The Initial Mass Function of Massive Stars in the Local Group

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Abstract. The present-day mass function of massive stars can be measured directly by counting stars as a function of bolometric luminosity; if we understand the star forming history of a region, we can convert this to an initial mass function (IMF). We discuss the complications inherent in this process: because of the high effective temperatures characteristic of massive stars, the bolometric corrections are a very steep function of $T_{\rm eff}$, but at the same time the near-UV and optical colors are quite insensitive to effective temperatures for $T_{\rm eff}$ >30,000°. Hence such diagnostic tools as the visual luminosity function are nearly useless as a measure of the IMF. However, the combination of spectroscopy with photometry provides the means of measuring the IMF of massive stars in the differing physical environments of Local Group galaxies. We find that the IMF slope is essentially Salpeter ($\Gamma \sim -1.3$) in the clusters and associations of the Magellanic Clouds and Milky Way (a factor of 10 spread in metallicity), and in both sparse OB associations and the "super star cluster" R136 (a factor of 1000 in the surface densities of stars). Nevertheless, the significant population of high mass stars that appear to be born in the *field*—presumably of very modest star-forming events—show a considerably steeper slope, with $\Gamma \sim -4$. The evidence suggests that there is no "upper mass cutoff" to the IMF, or at least none that have been detected observationally: the richer and younger a star-forming region, the higher the highest mass star we see, and it is statistics, rather than physics, that limits the mass we see. We discuss the star formation history of OB associations, and find that the formation of intermediate mass stars commences before the formation of high mass stars by several million years. The vast majority of the high mass stars form over a very short period ($\Delta\tau$ <1–2 Myr). Once the high mass stars form, the production of intermediate mass stars usually continues, until the gas is cleared by subsequent supernovae. There are exceptions, however: in the case of very rich clusters (such as R136), the actions of stellar winds is apparently sufficient to clear out the gas, and the production of stars of intermediate mass stops within a million years or so of the formation of the highest mass stars. Finally, we consider what we know directly and indirectly about the IMF of the more distant members of the Local Group: NGC 6822, IC 1613, M 31, M 33, and IC 10. Of these, only the "starburst" galaxy IC 10 shows signs of anything odd in its massive star population.

1. Introduction

The highest mass stars invite many other superlatives. They are the most bolometrically luminous, although, as we will see, not necessarily the most *visually* luminous. They have the shortest life spans. When still near the zero-age main-sequence (ZAMS), they have the highest effective temperatures. Through the actions of their strong stellar winds, and their eventual disruption as supernova, massive stars provide most of the heavy-metal enrichment and energy deposition into the interstellar medium (Abbott 1982).

The last 15 years has seen a stunning improvement in our direct knowledge of the initial mass function of high mass ($> 15\ \mathcal{M}_{\odot}$) stars. Reviews of older observational data were given by Garmany, Conti, & Chiosi (1982) for the Milky Way, and by Fitzpatrick & Garmany (1990) for the LMC. Since that time, however, there has been a true revolution in the instrumentation used by observers: CCDs allow us to obtain good photometry in crowded OB associations, and permit spectroscopy to fainter magnitudes. In addition, the deployment of multifiber positions have allowed us to obtain spectra of a multitude of stars in associations in an efficient manner. *HST* now permits us to obtain spectroscopy of a complete sample of massive stars in a region even as crowded as R136 in 30 Doradus. The advent of IR arrays both for imaging and K-band spectroscopy opens up further avenues in highly reddened associations (for instance, the study done for M 17 by Hanson, Howarth, & Conti 1997), and soon *Hipparcos* results may help us improve distances to OB associations, providing further refinement in our knowledge of the IMF. These observational advances have been coupled with significant improvements in theory: evolutionary tracks by the Geneva group (Maeder & Meynet 1988; Schaller et al. 1992; Schaerer et al. 1993) now provide excellent agreement with the observational data (Massey et al. 1995c), eliminating the so-called main-sequence widening problem (Fitzpatrick & Garmany 1990). Non-LTE stellar atmospheres have also come a long ways since the Auer & Mihalas (1970) models first allowed Conti (1973) to define an accurate effective temperature scale for O-type stars; see the review by Kudritzki & Hummer (1990), and the recent analysis by Puls et al. (1996).

Specific questions that we will address in this review are:

1. How do we determine the initial mass function for massive stars?

2. What is the current data telling us about the slope of the initial mass function, and how this changes as a function of physical conditions (metallicity, density)?

3. What is the upper mass “cut-off” for massive stars, and how does this depend upon physical conditions?

4. Over what sort of time scales do massive stars form in a cluster? What do we know about the history of star formation within a group?

2. Determining the IMF for Massive Stars

2.1. Definitions: Converting the Number of Stars as a Function of Mass to an IMF

For massive stars in star-forming regions of the Milky Way or Magellanic Clouds, we can count their actual numbers as a function of mass: this is the present day mass function (PDMF), and we can determine it directly as detailed in the next section. How these counts are related to the initial mass function (IMF) depends, however, on the history of star formation for the region in question. Here we will consider two extreme "stories" of star formation: (1) that all the stars in a region are born at the same time, and (2) that star formation has been continuous.

