.064 | 0.004 | 13.419 | 0.054 |
| 13:11:41.200 | -62:32:59.20 | 13.887 | 0.004 | 12.878 | 0.004 | 13.454 | 0.012 |
| 13:11:38.522 | -62:33:20.38 | 15.966 | 0.017 | 14.812 | 0.013 | 13.482 | 0.011 |
| 13:11:42.481 | -62:33:13.12 | 15.335 | 0.007 | 14.117 | 0.005 | 13.488 | 0.018 |
| 13:11:40.482 | -62:32:53.83 | 19.190 | 0.398 | 17.987 | 0.315 | 13.649 | 0.000 |
| 13:11:34.324 | -62:33:08.70 | 15.574 | 0.010 | 14.271 | 0.010 | 13.696 | 0.007 |
| 13:11:34.976 | -62:33:00.49 | 15.383 | 0.007 | 14.369 | 0.063 | 13.718 | 0.060 |
| 13:11:42.104 | -62:32:54.73 | 15.935 | 0.095 | 14.319 | 0.007 | 13.749 | 0.072 |
| 13:11:39.128 | -62:33:35.01 | 15.994 | 0.124 | 14.525 | 0.006 | 13.757 | 0.063 |
| 13:11:40.731 | -62:33:00.95 | 16.011 | 0.032 | 14.742 | 0.033 | 13.772 | 0.018 |
| 13:11:40.916 | -62:33:05.08 | 15.901 | 0.015 | 14.624 | 0.020 | 13.850 | 0.018 |

Table 3.3-Continued

| RA (J2000) | DEC (J2000) | J | $\Delta \mathrm{J}$ | H | $\Delta \mathrm{H}$ | K | $\Delta \mathrm{K}$ |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| 13:11:43.215 | -62:33:16.25 | 16.627 | 0.027 | 14.772 | 0.008 | 13.852 | 0.018 |
| 13:11:38.346 | -62:33:07.82 | 15.943 | 0.021 | 15.327 | 0.030 | 13.936 | 0.043 |
| 13:11:42.225 | -62:33:03.31 | 16.072 | 0.013 | 14.719 | 0.008 | 13.945 | 0.018 |
| 13:11:39.514 | -62:32:57.00 | 16.277 | 0.344 | 14.673 | 0.007 | 13.988 | 0.018 |
| 13:11:37.032 | -62:33:10.62 | 13.976 | 0.005 | 12.769 | 0.005 | 14.080 | 0.016 |
| 13:11:41.817 | -62:33:10.73 | 14.808 | 0.007 | 13.539 | 0.006 | 14.099 | 0.018 |
| 13:11:36.506 | -62:33:17.45 | 17.022 | 0.040 | 15.022 | 0.021 | 14.159 | 0.019 |
| 13:11:40.797 | -62:33:17.11 | 16.149 | 0.024 | 14.910 | 0.021 | 14.196 | 0.018 |
| 13:11:37.887 | -62:33:19.60 | 14.781 | 0.010 | 13.549 | 0.004 | 14.217 | 0.054 |
| 13:11:39.511 | -62:32:51.64 | 15.999 | 0.098 | 14.909 | 0.009 | 14.391 | 0.018 |
| 13:11:34.336 | -62:33:15.46 | 15.337 | 0.067 | 14.803 | 0.008 | 14.427 | 0.018 |
| 13:11:36.437 | -62:33:29.40 | 15.331 | 0.010 | 14.520 | 0.005 | 14.478 | 0.107 |
| 13:11:36.523 | -62:33:21.46 | 17.406 | 0.051 | 16.853 | 0.078 | 14.501 | 0.019 |
| 13:11:42.858 | -62:33:16.72 | 18.010 | 0.079 | 17.233 | 0.076 | 14.504 | 0.018 |
| 13:11:36.825 | -62:33:25.01 | 17.507 | 0.031 | 15.818 | 0.016 | 14.506 | 0.013 |
| 13:11:41.219 | -62:33:27.39 | 15.826 | 0.014 | 15.700 | 0.034 | 14.540 | 0.018 |
| 13:11:37.697 | -62:33:29.49 | 15.620 | 0.007 | 14.801 | 0.008 | 14.542 | 0.018 |
| 13:11:33.579 | -62:32:56.13 | 18.305 | 0.075 | 17.463 | 0.073 | 14.577 | 0.026 |
| 13:11:35.662 | -62:33:01.78 | 14.779 | 0.006 | 13.466 | 0.005 | 14.639 | 0.016 |
| 13:11:43.148 | -62:32:55.34 | 16.616 | 0.019 | 15.964 | 0.017 | 14.687 | 0.025 |
| 13:11:41.833 | -62:33:36.71 | 15.949 | 0.085 | 14.985 | 0.006 | 14.727 | 0.010 |
| 13:11:37.306 | -62:32:55.60 | 18.370 | 0.073 | 17.131 | 0.029 | 14.758 | 0.022 |

Table 3.3-Continued

| RA (J2000) | DEC (J2000) | J | $\Delta \mathrm{J}$ | H | $\Delta \mathrm{H}$ | K | $\Delta \mathrm{K}$ |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| 13:11:40.602 | -62:33:28.71 | 16.143 | 0.142 | 15.212 | 0.014 | 14.769 | 0.018 |
| 13:11:42.593 | -62:33:33.58 | 15.585 | 0.007 | 15.018 | 0.006 | 14.844 | 0.018 |
| 13:11:38.673 | -62:33:13.70 | 17.234 | 0.098 | 15.935 | 0.085 | 14.866 | 0.018 |
| 13:11:39.064 | -62:33:02.35 | 15.945 | 0.012 | 15.485 | 0.017 | 14.886 | 0.018 |
| 13:11:33.958 | -62:33:23.58 | 16.264 | 0.019 | 15.592 | 0.022 | 15.018 | 0.018 |
| 13:11:35.334 | -62:33:16.22 | 17.619 | 0.042 | 15.712 | 0.023 | 15.178 | 0.017 |
| 13:11:35.168 | -62:33:35.69 | 18.451 | 0.064 | 17.257 | 0.049 | 15.207 | 0.017 |
| 13:11:32.967 | -62:32:57.08 | 17.237 | 0.034 | 16.541 | 0.042 | 15.235 | 0.047 |
| 13:11:36.074 | -62:33:33.56 | 16.513 | 0.014 | 15.713 | 0.011 | 15.307 | 0.018 |
| 13:11:35.630 | -62:33:33.46 | 17.193 | 0.024 | 16.109 | 0.015 | 15.353 | 0.018 |
| 13:11:33.120 | -62:33:02.23 | 16.745 | 0.018 | 15.919 | 0.013 | 15.410 | 0.018 |
| 13:11:37.787 | -62:32:55.19 | 17.263 | 0.038 | 16.257 | 0.033 | 15.427 | 0.018 |
| 13:11:42.040 | -62:33:29.12 | 14.055 | 0.003 | 13.571 | 0.004 | 15.433 | 0.027 |
| 13:11:41.102 | -62:33:29.12 | 16.458 | 0.022 | 15.874 | 0.034 | 15.447 | 0.018 |
| 13:11:35.050 | -62:33:06.75 | 18.794 | 0.176 | 16.753 | 0.069 | 15.455 | 0.018 |
| 13:11:40.603 | -62:33:33.71 | 16.584 | 0.014 | 15.839 | 0.015 | 15.483 | 0.018 |
| 13:11:41.752 | -62:33:24.65 | 17.106 | 0.027 | 16.257 | 0.030 | 15.492 | 0.018 |
| 13:11:37.052 | -62:33:14.86 | 17.400 | 0.073 | 16.320 | 0.080 | 15.526 | 0.018 |
| 13:11:35.505 | -62:33:24.18 | 16.469 | 0.019 | 15.808 | 0.015 | 15.535 | 0.018 |
| 13:11:38.514 | -62:32:52.00 | 17.306 | 0.033 | 16.343 | 0.045 | 15.598 | 0.018 |
| 13:11:43.176 | -62:33:34.55 | 16.697 | 0.015 | 16.026 | 0.019 | 15.632 | 0.018 |
| 13:11:36.480 | -62:33:24.14 | 18.355 | 0.126 | 16.536 | 0.039 | 15.633 | 0.018 |

Table 3.3-Continued

| RA (J2000) | DEC (J2000) | J | $\Delta \mathrm{J}$ | H | $\Delta \mathrm{H}$ | K | $\Delta \mathrm{K}$ |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| 13:11:42.166 | -62:32:57.78 | 17.266 | 0.054 | 15.944 | 0.029 | 15.640 | 0.034 |
| 13:11:34.992 | -62:33:09.74 | 19.001 | 0.372 | 17.187 | 0.185 | 15.641 | 0.018 |
| 13:11:40.881 | -62:32:53.49 | 16.019 | 0.013 | 14.877 | 0.012 | 15.653 | 0.058 |
| 13:11:41.177 | -62:33:25.05 | 15.826 | 0.014 | 15.164 | 0.021 | 15.702 | 0.043 |
| 13:11:34.257 | -62:32:54.79 | 20.060 | 0.353 | 17.158 | 0.060 | 15.727 | 0.018 |
| 13:11:43.097 | -62:32:57.73 | 17.839 | 0.043 | 17.029 | 0.048 | 15.811 | 0.018 |
| 13:11:43.288 | -62:32:57.13 | 17.839 | 0.043 | 16.953 | 0.040 | 15.811 | 0.018 |
| 13:11:40.725 | -62:33:25.98 |  | . | 18.059 | 0.464 | 15.900 | 0.018 |
| 13:11:43.017 | -62:33:32.66 |  |  | 16.922 | 0.048 | 15.907 | 0.033 |
| 13:11:33.512 | -62:33:23.69 | 18.259 | 0.078 | 17.578 | 0.083 | 15.924 | 0.042 |
| 13:11:33.400 | -62:33:32.23 | 18.288 | 0.079 | 16.771 | 0.037 | 15.988 | 0.018 |
| 13:11:34.286 | -62:33:05.34 | 18.841 | 0.107 | 17.051 | 0.037 | 16.055 | 0.018 |
| 13:11:41.668 | -62:33:01.67 | 19.070 | 0.319 | 16.667 | 0.162 | 16.057 | 0.080 |
| 13:11:34.549 | -62:33:20.84 | 18.643 | 0.094 | 16.735 | 0.050 | 16.234 | 0.035 |
| 13:11:35.534 | -62:33:29.34 | 18.549 | 0.072 | 16.926 | 0.027 | 16.281 | 0.052 |
| 13:11:33.766 | -62:33:13.88 | 17.567 | 0.041 | 16.791 | 0.044 | 16.337 | 0.018 |
| 13:11:35.640 | -62:32:55.46 | 16.806 | 0.026 | 15.517 | 0.013 | 16.433 | 0.077 |
| 13:11:35.929 | -62:32:59.16 | 14.779 | 0.006 | 16.188 | 0.044 | 16.591 | 0.085 |
| 13:11:35.269 | -62:32:56.92 | 20.171 | 1.344 | 16.995 | 0.133 | 16.675 | 0.091 |
| 13:11:43.085 | -62:33:20.66 |  | $\ldots$ | 16.479 | 0.025 | 16.733 | 0.075 |
| 13:11:36.278 | -62:32:55.09 | 19.388 | 0.197 | 17.669 | 0.077 | 16.848 | 0.096 |
| 13:11:37.055 | -62:33:35.61 | 17.021 | 0.022 | 16.354 | 0.027 | 16.979 | 0.161 |