In the general case, let

$$\phi(m)\,\psi(t)\,dm\,dt$$

be the number of stars born per unit area between times t and $t+dt$ with mass between m and $m+dm$. Then $\phi(m)$ is the initial mass function (really just the probability of forming a star with a mass m), and $\psi(t)$ is the star formation rate. (In principle, $\phi(m)$ could also be a function of time, but I will assume that any dependence on t in ϕ is very slow over any time interval we will be considering.)

The number of stars seen at time t with masses between m_1 and m_2 will be

$$n_{m_1}^{m_2}(t) = \int_{m_1}^{m_2} \int_{t_0}^{t} \phi(m)\,\psi(t)\,dm\,dt, \tag{1}$$

if star-formation began at time $t = t_0$, and if the main-sequence life-time τ_{ms} of a star with mass m_2 is less than t.

Let us further make the usual simplifying assumption that the initial function is a power law, at least over the limited mass-range we are considering:

$$\phi(m) = Am^{\gamma}.$$

Star formation strictly coeval: In the case that all the stars in a region are born in a single event (at time $t = t_0 = 0$), the star-formation rate $\phi(m)$ is simply a delta-function, and at any time $t < \tau_{ms}(m_2)$ the integral in Equ. 1 simplifies to

$$n_{m_1}^{m_2} \propto m^{\gamma+1} \mid_{m_1}^{m_2} = m^{\Gamma} \mid_{m_1}^{m_2}. \tag{2}$$

It is common to refer to $\Gamma = \gamma + 1$ as the "slope" of the initial mass function for the following reason: if we take the log of both sides, we find

$$\log n = \Gamma \log m.$$

A Salpeter (1955) IMF has $\Gamma = -1.35$.

For the case of coeval star-formation, the present-day mass function *is* the initial mass function, except that stars whose mass-sequence life-times τ_{ms} are less than t will have disappeared. Nevertheless, the slope of the PDMF has the same slope as the IMF for stars whose mass is below this limit.

As long as we are careful to measure the age (or just ignore the highest mass mass-bin), then the slope we see today is a true reflection of the IMF. Another way of saying this is that if the mass bins are small enough, then evolution will have no effect on the measured slope of the IMF, but will simply selectively remove the highest mass stars.

Continuous star formation: For contrast, let us consider the seemingly unlikely scenario that star-formation has been continuous in the region in question. The star-formation rate $\phi(t)$ then is simply a constant, and the situation becomes steady-state: for a particular mass, the number of stars that are born each year is the same as the number of stars that leave the main-sequence each year. After a time t greater than the main-sequence life-time of the lowest mass we are considering [i.e., $t > \tau_{ms}(m_1)$], the number of stars we observe as a function of mass is just the product of the initial mass function with the main-sequence lifetime τ_{ms} at a particular mass m:

$$n_{m_1}^{m_2} \propto [m^{\Gamma} \times \tau_{ms}(m)]_{m_1}^{m_2} \tag{3}$$

What physical situations might we expect to correspond to these two cases? We will find in Sec. 3.2. and 3.3. that the massive stars in OB associations seem to have formed quite coevally—over less than a million years.

The second situation described above, that of continuous, steady-state star formation, does have physical application. If we average over a large enough number of star-forming regions of differing ages (such as the entire LMC, say) then our discussion is valid—*if* the global (average) star-formation rate has remained constant for at least $\tau_{ms}(m_1)$, where m_1 is the lowest mass we are considering. We will employ this approximation when describing the *field* massive star population of the Milky Way and Magellanic Clouds in Sec. 4.

Let me note finally that in practice when we measure the PDMF we count the number of stars on the HRD in various mass bins: between $85\mathcal{M}_\odot$ and $100\mathcal{M}_\odot$, say, and between $60\mathcal{M}_\odot$ and $85\mathcal{M}_\odot$, and so on. Following the notation of Scalo (1986), we introduce $\xi(\log m)$ to be $\phi(m)$ corrected for the size of the mass bin; i.e., the number of stars born per unit logarithmic (base ten) mass interval per unit area (kpc^2) per unit time (Myr). Taking the log of both sides of either Equ. 2 or 3 then leads to

$$\Gamma = d \log \xi(\log m)/d \log m.$$

In the case of an instantaneous burst, ξ is just proportional to n divided by the size of the jth mass bin $[\log(m_{j+1}/m_j)]$. In the case of continuous star formation, n also needs to be divided by the main-sequence lifetime τ_{ms} appropriate to the jth mass bin.

2.2. Complications for Massive stars: the Need for Spectroscopy

One of the great advantages we have in studying the massive star IMF is that we always have an instantaneous snapshot of current star-formation— there is no such thing as an "old" massive star. But, there are two distinct disadvantages, not always appreciated:

- Because of the extremely high effective temperatures of main-sequence massive stars, most of their luminosity is in the far-UV. Thus, the bolometric corrections are highly significant, and a very strong function of effective temperature. As a massive star evolves slightly to cooler temperatures, its absolute visual magnitude increases. Thus among high-mass stars one cannot distinguish bolometric luminosity (mass) on the basis of visual absolute magnitude. This means that luminosity functions have very little sensitivity to the IMF for massive stars.

- The extremely high temperatures mean that optical and near-UV colors only sample the extreme tail of the Rayleigh-Jeans distribution: there is virtually no difference in the colors of stars with $T_{\rm eff}$=50,000° and $T_{\rm eff}$=30,000°. Even going into the near-UV (λ1700, say) is insufficient to reliably distinguish early-type O stars from late-type O stars (Massey et al. 1996). Yet there is a factor of 1.5 difference in the masses of early-type and late-type O stars near the ZAMS. Relying upon photometry alone in determining the IMF for massive stars is doomed to failure.

To emphasize the first of these points, I have used the $z = 0.02$ (Galactic metallicity) evolutionary tracks of Schaller et al. (1992) to illustrate how stars evolve in Table 1. These tracks include the effects of mass-loss, and at least the lower metallicity tracks ($z = 0.008$ from Schaerer et al. 1993, and $z = 0.001$ from Schaller et al. 1992) have been shown to do an excellent job of reproducing the distribution of stars in the HRD of the LMC and SMC (Massey et al. 1995c). There is a significant discrepancy in the surface gravity deduced from these evolutionary tracks and from model atmospheres (see Vacca, Garmany, & Shull [1996] and Burkholder, Massey & Morrell [1997] for recent discussions of this problem); the empirical data from O-type binaries seems to favor the evolutionary tracks (Burkholder et al. 1997). The absolute visual magnitude has been computed by using the effective temperature to determine the bolometric correction. The spectral types have been assigned using the effective temperature scale of Vacca et al. (1996), which is primarily based upon the work of Puls et al. (1996) and references therein. I have assumed that stars with $\log g = 3.7$ to 3.9 are luminosity class "III", with higher surface gravity stars luminosity class "V" and lower surface gravities luminosity class "I" objects.

A 120 $\mathcal{M}_\odot$ star has a main-sequence lifetime $\tau_{ms} = 2.6$ Myr. The table shows that during its main-sequence evolution that the bolometric luminosity stays virtually unchanged at $M_{bol} = -11$, as would be expected from naive considerations. Yet, during this same time period the absolute visual luminosity *increases by nearly a factor of 10*, from $M_V = -6.2$ to $M_V = -8.6$. In a *mixed-age* population (such as found in a global study of a nearby galaxy), there would be a hopeless mixture of stars of different masses at a particular M_V: for instance, a 4 Myr old 40 $\mathcal{M}_\odot$ star will be *visually* brighter than a ZAMS (0 Myr) 120 $\mathcal{M}_\odot$ star.

Exactly how sensitive, then, would we expect a luminosity function to be to the IMF? We can easily construct luminosity functions for various IMFs using the evolutionary tracks. For a steady-steady population, we simply sample the tracks at equally spaced isochrones, weighing the contribution at each mass by a factor determined by the IMF. In such a population the same number of stars of a given mass will be found between any two equally spaced isochrones; i.e., there should be the same number of 40 $\mathcal{M}_\odot$ stars between 0 and 2 Myr as between 4 and 6 Myr. The luminosity functions are shown in Fig. 1. We

Table 1. Evolution of Massive Stars at Galactic Metallicity According to the $z = 0.02$ Tracks of Schaller et al. (1992)

120 $\mathcal{M}_\odot$
$\tau_{ms} = 2.56$ Myr

	Age (Myr)					
	0.0	0.5	1.0	1.5	2.0	2.5
Log T_{eff}	4.727	4.706	4.683	4.648	4.635	4.420
M_{bol}	−10.9	−10.9	−11.0	−11.0	−11.1	−11.1
log g [cgs]	4.1	4.0	3.9	3.7	3.6	2.6
M_V	−6.2	−6.4	−6.6	−6.9	−7.0	−8.6
Sp Type	O3 V	O3 V	O4 III	O5.5 III	O5 If	WNL

85 $\mathcal{M}_\odot$
$\tau_{ms} = 2.82$ Myr

	Age (Myr)					
	0.0	0.5	1.0	1.5	2.0	2.5
Log T_{eff}	4.705	4.691	4.673	4.645	4.577	4.449
M_{bol}	−10.3	−10.3	−10.4	−10.5	−10.6	−10.6
log g [cgs]	4.1	4.0	3.9	3.8	3.5	2.9
M_V	−5.7	−5.9	−6.1	−6.4	−6.9	−7.9
Sp Type	O3 V	O4 V	O4 III	O5.5 III	O7 If	B0 I

60 $\mathcal{M}_\odot$
$\tau_{ms} = 3.45$ Myr

	Age (Myr)						
	0.0	0.5	1.0	1.5	2.0	2.5	3.0
Log T_{eff}	4.683	4.671	4.660	4.645	4.620	4.576	4.464
M_{bol}	−9.6	−9.6	−9.7	−9.8	−9.9	−10.0	−10.1
log g [cgs]	4.2	4.1	4.0	3.9	3.8	3.6	3.1
M_V	−5.2	−5.4	−5.5	−5.7	−5.9	−6.3	−7.2
Sp Type	O4 V	O5 V	O5 V	O5.5 III	O6.5 III	O7.5 If	B0 I

40 $\mathcal{M}_\odot$
$\tau_{ms} = 4.30$ Myr

	Age (Myr)				
	0.0	1.0	2.0	3.0	4.0
Log T_{eff}	4.640	4.626	4.609	4.570	4.443
M_{bol}	−8.7	−8.8	−9.0	−9.1	−9.3
log g [cgs]	4.2	4.1	3.9	3.7	3.1
M_V	−4.6	−4.8	−5.1	−5.5	−6.6
Sp Type	O6 V	O6.5 V	O7 III	O8 III	B0.5 I