Table 3.3-Continued

|  |  |  |  |  |  |  |  |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| RA (J2000) | DEC (J2000) | J | $\Delta \mathrm{J}$ | H | $\Delta \mathrm{H}$ | K | $\Delta \mathrm{K}$ |
| 13:11:33.733 | $-62: 33: 00.39$ | 17.536 | 0.032 | 16.652 | 0.029 | 17.264 | 0.248 |
| $13: 11: 36.172$ | $-62: 33: 01.75$ | 16.469 | 0.032 | 16.023 | 0.046 | 18.288 | 0.393 |

### 3.3.2.1 G305.3+0.2

The color composite of the full $J, H$, and $K_{s}$ images is presented in Fig. 3.1; the cluster alone is shown in Fig. 3.4, with the spectroscopic targets marked. The cluster is clearly visible in the full-size image with a concentration of nebular emission to the northwest. In order to help determine whether the nebular emission is physically associated with the cluster, we overplotted the contours at $8 \mu \mathrm{~m}$ from the MSX mission ${ }^{2}$. (Fig. 3.5). The ridge of near-IR nebulosity corresponds to the brightest portion of a roughly circular structure of mid-IR emission, with the cluster located in the interior where there is no mid-IR emission present. The general appearance is that of a wind-blown bubble, and the $8 \mu \mathrm{~m}$ emission wraps entirely around the cluster at a lower level. The cluster is located off-center in this structure, near the brightest portion of the mid-IR emission, but there is no mid-IR emission and no near-IR nebulosity present in the area of the cluster itself. The cluster is dense and well-defined, with stellar density much higher than in the field.

The $K$ versus $H-K$ color-magnitude diagram of the cluster region is shown in Figure 3.9. At radii of approximately $30^{\prime \prime}$ in the east-west direction and $20^{\prime \prime}$ in the north-south direction from the cluster center the stellar density has fallen to

[^1]

Figure 3.9: K vs H-K for the G305.3+0.2 cluster region, with all sources plotted. Typical error bars are smaller than the circles. Filled symbols designate spectroscopic targets A1-A3.
that of the field, which we used to define the cluster region. Foreground stars are apparent in the color-magnitude diagram at $H-K \sim 0.3$; in this cluster there is no clear separation in color between cluster and field stars, just an overdensity of redder stars in the cluster; as a result, we cannot impose a firm color cut to separate field stars from cluster stars. A color-magnitude diagram of a randomly selected control field with the same area as the cluster is shown in Fig. 3.10; many fewer stars are present, especially at bright magnitudes and moderately red colors.

In order to account for field star contamination within the cluster region, we


Figure 3.10: K vs H-K for a randomly selected control field for G305.3+0.2, with the same area as the cluster field.
determined the average number of stars per square arcminute in the image outside the cluster region in color-magnitude bins of $\Delta K=0.5, \Delta(H-K)=0.5$ and randomly selected the appropriate number of stars from the cluster field for removal. This is similar to the procedure employed by, among others, Blum, Conti, \& Damineli (2000) and Figuerêdo et al. (2002). In cases where less than one star was expected in the cluster field in a particular color-magnitude bin, the number expected was used as a probability for removing a star. A total of 24 "field" stars were removed, leaving 115. The main concentration of cluster stars is at about $H-K=0.8$, with a gradually declining number present out to $H-K \sim 4$. The resulting cluster CMD with the field stars statistically removed is shown in Fig. 3.11. Given the spectroscopically confirmed presence of OB stars in the cluster, as well as the lack of an obvious color gap, we consider it more likely that these very red sources are either background sources or sources with a near-IR excess due to local dust than that they represent a separate cluster giant branch. The red sources are not concentrated toward any part of the cluster, though they may occur more frequently on the outskirts (as would be expected if they are background objects). Sources redder than $H-K=1.5$ were excluded from analysis of the cluster KLF and IMF; they are unlikely to be main-sequence cluster members. If they are included and assumed to be on the main sequence, the resulting extinction correction would give very large values for the masses and an overly flat slope to the IMF. If these sources are cluster members, they are pre-main-sequence objects, and their masses are difficult to determine from $H$ and $K$ photometry alone. Thus, including them in the IMF determination would give an inaccurate result whether or not they are cluster members, and they have been excluded. Finally, the crowded nature of the cluster region means that these very red sources may suffer from poor photometry.


Figure 3.11: Distance- and reddening-adjusted K vs. H-K for the G305.3+0.2 cluster region, statistically corrected for field star contamination. The ZAMS from the Meynet \& Maeder (2003) evolutionary models has been transformed to the observed quantities and overplotted.

The $J-H$ vs. $H-K$ color-color diagram (Fig. 3.12) is of limited utility in identifying cluster members or determining whether some cluster members are pre-main-sequence objects. Since many sources were undetected in $J$, it will not represent all cluster members, and faint red sources (where we would expect to find the relatively low-mass, pre-main-sequence objects) would be most commonly missed in the color-color diagram. A cut based solely on $H-K$ must still be applied to exclude background sources. Fig. 3.12 shows few sources in the area occupied by pre-main-sequence objects. Of those sources separated from reddened main-sequence stars by more than $3 \sigma$, three are relatively faint sources adjacent to bright sources and one is in a particularly crowded region. The remaining three could potentially be pre-main-sequence objects. However, due to the lack of observed gaseous emission from the cluster, we consider it unlikely that these are truly pre-main-sequence stars, and exclude them from the analysis along with the objects in the unphysical blue region of the color-color diagram as likely suffering from blending or a mismatch between sources in the different bandpasses. There are few enough sources in this region that we do not expect their inclusion or exclusion to greatly affect the IMF determination.

### 3.3.2.2 G353.4-0.36

The $J, H$, and $K$ color composite of G353.4-0.36 is presented in Fig. 3.2. The youth of this cluster is immediately apparent from its heavily embedded nature and the dense molecular cloud that surrounds it. This region has long been known to be a site of massive star formation, and it has been studied extensively in the radio and sub-mm, including continuum observations at $1.5 \mathrm{GHz}, 5 \mathrm{GHz}$ (Becker et al., 1994), and $850 \mu \mathrm{~m}$ (Carey et al., 2000) as well as molecular line observations in CS (Gardner \& Whiteoak, 1978), CO (Whiteoak, Otrupcek, \& Rennie, 1982), $\mathrm{H}_{2} \mathrm{CO}$ (Gardner \& Whiteoak, 1984), HNCO (Zinchenko, Henkel, \& Mao, 2000)


Figure 3.12: Distance- and reddening-adjusted $J-H$ vs. $H-K$ for the G305.3+0.2 cluster region, statistically corrected for field star contamination. Error bars are comparable to the size of the points or smaller. Very few sources fall outside the region of reddened main-sequence stars by more than $2 \sigma$.
(identified as a dense molecular core), and SiO (Harju et al., 1998). These signatures of ongoing star formation, combined with the strong nebular emission still present around the sources observed spectroscopically, suggest that the cluster is quite young, without main-sequence stars. Many of the continuum and molecular line observations quote slightly different positions for the source peak, and sources separated by several tens of arcseconds are all identified with the IRAS point source 17271-3439. Since the beam sizes in many instances are comparable to the size of the NIR-bright nebulosity and to the separation between sources, it is likely that the extended source measurements are observing the same complex, which may peak at different locations in different wavelengths. Many of the radio data are tabulated by Chan, Henning, \& Schreyer (1996), who identify a massive YSO in the region based on the IRAS colors. It is obvious from the NIR imaging that this source is not a single point source; in addition to the NIR sources, there are at least four separate sets of masers (e.g. Caswell et al., 2000; Argon et al., 2000; Val'tts et al., 2000), one of which is associated with an UCHII (Forster \& Caswell, 2000). Positions of the masers are indicated in Fig. 3.7. We note that the masers occur in regions which are heavily extincted in the near-IR. $\mathrm{OH}, \mathrm{H}_{2} \mathrm{O}$, and $\mathrm{CH}_{3} \mathrm{OH}$ masers are all known in the region; the latter in particular are indicative of ongoing massive star formation. Clearly the sources visible in the near-infrared are only the tip of the iceberg, with other massive stars still in the process of formation. Higher-resolution maps at radio and sub-mm wavelengths are necessary to obtain a full understanding of this region.

In the region of the large dark molecular cloud, only foreground stars are visible. This implies $A_{V}>50$ in order to completely obscure the stars even in $K$, assuming a $K$-band detection limit of 17 and a distribution of $K$ magnitudes similar to the rest of the field. The less heavily extincted region in which the
cluster is visible in the near-IR must have been partially cleared out by stellar winds and ionization from massive stars. The relative position of the NIR stars and the methanol masers (which lie in regions of higher extinction) suggest that we are observing stars nearer the main sequence which are emerging from the dust, while objects at an earlier evolutionary state are offset from this region, indicating ongoing star formation.

The color-magnitude diagram of the G353.4-0.36 cluster is presented in Figure 3.13. The cluster sequence is much narrower and more well-separated than in the G305.3+0.2 cluster, allowing for reliable separation of foreground objects based solely on $H-K$. Thus, we did not carry out a statistical removal of foreground objects for this cluster, instead considering only the objects well-separated from the foreground sequence. Due to the high extinction toward this cluster, a large number of objects in the cluster area were detected only in $K$ (shown as limits in Figure 3.13). The KLF is thus likely to be more reliable than the colormagnitude and color-color diagrams in determining cluster characteristics.

### 3.4 The K Band Luminosity Function and the Initial Mass Function

Once field stars have been rejected as described in $\S 3.3 .2 .1$, we can compute the KLF for both clusters. For the G305.3+0.2 cluster, which has more than 100 stars remaining, we additionally compute the initial mass function (IMF) using two different techniques, the first using the KLF and the second using the colormagnitude diagram and the spectroscopy of the massive stars. The KLF is commonly used to determine the IMF even when multi-color photometry is available; we take this opportunity to test the robustness of this method and compare the results between this simple and commonly used method and the more involved method using the color-magnitude diagram. This will help to understand the un-


Figure 3.13: K vs H-K for the G353.4-0.36 cluster region. Note the clear separation in color between cluster and foreground sources. Sources B1 and B2 are denoted by filled symbols; Source B3 was not detected in H.
certainties and systematic errors that may be a factor when only the KLF method can be used to derive an IMF. There were too few stars to robustly compute the IMF for the G353-0.4 cluster, so we compute only the KLF in this case.

### 3.4.1 The G305.3+0.2 Cluster

To provide a robust determination of the KLF and the IMF, we must determine the completeness of our data, which we established by performing artificial star tests. Five artificial stars at a time were inserted into the cluster region; the small number was chosen to avoid significantly changing the crowding characteristics. IRAF-DAOPHOT was then run on the images to determine the number of artificial stars that were successfully recovered. The procedure was repeated 50 times for each magnitude bin $(\Delta m=0.5)$, for a total of 250 artificial stars added in each bin in $H$ and in $K$. Figure 3.14 shows the results; completeness falls sharply to about $25 \%$ at $H \sim 16.5, K \sim 15.5$. We can compare these magnitudes with the turnover in the "field luminosity function", which also probes incompleteness. The counts in the field turned over sharply at $K \simeq 16$, in reasonable agreement with the artificial star estimate of incompleteness.