25 $\mathcal{M}_\odot$
$\tau_{ms} = 6.40$ Myr

	Age (Myr)						
	0.0	1.0	2.0	3.0	4.0	5.0	6.0
Log T_{eff}	4.579	4.570	4.563	4.552	4.535	4.504	4.438
M_{bol}	−7.5	−7.6	−7.7	−7.8	−7.9	−8.1	−8.3
log g [cgs]	4.2	4.1	4.1	4.0	3.8	3.7	3.3
M_V	−3.8	−4.0	−4.1	−4.3	−4.6	−4.9	−5.6
Sp Type	O8 V	O8 V	O9 V	O9 V	O9.5 III	O9.5 III	B0.5 I

20 $\mathcal{M}_\odot$
$\tau_{ms} = 8.14$ Myr

	Age (Myr)				
	0.0	2.0	4.0	6.0	8.0
Log T_{eff}	4.640	4.626	4.609	4.570	4.443
M_{bol}	−6.9	−7.0	−7.2	−7.4	−7.8
log g [cgs]	4.2	4.1	4.0	3.8	3.3
M_V	−3.5	−3.7	−4.0	−4.4	−5.3
Sp Type	O9.5 V	O9.5 V	B0 V	B0 III	B1 I

see that only a very modest change in slope of the luminosity function occurs over a wide range of IMF slopes, in accord with what we might expect from evolutionary considerations (Table 1). We have also included a version of the luminosity function if we "cut" the data to retain only the blue stars, similar to what real observers tend to do (see, for example, the determination of the IMF via luminosity functions by Freedman [1985]), but there is still little sensitivity to Γ in the luminosity function slope.

What if we observe a coeval population? The data in Table 1 does suggest that for a particular age, the bolometric luminosity is correlated with absolute visual luminosity, but there still is not very much sensitivity to the IMF slope, as we see in Fig. 2. The truncation of the luminosity function does tell us something about *age*, but the slopes of the luminosity still do not allow us to distinguish between IMF *slopes* in a meaningful way.

The second complication that we face in our study of the massive star IMF is the high effective temperatures of these stars and how little photometry can tell us about $T_{\rm eff}$ and hence the bolometric correction (BC). This point has been emphasized by Massey (1985), and it has been demonstrated that relying upon photometry alone leads to substantially wrong IMF slopes (Massey et al. 1989, 1995c). Why is this?

Let us consider the equations that allow the observer to go from observational quantities to an IMF slope, and see how we expect errors to propagate. Since massive stars evolve at nearly constant bolometric luminosity (Table 1), the inferred mass of a star is basically independent of the assigned effective temperature, were it not for how steep a function the BC is with $T_{\rm eff}$. Simply using the data in Table 1 (for instance, for 2 Myr) leads us to the mass-luminosity relationship for very massive stars:

$$L \propto m^{2.0}.$$

(This is less steep than that of solar-mass stars, where $L \propto m^{3.5}$.) Since $M_{bol} = -2.5 \log L + 4.75$, we find that

$$\Delta \log m = -0.2 \times \Delta M_{bol}. \tag{4}$$

Note, though, that a constant error in M_{bol} (due say, to an error in the adopted distance modulus) would lead to *no* error in Γ: stars might fall a little to the left or a little to the right of the proper isochrone, but the change in the number of stars with $\log m$ would remain the same. Instead, the sort of errors that will propagate into Γ are those which will affect our counts at one end of the range of masses but not at the other. This is just what we expect from the bolometric corrections: at the highest effective temperatures (and most luminous M_{bol}'s), the bolometric corrections are the most significant—and the photometry least sensitive to temperature.

We will adopt the relation between BC and $\log T_{\rm eff}$ found by Vacca et al. (1996): $\mathrm{BC} = 27.66 - 6.84 \times \log T_{\rm eff}$, and hence

$$\Delta \mathrm{BC} = -6.84 \times \Delta \log T_{\rm eff} \tag{5}$$

An error of 0.1 in our determination of $\log T_{\rm eff}$ leads to an error of -0.68 mag in the BC, which (according to Equ. 4) translates to an error of 0.14 in $\log m$.

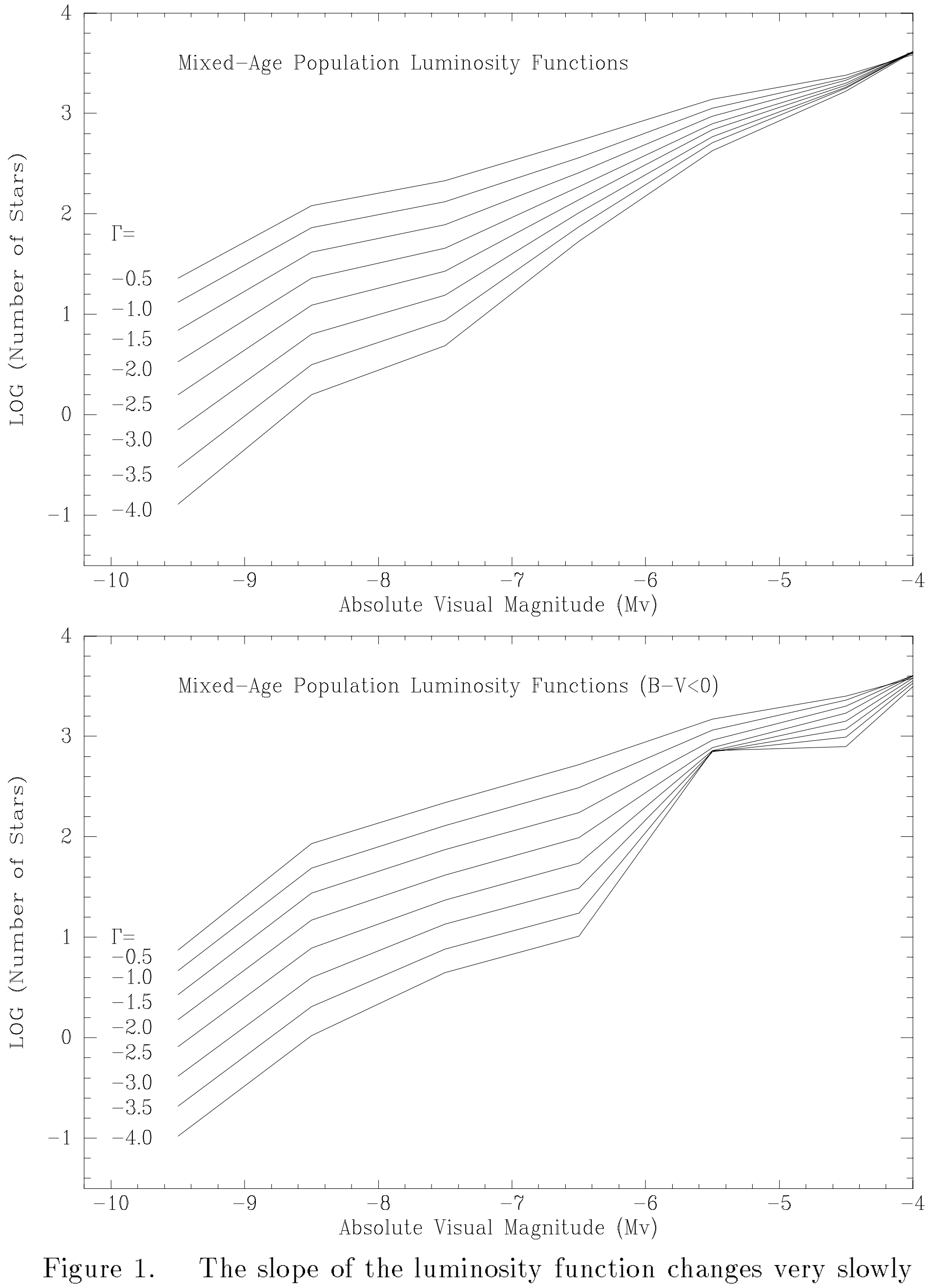

Figure 1. The slope of the luminosity function changes very slowly with changing IMF slope for a mixed-age population. The figure on the top shows the luminosity functions for stars of all colors; the one on the bottom shows what happens if we count only the blue stars ($B - V < 0$). In neither case is the slope of the luminosity function very sensitive to the IMF slope Γ. Each luminosity function has been normalized to a total of 10,000 stars with $M_V < -3$.

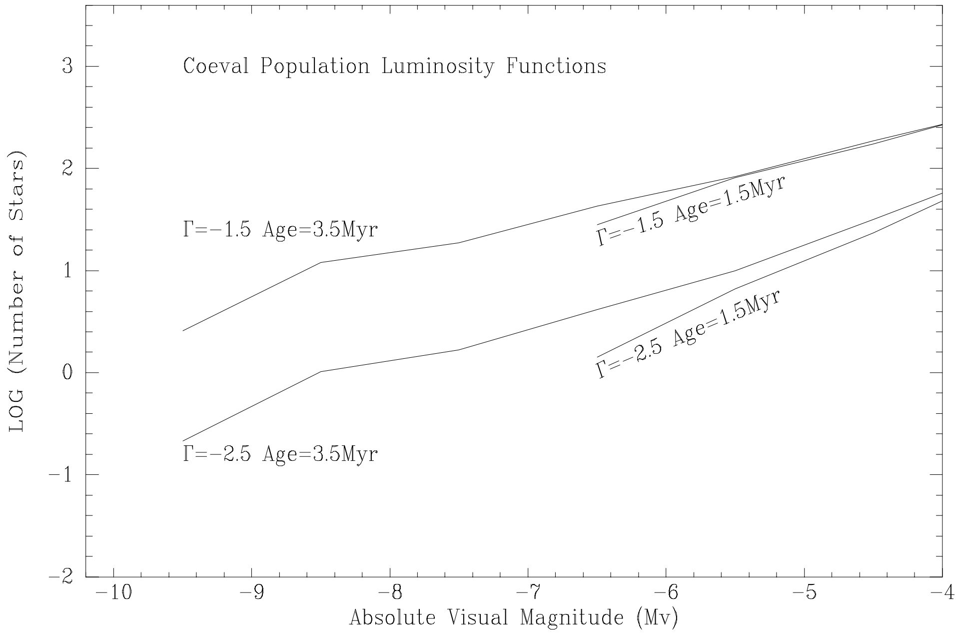

Figure 2. The luminosity functions for two different ages and IMF slopes.