### 3.4.1.1 The $K$ Luminosity Function

Knowing our incompleteness, we can calculate the KLF for the cluster. Figure 3.15 shows the uncorrected data, with the field "luminosity function" normalized to the same total number of stars overplotted for comparison. Figure 3.16 shows the results after correcting for incompleteness by dividing the number of stars in each magnitude bin by the recovered fraction of artificial stars. As expected, there is an overabundance of bright stars $(K<14.5)$ in the cluster region relative to the field. This is not an artifact of incompleteness; the completeness fraction at this magnitude is $\sim 90 \%$, and we expect incompleteness to be higher


Figure 3.14: Completeness fraction determined by artificial-star tests for the G305.3+0.2 cluster region ( $K=$ solid line, $H=$ dashed line).


Figure 3.15: K-band luminosity functions for G305.3+0.2 cluster (solid) and field (dashed), normalized to the same total number of stars. Note the peak is shifted to brighter magnitudes for the cluster.
in the cluster than the field due to the effects of crowding. Using the number counts corrected for field star contamination (as discussed in §3.3.2.1) and incompleteness, we fit a slope to the number counts in bins of $\Delta K=0.5$. We excluded sources fainter than $K=15.5$ from the fit since errors in the incompleteness determination are likely to dominate the number counts. We derived a slope of $0.21 \pm 0.06$ for $\log N_{*}$. This slope is somewhat flatter than the KLFs derived for more massive embedded clusters (e.g. $0.41 \pm 0.02$ for NGC 3576 from Figuerêdo et al. (2002), $0.40 \pm 0.03$ for W42 from Blum et al. (2000)). This suggests that this cluster is more weighted toward massive stars than the norm.

### 3.4.1.2 The Initial Mass Function

In order to better compare our results with the literature, and to explore how much of a difference the use of multi-color photometry and spectra of the massive stars make in the determination of the IMF, we used two methods to derive


Figure 3.16: Completeness-corrected K-band luminosity function for G305.3+0.2 cluster (dashed). The uncorrected KLF is overplotted (solid) for comparison.
an IMF for the G305+00.2 cluster. For both methods we use a distance to the cluster of 4.0 kpc (as discussed in §3.3.1.1). The first IMF-determination method, which uses only the KLF, is commonly employed even when multi-color photometry and spectra are available (e.g. Figuerêdo et al., 2002; Blum et al., 2000). This method is simply a transformation from $K$ magnitude bins to mass bins. To make this transformation, we first correct the observed $K$ for distance and extinction as discussed in §3.3.1.1. Using the stellar evolutionary models of Meynet \& Maeder (2003) for solar metallicity, we relate the mass for each track to an absolute $K$ magnitude for a star on the ZAMS. We transformed $L_{b o l}$ to $K$ using the bolometric corrections from Vacca, Garmany, \& Shull (1996) for the early spectral types and Malagnini et al. (1986) for later spectral types. We then use the intrinsic $V-K$ colors from Bessell \& Brett (1988) for A-M stars and from Wegner (1994) for O and B stars. Finally we interpolate linearly between the masses available on the evolutionary tracks to find the masses corresponding to our magnitude bins, and fit a power law to the resulting mass function. Our resulting IMF slope
is $\Gamma=-1.5 \pm 0.3$, excluding the two lowest-mass bins where incompleteness is significant.

Our second method of determining the IMF made use of our multi-color photometry and spectra to estimate individual extinctions and masses for cluster members. Spectral typing of the brightest cluster stars allows their mass to be determined fairly accurately for a given stellar evolutionary model. For the models described above, the mass of an O6V star is approximately $40 \mathrm{M}_{\odot}$, that of a B0V star is $15 \mathrm{M}_{\odot}$, and that of a B2V star is $8 \mathrm{M}_{\odot}$. Although spectra are not available for most of the cluster stars, their masses, as well as extinctions to the individual stars, can be estimated from the accurate relative photometry. The presence of an O supergiant in the cluster suggests that, while the most massive stars have begun to evolve away from the main sequence, none have yet gone supernova, and less massive stars should still be on the zero-age main sequence. Therefore, with the exception of the few most massive stars (for which we can estimate masses from their spectral types) the cluster stars should be scattered around the zeroage main sequence (ZAMS) primarily by differential extinction and rather than the effects of stellar evolution. We can then use the same models and conversions from theoretical to observed quantities described for the KLF method, with additional transformations from $T_{\text {eff }}$ to $H-K$ using intrinsic colors from from Bessell \& Brett (1988) and Wegner (1994) and from $T_{\text {eff }}$ to spectral type from Repolust et al. (2004) or Johnson (1966).

This transformation from theoretical to observed quantities allows us to place the ZAMS on our CMD. If the cluster is sufficiently young that we can neglect the effects of stellar evolution, as discussed in the previous paragraph, we expect the ZAMS will lie in the middle of the distribution of cluster stars. The ZAMS derived from the evolutionary tracks of Meynet \& Maeder (2003) is overplotted on
the distance and extinction-corrected CMD in Figure 3.11. A significant number of stars are bluer than the ZAMS on this plot. We interpret these as stars which are less extincted than those used to determine the average cluster extinction and thus have been over-corrected by using the mean extinction. The scatter of stars around the ZAMS suggests that the extinction varies across the cluster region. To correct for this, we move the stars along the direction of the reddening vector until they lie on the ZAMS. If the resulting extinction differs from the mean cluster value by more than $A_{V}=5$ for a given star, we exclude the star from the analysis, as it probably suffers from poor photometry. Examination of the color image of the cluster region (Figure 3.1) suggests that the variation in internal extinction in this region is relatively small; no dust lanes or color variations across the cluster are visible to the eye. The exact value selected for the cutoff is somewhat arbitrary, but does not greatly affect the results; most of the sources thus excluded have derived extinctions that differ from the median value by $A_{V}=10$ or more.

Using the positions of the extinction-corrected photometry along the ZAMS, we are able to more accurately place stars in mass bins. The endpoints of the bins were determined by the masses for which theoretical tracks are present in the models we used. In order to have an adequate number of stars in each bin we constructed bins using alternate tracks for the endpoints, rather than every track. The analysis was repeated for three different metallicities ( $Z=0.1,0.02,0.001$ ) using the evolutionary tracks of Mowlavi et al. (1998, $\mathrm{Z}=0.1$ ), Schaller et al. (1992); Meynet \& Maeder (2003, $\mathrm{Z}=0.02$ ) and Schaller et al. (1992, $\mathrm{Z}=0.001$ ). For the solar-metallicity case the high-mass points $\left(M>9 M_{\odot}\right)$ are from Meynet \& Maeder (2003) while the lower-mass points are from Schaller et al. (1992). The difference in $K$ for the two solar-metallicity tracks is always less than 0.1 magnitudes for the masses where the two sets of tracks overlap and for most masses is
less than 0.03 magnitudes. The high metallicity model should be considered only as a limiting case since such a high metallicity is not expected. The use of such a wide range of metallicities allows us to estimate the importance of this parameter on the final IMF determination.

Given these sets of mass bins, for each metallicity we determine the number of stars per unit logarithmic mass interval after correcting for completeness. We then fit a power law to the data. The two lowest-mass bins ( $M<2 M_{\odot}$ ), where incompleteness was significant, were excluded from the fit; uncertainty in the completeness correction applied could significantly influence the results in these mass bins. The resulting completeness-corrected IMF for the cluster is plotted in Figure 3.17. The solar-metallicity models yield an IMF slope $\Gamma=-0.98 \pm 0.2$, where the quoted errors are only the formal fit errors and should be considered an underestimate. The low-metallicity tracks yield $\Gamma=-1.01 \pm 0.2$ for the same distance, suggesting that the cluster IMF determination is insensitive to metallicity for solar and sub-solar values. The $Z=0.1$ tracks give $\Gamma=-0.88 \pm 0.15$.

### 3.4.1.3 Comparison of the IMF Methods

The IMF slopes we derive using these two methods are marginally consistent within the error bars: $\Gamma=-1.5 \pm 0.3$ for the KLF method, and $\Gamma=-0.98 \pm 0.2$ for the CMD + spectroscopy method assuming solar metallicity. Comparing these results individually to the Salpeter slope would lead to different conclusions, however. The KLF method produces a slope that is very close to the Salpeter value, while the slope from the CMD + spectroscopy method differs from Salpeter by about $2 \sigma$. While this difference in slopes could arise purely from statistical uncertainty, various systematic effects should cause the KLF-derived slope to be steeper than the CMD-derived slope, as we observe. If the more massive stars are preferentially located toward the center of the cluster, as expected due to mass


Figure 3.17: The completeness-corrected IMF for the G305.3+0.2 cluster. The fitted values are the masses derived from the Schaller et al. (1992) stellar evolutionary tracks with a distance to the cluster of 3.4 kpc . The plotted error bars are given by assuming the error in the number of stars in a mass bin is equal to the square root of the number of stars in the bin. The fitted line has a slope of -0.96 .
segregation, and if the extinction is higher in the center of the cluster, the mean extinction used in the KLF determination would be systematically low for the more massive stars. This method would then underestimate the masses the highest mass stars, thus steepening the slope of the IMF. Evidence that this effect may be at work is provided by the six brightest cluster members, all of which lie redward of the ZAMS in Figure 3.11 while the fainter members are scattered more evenly. A difference in $A_{K}$ (and thus $M_{K}$ ) of 0.2 corresponds to 1-2 subtypes for massive stars and thus to a difference in the derived mass of at least $2 \mathrm{M}_{\odot}$.

However, mass segregation can only provide a partial explanation for the difference in the IMF slopes; the stars for which we obtained spectra are not in the very center of the cluster (since crowding in the 2MASS image used to select spectroscopic targets prevented us from selecting targets in the cluster core). An additional possible source of systematic error in the KLF method relative to the CMD method lies in field star rejection. In addition to the statistical field star rejection described in §3.3.2.1, which was done before any further analysis and thus applies to both methods, the CMD method has color-based field star rejection. The CMD method can reject foreground objects, which due to lower extinction are bluer than cluster objects, as well as background objects which are redder than the cluster. The KLF method includes these objects, which tend to be fainter on average than the cluster stars (since they are either at a greater distance or are low-mass foreground stars) and thus finds an artificially high number of lowmass stars. We find the use of $K$ photometry alone to derive the IMF is likely to produce an overly steep IMF in regions with significant field contamination or variable extinction.