If this happens only at one end of the mass range we are considering, we would expect it to have a 14% error on Γ; i.e., an error of 0.2 if $\Gamma \sim -1.3$. This is comparable to or greater than the usual statistical uncertainty we encounter in determining the IMF slope, as we will see in Sec. 3.1. How difficult is it to determine $\log T_{\rm eff}$ to 0.1 dex?

In Table 2 we show the colors determined from Kurucz model atmospheres as a function of effective temperature. In addition to the standard UBV colors, we also include the colors that would be measured on the *HST* WFPC2 system extending into the near-UV (1700Å): F170W, F336W, and F555W. (The latter two are nearly the same as U and V.) The *HST* compilation comes from the Hunter et al. (1997) study of the R136 cluster in the LMC and are based upon Kurucz (1992) atmospheres, while the UBV values are taken directly from Table 5 of Kurucz (1979). Surface gravities corresponding to main-sequence stars were used.

We can see at once from this table the validity of our claim that optical (and near-UV) colors are not very sensitive to $T_{\rm eff}$: A change of 0.11 mag in $(B-V)_o$ spans the *entire range* in effective temperatures tabulated here, a change of 0.4 in $\log T_{\rm eff}$, and hence a change in the BC of 2.7 mag, or 0.5 in $\log m$ (Equ. 4 and 5). As we might expect, we fare a bit better as we push to shorter wavelengths, with a change in $(U-B)_o$ of 0.38 and a change of 0.68 in $(F336W-F555W)_o$ for the same range in $T_{\rm eff}$.

In practice, though, we must deal with the complication of reddening. Even if we knew *a priori* the average reddening of an OB association, the intrinsic

Table 2. Intrinsic colors and reddening-free parameters as a function of effective temperature as determined from Kurucz (1979, 1992) atmospheres.

		z ∼solar			z ∼LMC		
$T_{\rm eff}$	$\log T_{\rm eff}$	(U-B)$_o$	(B-V)$_o$	Q	$(F170W-F555W)_o$	$(F336W-F555W)_o$	Q_{F170}
50K	4.699	−1.151	−0.326	−0.916	−2.351	−1.903	3.168
45K	4.653	−1.139	−0.320	−0.909	−2.308	−1.882	3.150
40K	4.602	−1.124	−0.311	−0.900	−2.229	−1.851	3.139
35K	4.544	−1.102	−0.298	−0.887	−2.128	−1.806	3.109
30K	4.477	−1.030	−0.286	−0.824	−2.088	−1.697	2.833
25K	4.398	−0.930	−0.255	−0.746	−1.902	−1.523	2.515
20K	4.301	−0.776	−0.217	−0.620	−1.671	−1.273	2.021

variations within an association are typically 0.2 to 0.3 in $E(B-V)$ (see Table 4 of Massey et al. 1995b). So, attempting to determine an IMF slope based upon a single color (such as $B-V$) is clearly hopeless. However, if we have two colors we can construct reddening-free indices: the well-known Johnson $Q = (U-B) - 0.72 \times (B-V)$ (Johnson & Morgan 1953) is tabulated in Table 2, along with the reddening-free UV parameter $Q_{F170} = (F170W - F555W) - 2.90 \times (F336W - F555W)$ introduced by Hunter et al. (1996).

In Fig. 3 we show how quickly the BCs change as a function of these reddening-free parameters. In practice, we can do no better than this in photometry. While Q_{F170} provides excellent separation in $\log T_{\rm eff}$ below 35,000°, both the Johnson Q and Q_{F170} provide very limited information at higher temperatures. Over the range 35,000° to 50,000°, we find that the Johnson Q parameter changes by only 0.03 mag per 1 mag change in the BC. For the UV-parameter Q_{F170}, we find a change of 0.05 mag per 1 mag change in the BC. These errors are comparable to the photometric accuracy we might hope to achieve. Note, however, that such large errors are only true for the hottest stars—for slightly cooler (lower-mass, lower-luminosity) stars there is good resolution—and hence it is easy to see how IMF slopes can become highly biased from extremely modest systematic problems in photometry, or by a slight bias in the transformation equations themselves.

Given that this leads to an unacceptable error for the IMF, what is the answer? Fortunately spectroscopy resolves this issue nicely. Classification spectra of the bluest stars (as judged from a reddening-free index) leads to accurate effective temperatures and bolometric corrections. Even at the highest effective temperatures a difference of one spectral subtype leads to differences in the bolometric correction of <0.2 mag; see Table 5-7 in Vacca et al. (1996).

We note that for $T_{\rm eff} = 35,000°$ and below the reddening-free indices *do* do a good job of determining the BCs. This temperature corresponds to a ZAMS mass of $20\mathcal{M}_\odot$ (at $z = .008 - .020$) or 15 $\mathcal{M}_\odot$ for $z = .001$. So, below 15-20 $\mathcal{M}_\odot$ we see that photometry alone (if two colors can be used to remove the effects of reddening) does a good job. However, around $T_{\rm eff} = 14,000°$ (3 $\mathcal{M}_\odot$ on the ZAMS), degeneracy sets in for Q.