### 3.4.1.4 Comparison With Other Young Stellar Clusters

Most studies of young star clusters have found an initial mass function consistent with the Salpeter slope of $\Gamma=-1.35$, generally with uncertainties of 0.1-0.2 (e.g. Figuerêdo et al., 2002; Massey, 1998; Hillenbrand \& Carpenter, 2000; Okumura et al., 2000), including the extremely massive R136 cluster in the LMC (Massey \& Hunter, 1998). A review of the results is provided in Massey (2003). In the case of NGC 6611, reanalysis of the same data by different authors has produced dramatically different results; an IMF of $-1.1 \pm 0.1$ was found by Hillenbrand et al. (1993), while a reanalysis with different treatment of extinction produced $-0.7 \pm 0.2$ (Massey et al., 1995), suggesting that the systematic effects are at work in IMF determinations that are at least as important as the statistical errors, as we see in this work. Slopes significantly flatter than Salpeter have been reported for the Arches cluster near the Galactic Center (Figer et al., 1999), though later work suggests that this result is an artifact of mass segregation; Stolte et al. (2002) found a very flat IMF in the core of the Arches Cluster with a steeper IMF at larger radii, with an overall slope consistent with a Salpeter value. The flatness we observe in both the KLF and the IMF for the G305+00.2 cluster using the CMD + spectroscopy method may similarly be due to mass segregation. In addition to the extinction effects mentioned previously, fainter stars in the outskirts of the cluster could be indistinguishable from the field star density (especially given the high field star density due to the location of the cluster in the Galactic plane) and not fall within the cluster boundaries we employ.

### 3.4.2 The KLF for the G353.4-0.36 cluster

Completeness tests were performed for the G353.4-0.36 cluster using artificial stars as discussed above, and the completeness-corrected KLF is plotted in Figure 3.18. Since the cluster is significantly less crowded and faint cluster stars less


Figure 3.18: The raw (solid line) and completeness-corrected (dashed line) KLF for the G353.4-0.36 cluster. Only sources with $H-K>1$, which fall in the cluster color sequence, have been included.
common, our detections in this cluster are nearly complete in $K$, even though our detection limit is brighter than in the G305.3+0.2 cluster. The turnover at $K=15.5$ appears to be genuine rather than an artifact of completeness. Perhaps lower-mass stars in this cluster are still more deeply embedded in the gas and dust, and thus we observe only the massive objects.

Due to the small number of stars detected in this cluster $(N=25$, only 7 of which were detected in $H$ ) and to the early evolutionary stage of the objects, we did not attempt to determine an IMF for this cluster or to place objects on the ZAMS. While the individual objects we observed in the G353.4-0.36 cluster were intriguing and worthy of further study, we cannot analyze the cluster as a whole because there are so few objects.

This cluster is a very promising target for study at other wavelengths more suited than the NIR to the study of YSOs and even earlier stages of star formation; the methanol masers and likely presence of massive YSOs suggest that several
stages of massive star formation can be studied in this region.

### 3.5 Summary

We present NIR images and spectroscopy of two young stellar clusters near radio sources G353.4-0.36 and G305+00.2. Our $K$-band spectrum of the brightest cluster star in the G305+00.2 cluster show it to be an O5Ib-O6Ib star. Although the range of luminosities of supergiants prevents us from determining an exact distance, this identification suggests a larger distance than radio distance to the nearby methanol masers (Walsh et al., 1997) of 3.3 kpc . We also obtained spectra of early two B stars in the cluster. There was no nebular emission present in the G305+00.2 cluster, though a ridge of nebular emission, coinciding with $8 \mu \mathrm{~m}$ emission and masers, is present $\sim 1^{\prime}$ away and may indicate sequential star formation, with the masers and gas indicating ongoing star formation and the cluster the result of earlier star formation. We computed the KLF and IMF of this cluster, and found them to be steeper than that reported for most young clusters ( $\Gamma=-0.98 \pm$ 0.2 for the more reliable CMD-based method) but generally consistent with the Salpeter value. We find that computing the IMF based only on a single color of photometry is prone to systematic errors when differential extinction and fieldstar contamination are significant.

Our $K$-band spectra of two of the three stars we observed in the G353.4-0.36 cluster were featureless, while the other showed CO absorption, which is consistent either with a cool foreground giant or a YSO. The absolute magnitudes derived based on the distance to the radio sources are too bright for these objects to be solar-mass YSOs. None of the objects were detected in our $J$-band photometry, making identification as YSOs based on NIR excess impossible. They remain candidate massive YSOs , and observations at other wavelengths are needed
to make a positive identification. The images of this cluster showed a region with intense nebular emission embedded in a very dark cloud where earlier stages of star formation are progressing.

## CHAPTER 4

## Northern 2MASS-Selected Young Stellar Clusters: Photometry and the Initial Mass Function

The contents of this chapter were previously published in Leistra et al. (2006).

### 4.1 INTRODUCTION

Embedded clusters are increasingly recognized as vital sites of star formation for both low- and high-mass stars. Recent studies indicate that clusters may account for $70-90 \%$ of star formation and that embedded clusters (those still partially or fully enshrouded in their natal molecular cloud) may exceed the number of older, non-embedded open clusters by a factor of $\sim 20$ (Elmegreen et al., 2000; Lada \& Lada, 2003). The stellar content of embedded clusters within well-known star formation regions can now be probed, where high extinction $\left(A_{V} \gtrsim 10\right)$ prohibits studies at optical wavelengths. The IMF of such clusters has generally been found to be consistent with a Salpeter value with a slope of $\Gamma=-1.35$ (e.g Okumura et al., 2000; Blum, Damineli, \& Conti, 2001; Figuerêdo et al., 2002) although outliers have been found as well, generally with flatter slopes than the Salpeter value (e.g. Porras et al., 1999).

Although near-infrared (NIR) spectral classification of massive stars is possible (Hanson et al., 1996), in most cases determinations of the IMF from NIR data rely heavily, if not exclusively, on photometry and use spectroscopy only to obtain reliable masses of the few brightest and most massive stars in a cluster if at all. Since these results thus depend on stellar evolutionary models as well as details of the handling of extinction, this raises concerns about to what extent the IMF depends on the methodology employed. Massey (2003) cites the example
of NGC 6611, where two separate analyses of the same data (Hillenbrand et al., 1993; Massey et al., 1995) using different treatments of extinction produced IMFs differing by more than the formal $1 \sigma$ errors would suggest $(\Gamma=-1.1 \pm 0.1$ and $\Gamma=-0.7 \pm 0.2$.) Similarly, the IMF for the G305+0.2 embedded cluster differs by more than the errors between the value derived from the K luminosity function (KLF; $\Gamma=-1.5 \pm 0.3$ ) and that derived using the color-magnitude diagram ( $\Gamma=-0.98 \pm 0.2$ ) (Leistra et al., 2005). Claims that variations in the IMF exist, whether based on individual extreme clusters such as the Arches or a general analysis of the data (Scalo, 1998), must thus be handled with care to compare only results based on similar methodology.

The final release of the Two Micron All Sky Survey (2MASS) has fostered studies (e.g. Dutra \& Bica, 2000, 2001; Bica et al., 2003; Ivanov et al., 2002) which can probe a much larger portion of the Galaxy for previously unknown embedded stellar clusters and significantly increased the number of known embedded clusters. A compilation of some of these results along with previously known embedded clusters is presented in Porras et al. (2003) who find that $\sim 80 \%$ of the stars in their sample are found in "large clusters" of more than 100 stars, despite the rarity of such clusters. However, these studies are not foolproof, and compilations of "embedded clusters" based purely on the 2MASS data without followup must be treated with caution. The studies based solely on stellar density criteria (e.g. Dutra \& Bica, 2000, 2001) have been found in followup work by different groups (e.g. Dutra et al., 2003; Leistra et al., 2005; Borissova et al., 2005) to have only about a $50 \%$ success rate toward the inner Galaxy where the stellar background is high. We have performed an independent search of the 2MASS archive, searching the Point Source Catalog for regions of higher stellar density than the background (determined locally within a $5^{\prime}$ radius) which are redder in
$H-K$ than the local field. This selects for embedded clusters, with the color criteria helping to eliminate chance superpositions and regions of low extinction. A large background radius and the use of color selection are critical to the automated identification of embedded clusters, but even color selection can fail in regions of high background stellar density, where patchy extinction can mimic clusters.

In $\S 4.2$ we present the observations and data reduction for four embedded clusters found in the 2MASS Point Source Catalog, in $\S 4.3$ we present the $K$-band luminosity functions (KLF) and initial mass functions (IMF) for the clusters, and in $\S 4.3 .5$ we address the issue of systematic differences between different methods of deriving the IMF for embedded clusters, for our clusters as well as IMFs from the literature.

### 4.2 OBSERVATIONS \& DATA REDUCTION

We selected five young stellar cluster candidates from the 2MASS Point Source Catalog based on color and density criteria as described in (Leistra et al., 2005). We selected regions with a higher stellar density than the locally defined field with redder $H-K$ color than the field to select embedded cluster candidates. The cluster candidates were observed using the PISCES instrument (McCarthy et al., 2001) on the 6.5 m MMT on Jan10-11, 2003. PISCES uses a $1024 \times 1024$ HAWAII array with a platescale of $0^{\prime} .185 /$ pixel on the MMT, providing a $3^{\prime} \times 3^{\prime}$ field of view. Images of all cluster candidates were obtained in $J, H$, and $K$ filters to a limiting magnitude of $J=19.5, H=18.5, K=18$. Four of our candidates (all except the one near Sh 2-217) were independently identified as cluster candidates by Bica et al. (2003), who used criteria based only on stellar density without color considerations. Our deeper, higher-resolution images suggest that four of these
candidates are genuine stellar clusters, while the fifth, near Sh 2-258, contains only a few stars in the PISCES images and could be either a very small cluster or a chance superposition. We present our results for the four confirmed clusters in this paper. Seeing conditions when the images were obtained were variable, ranging from $0^{\prime \prime} .5$ to $1^{\prime \prime} .1$ at $K$.

All images were reduced and combined using IRAF routines. The distortions of the PISCES camera were mapped by imaging the globular cluster NGC 4147 and mapping the observed locations to the USNO-B known coordinates, then constructing a transformation function using the IRAF task geomap and correcting the images using IRAF geotran. There are too few USNO-B stars in the heavily extincted regions we observed to provide suitable distortion corrections from the fields themselves. The individual images, taken with a spiral dither pattern, were then combined. We fit a PSF to each image using the IRAF task $p s f$, allowing the PSF to vary across the field to compensate for residual distortions. Photometry was then done using IRAF-DAOPHOT.

Photometric calibration was performed using the 2MASS magnitudes of field stars. Stars used for calibration were selected to be well-separated from other stars and from nebulosity in the PISCES images to ensure that they were uncontaminated in the lower-resolution 2MASS images, and to have magnitudes bright enough to have good photometry in 2MASS $(K<14)$. We chose to use a relatively large number of calibration stars rather than selecting the few most isolated stars to reduce effects of potential variability and photometric outliers among the calibration stars. The scatter in the photometric calibration derived from comparison to 2 MASS is the dominant source of photometric error, contributing two to three times the measurement errors as reported by DAOPHOT. DAOPHOT errors were $\sim 0.03$ magnitudes while the calibration uncertainties
were $\sim 0.1$ magnitudes. Quoted errors in the 2MASS photometry were negligible, with most stars having an error of $\pm 0.003 \mathrm{mag}$ or less in all bands. Thus, the quoted error should be considered an overestimate when considering the relative photometry of stars within either cluster; the calibration errors from comparison to the 2MASS photometry will shift all our measurements by the same amount. No trend with location on the chip was observed in the calibration for any of the clusters, though the scatter between the PISCES and 2MASS magnitudes becomes significant in the outermost 15 "; we thus exclude these sources from the analysis.

### 4.3 Analysis

### 4.3.1 Sh 2-217 Cluster

We present a $K$-band image of the cluster near Sh2-217 in Figure 4.1. The cluster is nearly circular in projection and is quite dense; even in the highest-resolution individual pointings we obtained it suffers from crowding in the central regions. The cluster extends over most, if not all, of the field of view. We present the $K$ image rather than a color frame because the seeing was significantly better at $K$. This cluster was analyzed in the NIR by Deharveng et al. (2003), who discuss the large uncertainties in the distance to Sh2-217. Based on the Lyman continuum fluxes from the main exciting star of Sh2-217 (located several arcminutes outside our field of view; the cluster is located on the periphery of the H II region) they adopt a distance of $5.0 \pm 0.8 \mathrm{kpc}$., which is consistent with the kinematic distance to the associated molecular gas. The cluster is coincident with a peak in the $8 \mu \mathrm{~m}$ emission as measured by the MSX mission, suggesting dust is still present in the cluster.

The $K$ vs $H-K$ color-magnitude diagram is shown in Figure 4.2. The stel-


Figure 4.1: $K$-band image of the Sh 2-217 cluster center (North=up, East=left). Image is approximately $120^{\prime \prime}$ on a side. The stellar density does not plateau in our entire $3^{\prime} \mathrm{FOV}$, suggesting the cluster outskirts continue at least to the edge of our image.
lar density (determined in $K$ where the seeing was best) does not plateau in the $3^{\prime} \times 3^{\prime}$ field of view, suggesting that the true field-star density level has not been reached, and cluster stars are still present out to the edges of the field. As a result, in order to correct for foreground contamination, we selected an adjacent field of the same size from 2MASS to use as a comparison field. We assumed the luminosity function of the field to be the same as that of the outer portions of the cluster (excluding the inner regions to minimize the results of crowding and mass segregation) in order to extrapolate from the limiting magnitude of 2MASS to that of our images. We then binned the field stars by $K$ and $H-K$ with a bin size of 0.5 magnitudes and randomly selected the appropriate number of stars for removal from each bin in the cluster region. This is similar to the procedure employed by, among others, (Blum et al., 2000) (who do not describe an extinction correction; this lack of correction is equivalent to assuming a common extinction) and Figuerêdo et al. (2002), and is the method we employed in (Leistra et al., 2005). The resulting statistically corrected CMD is shown in Figure 4.3. A total of 62 stars were removed in this procedure, out of an initial total of 236. The fairly wide distribution in $H-K$ for cluster stars is likely due to a combination of factors, notably an actual spread due to differential extinction to different regions of the cluster and to the greater influence of crowding in $H$ (where the seeing was poorer). Although we find that individually correcting extinctions generally provides a superior estimation of the IMF compared with using only the $K$ data and a single extinction for the cluster as a whole, in this situation the lower quality (in particular the poorer seeing and consequently more severe crowding) of the $H$-band data leads us expect that the $K$ luminosity function (KLF) will produce a better estimate of the IMF for this cluster than the CMD will. We have previously compared these two methods of determining the IMF for embedded clusters in

Leistra et al. (2005); in that case, we found they gave different results, with the "CMD" method (which we anticipate will be more reliable in most cases, especially where variable extinction is present) yielding a flatter slope.

### 4.3.1.1 The KLF

In order to obtain a robust determination of the KLF and the IMF, we need to determine the completeness of our data. To do this, we performed artificial star tests. We inserted five artificial stars, each of the same magnitude, at a time into the cluster region, then ran IRAF-DAOPHOT with the same parameters as we used for the initial analysis. This procedure was repeated 50 times for each magnitude bin $(\Delta m=0.5)$, for a total of 250 artificial stars added in each bin in $H$ and in $K$. The stars were added in small numbers at a time to avoid having the artificial stars significantly change the crowding characteristics and thus influence the completeness. The artificial star tests indicate a high level of completeness down to $K=17.5$. The actual completeness is most likely slightly lower, since in the crowded central region of the cluster our method may produce false positives when the artificial star is placed on top of a real star of approximately the same magnitude. Despite this concern we have used the calculated incompleteness in correcting the KLF; however, we have excluded the $K=17.5$ bin from consideration, both because this effect will be most pronounced at faint magnitudes, and because statistical uncertainties in the incompleteness will be significant.

Knowing our incompleteness, we can calculate the KLF for the cluster. Figure 4.4 shows both the uncorrected and completeness-corrected versions of the KLF, as derived from all sources detected in $K$. The slope of the KLF is $0.31 \pm 0.04$, with no extinction correction applied.


Figure 4.2: $K$ vs. $H-K$ color-magnitude diagram for the cluster near Sh2-217. Average DAOPHOT error bar is the size of the plot symbols or smaller. Overall uncertainties including calibration (which include terms that will not affect the relative position of the points) are indicated by the symbol in the upper right.


Figure 4.3: Statistically-corrected $K$ vs. $H-K$ color-magnitude diagram for the cluster near Sh2-217. Statistical correction was done based on the 2MASS Point Source Catalog for an adjacent field, extrapolating to the our limiting magnitude based on the luminosity function.


Figure 4.4: $K$ luminosity function for the Sh2-217 cluster. (Solid is measured; dashed is corrected for incompleteness.) Since the cluster fills the field of view and the comparison 2MASS data do not go as deep, we do not show a field sample for comparison.

### 4.3.1.2 The IMF

We derived an IMF for the Sh 2-217 cluster by two methods. For both methods we used a distance to the cluster of 5 kpc (Deharveng et al., 2003). In general we believe that the "CMD method" for deriving the IMF, which uses individuallyderived extinctions for each star in the cluster, to be more reliable than the "KLF method" which assumes a common extinction to all stars in the cluster, since variable extinction is frequently apparent in the NIR images of embedded clusters. However, in this case our $K$ data is superior to our $H$ data due to the difference in the seeing, which was $\sim 0 .^{\prime \prime}$ in $K$ and $\sim 1.2^{\prime \prime}$ in $H$. This suggests that despite the general drawbacks of the KLF method it may be preferable in this situation; including the $H$-band data adds nothing if it is of poor quality. At the least, using both methods will provide information on potential systematic effects in the IMF determination that depend on the methodology used.

The KLF method of determining the IMF is sometimes employed even when multi-color photometry and spectra are available (e.g. Blum et al., 2000), and is simply a transformation from $K$ magnitude bins to mass bins. The major problem with this is that of variable extinction, which for many embedded clusters is significant even in $K$. To make this transformation, we first correct the observed $K$ for distance and extinction. Without spectra, we cannot obtain a precise estimate for the extinction; instead, we compute an average extinction correction based on the observed $J-H$ and $H-K$ colors of the brighter stars. Since there is little difference in the intrinsic $H-K$ color of stars of spectral type F5 and earlier, this estimate of an average extinction is not sensitive to minor errors in the distance estimate. Using the stellar evolutionary models of Meynet \& Maeder (2003) for solar metallicity, we relate the mass for each track to an absolute $K$ magnitude for a star on the ZAMS. We transformed $L_{b o l}$ to $K$ using the bolometric corrections from Vacca, Garmany, \& Shull (1996) for the early spectral types and Malagnini et al. (1986) for later spectral types. We then use the intrinsic $V-K$ colors from Bessell \& Brett (1988) for A-M stars and from Wegner (1994) for O and B stars. Finally we interpolate linearly between the masses available on the evolutionary tracks to find the masses corresponding to our magnitude bins, and fit a power law to the resulting mass function. The IMF slope we derive by this method is $\Gamma=-2.7 \pm 0.25$, excluding bins corresponding to $K>17.5$ where incompleteness becomes significant. The slope we fit to the KLF itself for sources detected in both $H$ and $K$ is $0.35 \pm 0.04$.

We have previously used multi-color photometry in conjunction with near-IR spectroscopy of the brightest stars to determine the IMF for embedded clusters (Leistra et al., 2005). Although we do not have spectra in this case, we can still derive extinctions for individual stars based on their near-IR colors. We use the


Figure 4.5: $K$ vs. $H-K$ color-magnitude diagram for the Sh2-217 cluster adjusted to a distance of 5 kpc . The ZAMS converted to observed quantities as described is overplotted.
same evolutionary models and conversions from theoretical to observed quantities described for the KLF method. The ZAMS derived from the evolutionary tracks of Meynet \& Maeder (2003) is overplotted on the distance and extinctioncorrected CMD in Figure 4.5. The ZAMS lies in the middle of the distribution of stars due to the method used to estimate an average extinction. The scatter around the ZAMS is rather large, and is likely due to a combination of variable extinction in the cluster region and poor photometry in $H$, especially in the central portion of the cluster.

Since extinction appears to vary across the cluster, we impose an extinction limit on the sample used for the CMD computation of the IMF. Since the most massive stars can be seen to greater extinction than less massive stars, neglecting to impose this constraint will produce an overly flat IMF. Thus we need to simultaneously limit our sample by extinction and by mass. We use a mean extinction to the Sh2-217 cluster of $A_{V}=9$ mag as determined by individually de-reddening sources until they reach the main sequence. With the exception of a few extreme outliers that most likely suffer from poor photometry, most sources with higher extinctions have $A_{V}<15$. At a distance of 5 kpc with our limiting magnitudes, we can observe stars earlier than G4 to an extinction of 9 mag and earlier than F0 to an extinction of 15 mag .

At a distance of 5 kpc with our limiting magnitude, an extinction of $A_{V}=9$ limits us to G4 and earlier stars, while $A_{V}=15$ limits us to F0 and earlier. We select $A_{V}=10$ and G 2 as our limits; stars with higher extinction or later spectral type cannot be detected over the entire range of mass or extinction included and thus are excluded. Approximately 34 stars have a higher extinction than this, including five with extreme calculated extinctions $\left(A_{V}>50\right)$ that most likely suffer from poor photometry and have unrealistic colors. This extinction limit will also have the effect of excluding stars with $K$-band excess from the IMF determination, since such sources would appear to be at high extinction. This will tend to push the IMF to flatter values, since lower-mass sources spend more time as IRexcess objects and thus are more likely to be ruled out by this criterion. However, we consider this to be a better approach than including the sources since: 1. the number of sources detected in all three bands in this cluster showing near-IR excess is small, suggesting such sources will not significantly influence the mass; 2. the majority are of a low enough mass to fall below the completeness limit, and
are thus excluded anyway regardless of the method; 3. including them (since not all sources are detected in $J$ ) would weaken the extinction limit and tend to again force the IMF to flatter values. In clusters where IR-excess sources dominate we do not fit an IMF (see Section 4.3.2.1 for further discussion).

Individual extinctions are derived by moving the stars along the direction of the reddening vector until they lie on the ZAMS. Once the stars have been corrected individually for extinction, they are placed in mass bins. We fit a power law to the data, excluding masses $<1.1 M_{\odot}$ from the fit since they cannot be seen over the entire range of extinctions in the cluster. The IMF slope we derive by this method is $\Gamma=-1.61 \pm 0.2$. As for the KLF method, the quoted errors represent only the formal errors in the fit and should be considered an underestimate.