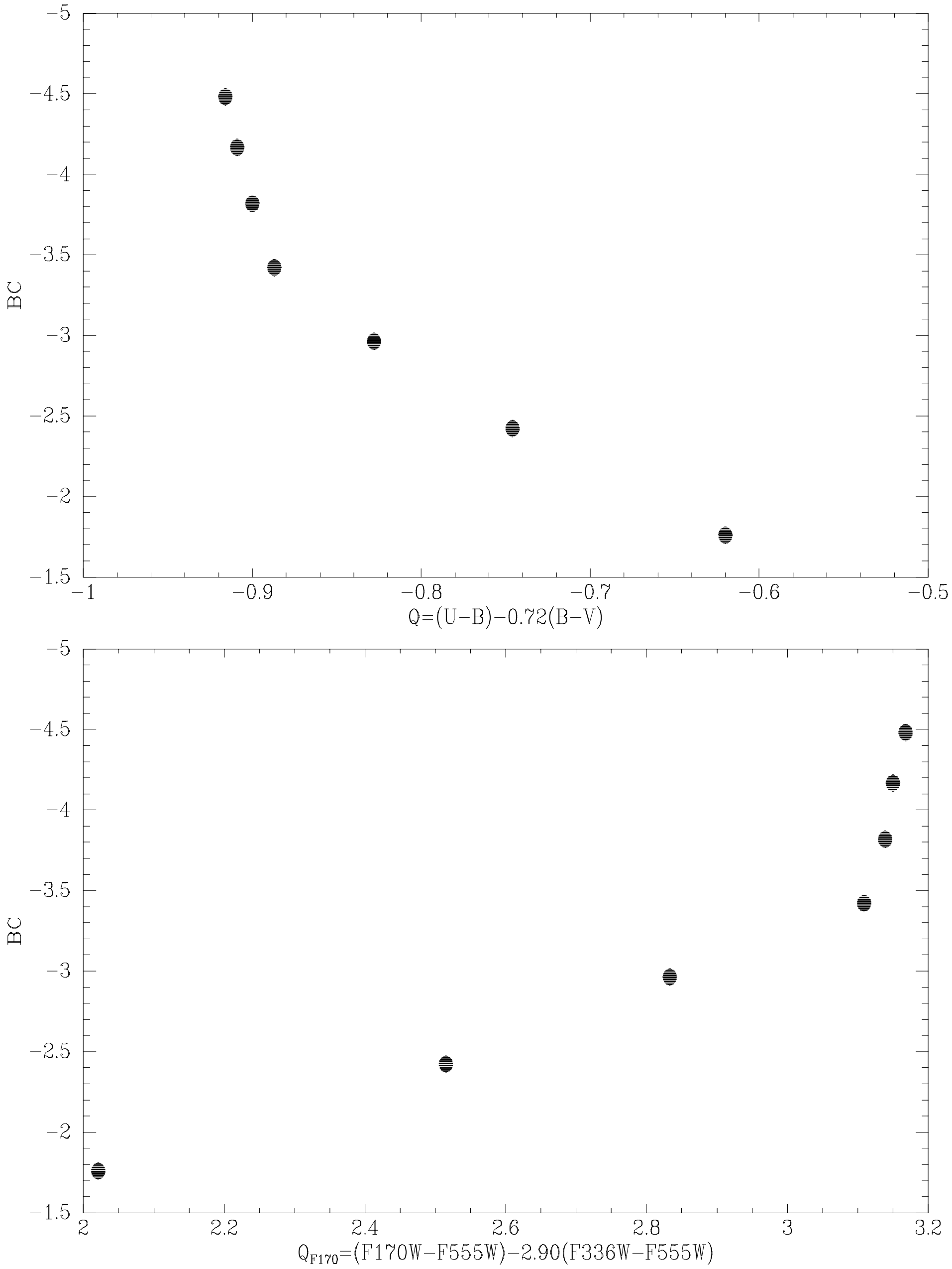

Figure 3. The reddening-free parameters Q and Q_{F170} provide only modest resolution of the bolometric correction at high effective temperatures. For the four hottest points, $\Delta BC/\Delta Q$ = 0.03 mag for the Johnson Q index, and $\Delta BC/\Delta Q_{F170}$=0.05 for the UV Q_{F170} index.

Let me close this section by citing an actual cautionary tale. In their study of the IMF in OB associations in the Magellanic Clouds, Massey et al. (1995c) combined spectroscopy of the bluest stars with CCD photometry to find an average IMF slope of $\Gamma = -1.3 \pm 0.3$, essentially Salpeter. Hill, Madore, & Freedman (1994a, b, c) had just completed their own study of OB associations in the MCs, based entirely upon UBV photometry, and derived a considerably steeper slope: $\Gamma = -2.0 \pm 0.5$. For the single region in common, LH 58, Massey et al. found $\Gamma = -1.4 \pm 0.2$, while Hill et al. found $\Gamma = -2.5 \pm 0.3$— a fairly significant difference! Massey et al. reanalyzed their data for LH 58 *ignoring the spectroscopy*, and adopting the transformation equations of Hill et al. Indeed, the IMF slope then steepened to $\Gamma = -2.0$. (The remaining difference with the Hill et al. result was presumably due to slight differences in the photometry.) This was in accord with the demonstration by Massey et al. (1989), who noted that the derived IMF slope for NGC 346 changed by 0.7 if the spectroscopy was ignored and only UBV photometry employed.