The difference between these two values for the IMF emphasizes that formal statistical errors significantly underestimate the true uncertainties in the IMF. In this case, as was the case in Leistra et al. (2005), the CMD method gives a noticeably flatter result than the KLF method. Clearly the individual extinction correction leads to the conclusion that more massive stars are present than an average correction does. This could be due to the effects of mass segregation, or to an incorrectly chosen extinction limit (so that we truly are sampling massive stars more completely than lower mass stars). It is difficult to understand which of these effects is most important without obtaining spectra for a significant number of stars in the cluster.

This cluster is also analyzed by Porras et al. (1999), who use a slightly different distance ( 5.8 kpc ) and extinction $\left(<A_{V}\right\rangle=5.3 \pm 3.7$ ) and derive an IMF slope of $\Gamma=-0.59$ based on 54 sources using the $J$ vs $J-H$ CMD to individually correct extinctions and compare with a theoretical JLF. This is a significant discrepancy from our result with either method. A number of factors may con-
tribute to this difference. The most significant, however, is likely to be due to a different choice of cluster boundaries. Their quoted cluster radius corresponds to only $50^{\prime \prime} 9$, compared to ours of $\sim 80^{\prime \prime}$. This suggests that their IMF will be more weighted toward the cluster core than ours. The value they quote for a field + cluster IMF is $\Gamma=-2.71 \pm 0.24$, much steeper and in fact quite close to the value we obtain using the KLF. If the cluster suffers from mass segregation, as we would expect given that it is observed even in very young clusters (e.g. the Arches cluster (Stolte et al., 2002)), we expect the core IMF to be quite flat. When we re-derive the IMF for this cluster using their radius with our data, we derive an IMF of $\Gamma=-1.55 \pm 0.22$, statistically indistinguishable from our original result. Porras et al. (1999) do not comment on issues of confusion or field star contamination, so we cannot evaluate how much of an effect it is likely to have on their result; we expect crowding to be a more significant issue, and a misidentification of blended sources as single stars by Porras et al. (1999) could account for their finding a steeper IMF than we do using the same method.

### 4.3.2 IRAS 06058+2158 Cluster

We present a three-color composite of the cluster near the IRAS source 06058+2138 in Figure 4.6. Bik et al. (2005) obtained VLT spectra in $K$ of several NIR point sources near the IRAS point source, which is located near the center of the cluster, and identified two candidate massive YSOs and an embedded early-B star. The spectrophotometric distance they derive from the B star is $1.0-1.5 \mathrm{kpc}$. This cluster is much more heavily embedded than the Sh 2-217 cluster, with significant nebular emission and prominent dust lanes. Numerous OH and methanol masers have been detected in this region, which along with the IRAS point source suggest ongoing star formation. (see, e.g., Caswell et al. (1995), Szymczak et al. (2000)). A peak in the $8 \mu \mathrm{~m}$ emission is observed in the MSX data, extending
from the region of NIR nebulosity to the southeast to the isolated bright source.
The embedded cluster here is described in the compilation of Lada \& Lada (2003), who quote a distance of 1.5 kpc (Carpenter et al., 1993). However, Hanson, Luhman, \& Rieke (2002) describe a UCHII region associated with the IRAS point source, and quote a distance of 2.2 kpc (Kömpe et al., 1989), and Bik et al. (2005) obtain a spectroscopic distance of $1.0-1.5 \mathrm{kpc}$. With no a priori reason to prefer one distance over the other, and no uncertainties associated with either, we use the average distance of 1.5 kpc for our analysis.

### 4.3.2.1 The KLF

We present color-color and color-magnitude diagrams for this cluster in Figures 4.7 and 4.8. The "cluster region" was defined to coincide with the extent of the nearIR nebulosity, and field stars were statistically corrected as described in 4.3.1. The CMD shows objects spanning a range of extinctions, with relatively few objects with colors consistent with unextincted main-sequence stars remaining in the statistically corrected data. Since the cluster in this case did not fill the field of view, we were able to use the data from the non-cluster portions of the region as our field, eliminating the need for an off-source 2MASS field and extrapolation based on the KLF. The $J-H$ vs. $H-K$ color-color diagram shows that the majority of sources in the cluster region exhibit $K$-band excess and fall in the region populated by reddened CTTS and YSOs, suggesting they may be pre-main-sequence objects. Because the amount of $K$-band excess is affected by many factors (see, e.g.,Meyer, Calvet, \& Hillenbrand (1997)), deriving masses for these objects is difficult. We thus derive only a KLF for this cluster, and do not convert it to an IMF or derive an IMF from the CMD. We determine and correct for our incompleteness as in 4.3.1. The KLF we derive using a statistically-corrected sample of all sources detected in $K$, with no attempt to correct for extinction due to the un-


Figure 4.6: Color composite ( $J=$ blue, $H=$ green, $K=$ red) of the IRAS 06058+21 cluster region (North=up, East=left). Image is approximately $140^{\prime \prime}$ on a side.
certainty of the intrinsic $H-K$ color of the YSOs that are present, has a slope of $0.30 \pm 0.03$. If only sources detected in $H$ are included, the KLF declines for $K>14.5$ since the high extinction means the fainter objects are less likely to be detected in $H$.

### 4.3.2.2 Pre-Main-Sequence Objects

A total of 37 out of 58 sources ( $63 \%$ ) detected in all three bands show a K-band excess in the color-color diagram, suggesting they are pre-main sequence objects. This is a lower limit on the actual pre-main-sequence fraction of the cluster, since objects with a sufficiently high IR excess may be detected in $K$ but not in $J$ or $H$. A total of 49 sources were detected in $K$ within the cluster region that were undetected in $J, H$, or both. Adding in these sources would give a PMS fraction of $80 \%$. The latter figure is an upper limit, since some of the $K$--only detections are likely to be knots of nebular $\operatorname{Br} \gamma$ emission or heavily extincted background stars. Comparing these values to the near-IR excess fraction of embedded clusters of known ages presented in Haisch, Lada, \& Lada (2001), we conclude that the age of the IRAS 6058 cluster is less than 3 Myr.

We observe a few sources with colors even redder than the reddened extension of the CTTS locus. Meyer, Calvet, \& Hillenbrand (1997) observe sources with similar colors, and suggest re-radiation by an extended envelope as an explanation.

### 4.3.3 IRAS 06104+1524 Cluster

The near-IR cluster image (Figure 4.9) of the cluster near IRAS 06104+1524 shows a clear separation into two subclusters separated by $\sim 2^{\prime}$. The southwest subcluster is dominated by two closely spaced bright sources while the northeast subcluster is denser and is not dominated by a single object. A ridge of marginally


Figure 4.7: Statistically field-star corrected $K$ vs. $H-K$ color-magnitude diagram for the cluster near IRAS 06058+2158. Significant and variable extinction is evident in this cluster. Typical errorbars are the size of the plot symbols. Overall uncertainties including calibration (which include terms that will not affect the relative position of the points) are indicated by the symbol in the upper right.


Figure 4.8: $J-H$ vs. $H-K$ color-color diagram for the cluster near IRAS $06058+2158$. The lines (parallel to the reddening vector) delineate the possible location of reddened main-sequence stars. Over half of our sources show $K$-band excess; their colors are not consistent with reddened main-sequence sources. Typical errorbars are the size of the plot symbols. Overall uncertainties including calibration (which include terms that will not affect the relative position of the points) are indicated by the symbol in the upper right. Points to the left of the reddening lines lie in crowded regions and may suffer from confused photometry. Points more than $3 \sigma$ to the left of the reddening line were excluded from the IMF determination.
higher density than the surrounding field appears to lie between the subclusters, though it is not apparent whether this is a real feature. The MSX $8 \mu$ m image similarly shows two separate peaks, with no indication of a connection. These are treated as separate clusters by Bica et al. (2003), and there are IRAS point sources associated with each of them (IRAS $06104+1524$ and IRAS 06103+1523, respectively). The IRAS point sources both have kinematic distances of 3.5 kpc (Wouterloot \& Brand, 1989), suggesting the subclusters are related. The radio kinematic distances derived to the two sources are the same, and the angular separation is comparable to the size of either subclump; in addition, a slight overdensity of stars can be seen in the $K$ image. These factors suggest that, at the very least, these clusters are related; they may differ in age, but they are likely to be part of the same general star-formation event. We see no difference between the two apparent in either the CMD or the color-color diagram (Figures 4.10 and 4.11); if they do differ in age, it is beyond the ability of our data to discern. Due to this apparent association and the small number of stars in each cluster, we analyze the two together as a single cluster. The CMD after statistical correction and adjusting to a distance of 3.5 kpc (Wouterloot \& Brand, 1989) is shown in Figure 4.12.

Using the color-magnitude diagram based method of deriving an IMF as described above, for a limiting extinction of $A_{V}=25$, we find an IMF slope of $\Gamma=-0.9 \pm 0.25$. Using the KLF method (with no extinction correction, to mimic the results of a study with only single-color photometry available) we arrive at $\Gamma=-2.6 \pm 0.3$. Even after allowing for the errors to be larger than quoted due to uncertainty in the photometry and the conversion to mass, these two slopes are inconsistent with each other, suggesting that systematic effects in one or both methods dominate over the statistical errors.


Figure 4.9: Color composite ( $J=$ blue, $H=$ green, $K=$ red ) of the IRAS 06103 +1523 / 06104+1524 cluster(s) (North=up, East=left). Image is approximately $120^{\prime \prime}$ on a side.


Figure 4.10: $K$ vs. $H-K$ color-magnitude diagram for the IRAS 06104+1524 / 06103+1523 cluster(s) adjusted to a distance of 3.5 kpc . Circles: NE cluster. Asterisks: SW cluster. There is no difference in the CMD apparent for the two clusters, so we treat them as one to improve the statistics. Typical errorbars are the size of the plot symbols. Overall uncertainties including calibration (which include terms that will not affect the relative position of the points) are indicated by the symbol in the upper right.


Figure 4.11: $J-H$ vs. $H-K$ color-color diagram for the IRAS 06104+1524 / $06103+1523$ cluster(s). Typical errorbars are the size of the plot symbols. Overall uncertainties including calibration (which include terms that will not affect the relative position of the points) are indicated by the symbol in the upper right.


Figure 4.12: Statistically corrected $K$ vs. $H-K$ color-magnitude diagram for the IRAS $06104+1524 / 06103+1523$ cluster(s) adjusted to a distance of 3.5 kpc . Typical errorbars are the size of the plot symbols.Overall uncertainties including calibration (which include terms that will not affect the relative position of the points) are indicated by the symbol in the upper right.

### 4.3.4 Sh 2-288

The near-IR cluster image (Figure 4.13) of the cluster near Sh2-288 shows a cluster with a dense core, crossed near the center by a dust lane. The center of the cluster is unresolved in our images, taken with a seeing of $0.7^{\prime \prime}$. This region was previously identified as an embedded cluster by (Dutra \& Bica, 2001). In their catalog of outer-Galaxy HII regions, Rudolph et al. (1996) quote widely disparate distances, with a radio kinematic distance of 7.2 kpc and a photometric distance of 3.0 kpc (Brand \& Blitz, 1993). The kinematic distance would make the sources we observe (several with $K<12$ ) extremely massive, it is thus far more likely that the photometric distance is correct, and we have used the photometric distance for our analysis.

The $8 \mu \mathrm{~m}$ image from the MSX mission shows a peak coincident with the nearIR nebulosity; there is not a significant amount of $8 \mu \mathrm{~m}$ emission from regions dark in the NIR. This suggests that there are not a significant number of sources so deeply embedded that they cannot be seen in $K$ present in this cluster.

The $K$ versus $H-K$ color-magnitude diagram of the cluster near the HII region Sh2-288 (Figure 4.14) clearly shows the effects of variable extinction; the stars separate into two groups, one nearly unextincted and one with $\sim A_{V}=5$. We correct for field star contamination using the region of the field outside the cluster region as described above. The $J-H$ versus $H-K$ color-color diagram (Figure 4.15) shows few stars separated from the main sequence locus by more than $2 \sigma$, suggesting that most stars in this cluster are on the main sequence. Extreme outliers in the color-color diagram were inspected individually; in general, they lie in the crowded central region of the cluster and most likely suffer from poor photometry due to the different PSFs in $H$ and $K$ that resulted from variations in seeing. Such sources were excluded from analysis of the KLF and the IMF.


Figure 4.13: Color composite ( $J=$ blue, $H=$ green, $K=$ red) of the Sh 2-288 cluster region (North=up, East=left).

Additionally, the brightest source in the cluster, which lies in the most crowded central region and has a FWHM slightly broader than most sources in the field, was excluded since it is quite likely to be a blend of multiple sources. We consider that the effects on the IMF are likely to be worse if a blend is included than if any single star, even the most massive, is excluded.

Using the photometric distance of 3.0 kpc from (Rudolph et al., 1996), we derive an IMF from the KLF of $\Gamma=-1.95 \pm 0.62$. To better compare the two methods of deriving the IMF, we included only those sources which were also detected in $H$, so that the same dataset would be used for both the KLF and CMD methods of deriving the IMF. We individually de-reddened sources in the $H-K$ CMD until they were on the main sequence, imposing an extinction limit as before, and derive an IMF of $\Gamma=-1.62 \pm 0.65$. Given the large uncertainties, these results are entirely consistent. The better agreement may be because the extinction bias is less in the latter case.

### 4.3.5 Comparison of Methods for IMF Determination

A summary of the IMFs derived for the three clusters without significant numbers of pre-main-sequence stars and the similar results from (Leistra et al., 2005) is shown in Table 4.1. In each case, there is a significant difference between the IMF derived from the KLF and that derived from the CMD. Does this reflect only uncertainties, or is one method in general more reliable than the other? A simple analysis would suggest that the CMD method is more reliable, simply because it uses more information; the extinction clearly varies across many embedded clusters (of those analyzed here, most notably Sh 2-288 and IRAS 06058+2158), and accounting for this should provide a more robust estimate of the true IMF. We cannot say for certain that this is the case, however, without obtaining spectra for most of the stars in each cluster, so that we can classify them spectroscopically


Figure 4.14: Statistically corrected $K$ vs. $H-K$ color-magnitude diagram for the Sh2-288 cluster. The bright central source, which we have identified as a blend of two or more stars, has been removed as well. Typical errorbars are the size of the plot symbols. Overall uncertainties including calibration (which include terms that will not affect the relative position of the points) are indicated by the symbol in the upper right.


Figure 4.15: Statistically corrected $J-H$ vs. $H-K$ color-color diagram for the Sh2-288 cluster. The lines (parallel to the reddening vector) delineate the possible location of reddened main-sequence stars and of reddened T Tauri stars. Typical errorbars are the size of the plot symbols. Overall uncertainties including calibration (which include terms that will not affect the relative position of the points) are indicated by the symbol in the upper right.

Table 4.1. IMFs for embedded stellar clusters

| Cluster | IMF: Common $A_{V}$ | IMF: Individual $A_{V}$ | Distance (kpc) | Source |
| :--- | :--- | :--- | :--- | :--- |
| Sh2-288 | $\Gamma=-1.95 \pm 0.82$ | $\Gamma=-1.62 \pm 0.5$ | 3.0 | This work |
| IRAS 06104+1524 | $\Gamma=-2.49 \pm 0.3$ | $\Gamma=-1.38 \pm 0.6$ | 3.5 | This work |
| Sh2-217 | $\Gamma=-2.7 \pm 0.25$ | $\Gamma=-1.61 \pm 0.2$ | 5.0 | This work |
| G305.3+0.2 | $\Gamma=-1.5 \pm 0.3$ | $\Gamma=-0.98 \pm 0.2$ | 4.0 | Paper I |

and obtain individual masses. We note that in each case, the IMF we derive from the CMD by individually correcting the extinction for each object is flatter than that we derive from the KLF by assuming a single extinction for the entire cluster. This suggests that more massive stars may preferentially lie in more heavily extincted regions in embedded clusters. Resolving this seeming discrepancy would require spectra for a large number of cluster members, so that the IMF derived from spectral classification of stars can be compared to that derived via the different photometric methods.

We examine the relation between derived mass and extinction (with the extinction limit imposed) for the Sh2-217 cluster, where we have the most data, in Figure 4.16. Such a relation appears to be present, albeit at low significance. This effect is opposite in sign to what would be expected from massive stars clearing their immediate environment more rapidly than their lower-mass counterparts, but the stars we observe are mostly of intermediate mass, rather than truly high mass, such that their winds are not as significant; the effects of mass segregation, placing the more massive stars at denser regions in the cluster, appear to dominate over the effects of clearing in these clusters. Since the clusters are numerically dominated by low-mass stars, the average extinction will be mostly determined by the average for the low-mass stars, changed slightly by the aver-
age for high-mass stars. If more massive stars are indeed preferentially found at higher extinction as our results suggest, this would mean the majority of (lowmass) stars are over-corrected for extinction by a small amount when using a single value, while a few (high-mass) stars are under-corrected (thus lowering the derived mass) by a large amount; thus, the effects of over-correcting and undercorrecting extinction do not fully cancel out, since the large under-corrections would be more likely to move stars between mass bins than the small overcorrections. If there is no relation between mass and extinction, we would expect these two results to cancel, since the average extinction for low-mass stars would be the same as that for high-mass stars.

### 4.4 Summary

We present NIR images of four embedded clusters in the outer Galaxy. In the case of the cluster near IRAS $06058+2158$ the number of stars with NIR excess indicates a pre-main-sequence fraction between $60 \%$ and $80 \%$ and an age of less than 3 Myr ; the other three clusters show less nebular emission and fewer stars with NIR excess indicating an older age. We compute the IMF for the three clusters dominated by main-sequence stars, in each case using both a KLF-based method relying on a single extinction value for the cluster and using only $K$ band data and a CMD-based method where an individual extinction value is calculated for each star. We found a statistically significant difference between the two values in two of the three cases, prompting us to examine IMF values of embedded clusters from the literature to determine whether systematic effects are at work. We found a significant difference in the mean value of the IMFs for embedded clusters derived from methods that handle extinction individually compared with those that adopt a single value for the extinction. Although a larger sample would help


Figure 4.16: Mean extinction derived from $H-K$ color for sources in each mass bin for the Sh2-217 cluster. An extinction limit of $A_{V}=10$ has been imposed; sources at higher derived extinction are not included in the calculation. More massive sources appear to preferentially lie at higher extinctions, albeit with low significance.
to make this claim more robust, since many of the results come from a single study (Porras et al., 1999) and methodological details of that work could affect the results, we consider it to be significant enough that IMFs obtained by different methods should not be compared in an attempt to search for variations in the IMF from region to region.

Truly reliable IMFs for embedded clusters will most likely require spectra for a large number of stars in the clusters; we are continuing to try to obtain spectra for these sources to better characterize the massive star population and the IMF of these clusters.

## CHAPTER 5

## Conclusions and Future Directions

I summarize the key results of the previous chapters, their relationship to other work, and outline some directions for future progress in this area.

### 5.1 The Initial Mass Function of Embedded Stellar Clusters

In Chapters 2 and 3 I presented the results of near-infrared imaging and spectroscopy of embedded cluster candidates, none of which had been known before the 2MASS Point Source Catalog became available. The clusters ranged in distance from 1.5 to 5 kpc , and covered a range of Galactic environments from the plane toward the Bulge (at a Galactic longitude of 353) to the outer disk. When we had spectra of the stars we were able to classify them using $K$-band spectroscopy following the massive star spectral atlas of Hanson et al. (1996). This enabled us to more precisely determine the masses of a few sources and to obtain a reliable distance to the clusters. When we did not have spectra available we needed to rely on radio-derived distances, although when kinematic distances were uncertain we could constrain the distance from our near-IR photometry. For two of the six clusters, the observed near-IR excess fraction was high, indicating a large number of pre-main-sequence sources. Since the mass-luminosity relation in the near-IR is complicated by the presence of disks in this situation, we were unable to determine an IMF for these clusters. For the four clusters that remained, we derived an IMF via two methods, both frequently used in the literature: individually correcting each source for extinction until it fell on the ZAMS in the $K$ vs. $H-K \mathrm{CMD}$, and by assuming a universal extinction for the cluster, deriving the average extinction to be used from the mean $H-K$ color for the brighter sources.


Figure 5.1: Initial mass functions determined in this work (Leistra et al., 2005, 2006).

Since the intrinsic $H-K$ color is essentially the same for all OB stars, $H-K$ color can be used as a proxy for extinction even when the exact type of the stars is unknown. The resulting IMFs are shown in Figure 5.1. In each case the IMF is consistent with the Salpeter value.

I noticed that in each case, the IMF I derived from the CMD by individually correcting the extinction for each object is flatter than that we derive from the KLF by assuming a single extinction for the entire cluster. This is suggestive of systematic effects, but it is far from conclusive, with a $12 \%$ chance that such an arrangement is coincidental. In order to obtain a larger sample to compare the two methods, I conducted a literature search for IMF slopes for embedded clusters derived via the two methods. While Lada \& Lada (2003) discuss the advantages and disadvantages of the two methods, they do not compare the results
from each. Although few clusters have IMF values via both methods available, if statistical effects are in place we should expect IMF slopes obtained from individually de-reddening stars to average flatter than those obtained from the KLF alone. Several authors derive the IMF by constructing a "corrected KLF" from individually de-reddened sources; we treat these as CMD-method sources. In situations where a form other than a single power law was used to describe the IMF (e.g. a multiple power law), we use the slope of the high-mass end for better comparison to our data.

The literature values, together with my results as presented in Chapters 2 and 3, are presented in Table 5.1. Considered all together, the average IMF slope is $\Gamma=-1.23 \pm 0.3$, while that for the subsample determined with the KLF method is $\Gamma=-1.69 \pm 0.3$ and that for the subsample determined with the CMD method is $\Gamma=-1.02 \pm 0.4$. The mean value is very similar to the Salpeter slope and to the average IMF of open clusters determined from optical photometry and spectroscopy (e.g. Massey (2003); Scalo (1998)).

These results suggest that the difference we observe between the IMF derived with and without individual extinction correction may in fact be systematic, rather than statistical. Individually correcting for extinction produces a flatter slope to the IMF than using a mean value for the entire cluster, though at low significance. If massive stars are found at higher extinctions, which may be observed in a non-extinction limited sample even if the true distribution is otherwise, such an effect would be expected. Such a correlation, if present, would be short-lived; while massive stars are expected from both theoretical (Bonnell et al., 2001) and observational (Stolte et al., 2002) evidence to form in the densest regions of clusters, the winds from massive stars will quickly clear out local dust and gas, placing the most massive stars in low-extinction regions; in the Arches

Table 5.1. IMFs for embedded stellar clusters taken from the literature. Distances quoted are those used by the authors.

| Cluster | IMF | Method | Distance (kpc) | Reference |
| :---: | :---: | :---: | :---: | :---: |
| NGC 1333 | $\Gamma=-1.35 * \pm 0.2$ | CMD | 0.3 | Lada et al. 1996 |
| IC 348 | $\Gamma=-1.35 * \pm 0.2$ | KLF | 0.3 | Muench et al. 1999 |
| Trapezium | $\Gamma=-1.35 * \pm 0.2$ | KLF | 0.4 | Muench et al. 1999 |
| IRAS 23545+6508 | $\Gamma=-1.42 \pm 0.18$ | CMD | 1.9 | Porras et al. 1999 |
| IRAS 06055+2039 | $\Gamma=-0.38 \pm 0.10$ | CMD | 2.0 | Porras et al. 1999 |
| IRAS 04324+5106 | $\Gamma=-0.99 \pm 0.10$ | CMD | 7.2 | Porras et al. 1999 |
| IRAS 06063+2040 | $\Gamma=-0.8 \pm 0.07$ | CMD | 2.8 | Porras et al. 1999 |
| IRAS 02236+6142 | $\Gamma=-0.63 \pm 0.15$ | CMD | 3.0 | Porras et al. 1999 |
| IRAS 02232+6138 | $\Gamma=-0.81 \pm 0.07$ | CMD | 3.0 | Porras et al. 1999 |
| IRAS 01198+6136 | $\Gamma=-0.26 \pm 0.16$ | CMD | 3.0 | Porras et al. 1999 |
| IRAS 22566+5828 | $\Gamma=-0.97 \pm 0.12$ | CMD | 3.5 | Porras et al. 1999 |
| IRAS 02044+6031 | $\Gamma=-0.76 \pm 0.21$ | CMD | 5.8 | Porras et al. 1999 |
| IRAS 04547+4753 | $\Gamma=-0.59 \pm 0.01$ | CMD | 5.8 | Porras et al. 1999 |
| Sh233IR | $\Gamma=-0.72 \pm 0.05$ | CMD | 1.8 | Porras et al. 2000 |
| W51 | $\Gamma=-1.8 \pm 0.1$ | CMD | 7 | Okumura et al. 2000 |
| NGC 3576 | $\Gamma=-1.62 \pm 0.1$ | KLF | 2.8 | Figueredo et al. 2002 |
| NGC 3603 | $\Gamma=-1.35 * \pm 0.2$ | KLF | 7 | Nürnberger et al. 2002 |
| Arches | $\Gamma=-0.7 \pm 0.2$ | CMD | 8 | Stolte et al. 2002 |
| G333.1-0.4 | $\Gamma=-1.1 \pm 0.1$ | CMD | 2.6 | Figueredo et al. 2005 |
| W49A | $\Gamma=-1.30 \pm 0.1$ | CMD | 11.4 | Homeier \& Alves 2005 |
| G305.3+0.2 | $\Gamma=-0.98 \pm 0.2$ | CMD | $\sim 4$ | Leistra et al. 2005 |
| G305.3+0.2 | $\Gamma=-1.5 \pm 0.3$ | KLF | $\sim 4$ | Leistra et al. 2005 |
| Sh2-288 | $\Gamma=-1.62 \pm 0.5$ | CMD | 3.0 | Leistra et al. 2006 |
| Sh2-288 | $\Gamma=-1.95 \pm 0.82$ | KLF | 3.0 | Leistra et al. 2006 |
| IRAS 06104+1524 | $\Gamma=-1.38 \pm 0.6$ | CMD | 3.5 | Leistra et al. 2006 |
| IRAS 06104+1524 | $\Gamma=-2.49 \pm 0.3$ | KLF | 3.5 | Leistra et al. 2006 |
| Sh2-217 | $\Gamma=-1.61 \pm 0.2$ | CMD | 5.0 | Leistra et al. 2006 |
| Sh2-217 | $\Gamma=-2.7 \pm 0.25$ | KLF | 5.0 | Leistra et al. 2006 |
| NGC 6611 | $\Gamma=-1.45 \pm 0.15$ | CMD | 1.7 | Bonatto et al. 2006 |

Cluster, with an age of $\sim 3 \mathrm{Myr}$, an extinction gradient is observed (Stolte et al., 2002) with the center of the cluster, where the massive stars are preferentially located, lying at lower extinction than the outskirts.

We do note that half of the CMD-method results are from a single analysis (Porras et al., 1999), and that problems with the methodology of this particular study could be at least in part responsible for the pronounced effect we observed; in the one situation where this study overlaps with ours, they had a significantly flatter IMF. Despite these caveats, it is clear that attempts to compare IMFs between regions are likely to be meaningful only if the IMFs are determined using the same methodology.

The different results of my two methods for IMF determination, and the suggestion of a similar result in the literature, point to the need for a more careful consideration of methodology in general when comparing photometricallyderived IMFs. In particular, more complete spectroscopically-determined IMFs, such as that of Hillenbrand (1997), which will be more accurate and less errorprone than either photometric method, would allow the true uncertainties and systematic errors of these methods to be explored in more detail.

### 5.2 Discovering Embedded Stellar Clusters using 2MASS: Completeness and Distribution

In Chapter 4 I described the methods used for finding clusters in the 2MASS database, and presented the resulting cluster candidates. We initially performed a targeted search of the 2MASS point source catalog in $3^{\prime}$ regions surrounding targets likely to host embedded star formation, based on their radio or optical properties. This search was the basis for our selection of candidates for followup as described in Chapters 2 and 3. These cluster candidates were selected based on
finding a stellar density enhancement where the stars in the dense region have $H-K$ colors redder than the local average. We found based on our followup observations that the $3^{\prime}$ window was frequently too small to rule out local fluctuations in stellar density due to dust lanes, and modified the size to use for the full-sky search. In the full-sky search we covered the entire sky twice with offset area boundaries. With both searches combined we discovered 87 previously unknown embedded stellar cluster candidates (including those which were simultaneously discovered by other groups using the 2MASS database, but were previously unknown.)

Comparison with catalogs of known embedded clusters and characterization of the different false-positive detections allowed us to determine the completeness of the embedded cluster catalog. We determined based on our success rate in recovering the Lada \& Lada (2003) sample that we could recover only clusters within $\sim 2.5 \mathrm{kpc}$, within which the existing sample including our new detections is $75 \%$ complete, compared with $50 \%$ complete before the recent 2 MASS cluster searches.

The incompleteness of our embedded cluster catalog raises the question of what would be required to construct a complete catalog of embedded clusters in the Galaxy. Such a catalog may never be possible - the density of foreground stars and of gas and dust in the Galactic Plane will always render clusters on the far side of the Galactic Plane inaccessible, so the most we may be able to hope for is a locally-complete sample. Additionally, the high stellar surface density in the Galactic Plane means identifying clusters based on density will be less effective in the plane. This is where most embedded clusters occur; in these region, detection based on density criteria is only feasible for the densest clusters. However, Portegies Zwart et al. (2001) point out that the surface density of the Quintuplet cluster
is less than that of the field (making the Quintuplet+field observed density less than twice that of the field), yet it was detected due to an enhancement in bright stars; real, magnitude-limited surveys will preferentially find young clusters due to their smaller mass-to-light ratio. Thus, deeper surveys may not help find more clusters in the crowded environment of the Galactic Plane and may even prove counterproductive for purely density-based searches, as they will increase the stars detected in the field both by going to farther distances and to lower-mass stars, while the signature massive stars in young clusters would have been detected even in shallower surveys.

The GLIMPSE (Galactic Legacy Infrared Mid-Plane Survey Extraordinaire) Survey, a Spitzer Legacy project to survey the Galactic Plane $\left(|b|<1^{\circ}\right)$ in the 4 IRAC bands from 3.5 to $8 \mu \mathrm{~m}$, together with the GLIMPSE II extension to the innermost Galaxy, offers another method for finding young stellar clusters. Preliminary searches (Mercer et al., 2005) have uncovered 92 previously unknown young clusters in the original GLIMPSE region. Based on the number of embedded clusters in the GLIMPSE area vs. the GLIMPSE II area and assuming a similar spatial distribution for GLIMPSE and known embedded clusters, the number found in GLIMPSE II would be $\sim 5$. The actual number will almost certainly be higher, since IRAC can more easily distinguish pre-main-sequence cluster members from field stars and will be less strongly affected by confusion in regions of high field star density than our near-IR methods. The IRAC bands, which are relatively insensitive to extinction but can easily distinguish pre-main-sequence sources from stars, and the $J H K$ bands which can distinguish between extincted and non-extincted stars but are subject to degeneracies between extinction and pre-main-sequence IR excess, have the potential to be used together to characterize extinction and ages to all the known embedded stellar clusters. The joint use
of the two surveys has enormous potential to provide homogeneous information on the vast majority of known embedded stellar clusters, and is an important future extension of the work done by various groups to discover young stellar clusters using either of the two surveys alone.

It is clear that searches for young stellar clusters using only density criteria are incomplete and subject to false detections, and that including color criteria helps but introduces its own false detections and fails to recover other known clusters. Different methods must be considered. Magnitude-weighted density searches that take advantage of the low mass-to-light ratio of young stellar clusters, are one avenue worth exploring, as are more sophisticated color-based searches that look for a spatially associated group of stars which are consistent with a young, single-age population (i.e. fall on the same isochrone).

In this thesis, I have demonstrated a new method for discovering embedded stellar clusters in the near-infrared that is complementary to the use of stellar density as a sole criterion, and have discovered numerous previously unknown stellar clusters. The embedded stellar clusters, like their better-known counterparts, trace regions known to host star formation, and do not appear in regions without radio or far-infrared indicators of ongoing star formation. Based on followup observations, we conclude that the initial mass functions of these embedded stellar clusters are consistent with the Salpeter value, though we caution that there appear to be statistically significant systematic effects of the method used, and that more spectroscopically-determined IMFs using spectra for a large number of stars are necessary to test the photometric methods employed in this thesis and in general. Finally, I discuss the completeness of the existing embedded cluster sample and offer suggestions for future cluster searches.

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[^0]:    ${ }^{1}$ FIGARO is part of the Starlink software package available at http:/ / star-www.rl.ac.uk/

[^1]:    ${ }^{2}$ On-line data are available from http:/ /www.ipac.caltech.edu/ipac/msx/msx.html