3. Results for OB Associations in the Milky Way, LMC, and SMC

3.1. The Initial Mass Function and Upper Mass "Cut-off"

During the last decade, my colleagues and I have been studying the massive star content of OB associations in the Milky Way (MW) and Magellanic Clouds (MCs) in order to answer the questions raised in the Introduction: what is the slope of the IMF for massive stars, and how does it differ as a function of physical conditions? The original motivation for this project was the intriguing finding by Garmany et al. (1982) of an apparent gradient in the IMF slope within the disk of the Milky Way. Using data taken from the literature, Garmany et al. concluded that the statistics on O-type stars were relatively complete to a distance of 3.0 kpc and found an IMF slope for this mixed-age population of $\Gamma = -1.6$. However, when they subdivided their sample into inwards and outwards of the solar circle, they found very different values for the IMF slope: $\Gamma = -1.3$ inwards of the solar circle, and $\Gamma = -2.1$ outwards of the solar circle.

One could imagine that in regions of higher star-formation (such as might be found in higher density regions that might occur inwards of the sun) one might preferentially make higher mass stars, given the expected relation between gas temperature and fragment mass proposed by Larson (1985, 1986). On the other hand, it had long been argued that higher mass stars would have increased difficulty forming in regions of higher metallicity; Shields & Tinsley (1976) had argued that the highest mass star formed should scale as $z^{-0.5}$. We realized that we could address this straightforwardly by a comprehensive program of UBV imaging and spectroscopy of SMC, LMC, and MW associations. The last sentence in the Garmany et al. paper counciled, "A beginning attack on these problems is now tractable for the brightest stars in nearby galaxies but will take a dedicated observational effort." Despite the effort required, the rewards are high: sampling the Magellanic Clouds provides a 10-fold progression in z from the SMC ($z = 0.002$) to the LMC ($z = 0.008$) to the solar neighborhood ($z = 0.02$).

The results of these studies have been summarized by Massey et al. (1995b, 1995c). We give the ages and derived IMF slopes in Table 3. Sally Oey has also completed a study of the stellar content of several LMC "superbubbles" (Oey & Massey 1995; Oey 1996); we have included only those regions from her studies containing bona fide OB associations with stars more massive than $25\mathcal{M}_{\odot}$.

In considering these numbers, we should understand that their principle strength is their accuracy in a differential, rather than absolute, sense. There was little we could do to account for the unseen binaries in these samples. As long as the binary frequency and mass distribution is similar, however, differential comparisons, between these different galaxies, should remain valid. Similarly, we cannot be sure that the Geneva group's evolutionary tracks give the right masses as a function of M_{bol}, although the data do tend to suggest that they do (Burkholder et al. 1997). Still, the differential comparisons should be valid.

There are two important conclusions that can be drawn from these studies:

- There is *no* difference in IMF slopes found between the Milky Way, the LMC, or SMC. The weighted average of the IMF slopes of the MW is -1.1 ± 0.1 (s.d.m.); that of the Magellanic Clouds is -1.3 ± 0.1. *Thus metallicity—at least over the factor of 10 spanned between the SMC and MW—does not affect the IMF of massive stars.*

- Similarly, there is *no* difference in the highest mass stars found in these associations: the youngest ones all contain massive stars $> 70 - 100\mathcal{M}_{\odot}$, regardless of metallicity. For the upper-mass limit dependence $m_{up} \propto z^{-0.5}$ suggested by Schilds & Tinsley (1976), we would expect a $\sim 3\times$ change in m_{up}. We can completely exclude this possibility: *radiation pressure on grains must not be the limiting factor in the mass of the highest mass star formed*—or at least on the highest mass star we see. Recent data on R136 (Sec. 3.3.) suggests that the limitation on m_{up} is statistical, rather than physical.

What then was the source of the Garmany et al. (1982) "gradient"? Massey et al. (1995b) proposed two explanations. First, our new data on Galactic OB associations demonstrated that prior knowledge of the massive star content of Galactic OB associations was incomplete. Humphreys & McElroy (1984) had argued that the Garmany et al. "slope" was due to studies of lower-mass (fainter) B stars being more selectively incomplete in the southern hemisphere. But the actual situation is more complex than this. For instance, in their study of the heavily reddened Cyg OB2 association, Massey & Thompson (1991) identified very many new O-type stars, some as early as O3! Yet the distance of Cyg OB2 is 1.7 kpc, well within the 3 kpc radius for which Garmany et al. had judged our knowledge of the O star content complete. In retrospect it is unsurprising that even relatively nearby associations harbored a significant number of unknown massive stars, due both to the complications of crowding and reddening. The second explanation offered by Massey et al. (1995b) to explain the gradient has to do with the inclusion of random "field stars" in the Garmany et al. sample. Evidence both from the Magellanic Clouds and Milky Way suggest that these massive stars in the field *do* have a very different IMF slope, $\Gamma \sim -3$ to -4. A relative change in the fraction of association and field stars with galactocentric distance would hence also explain the Garmany et al. gradient.

In an extremely intriguing paper, Elmegreen (1997) is able to produce the IMF Salpeter slope for all masses above some lower limit as a natural conse-

Table 3. The age, highest mass star (M_{up}), and the IMF slope Γ for OB associations in the Magellanic Clouds and Milky Way.

Association	Age (Myr)	m_{up} ($\mathcal{M}_\odot$)	Γ
		SMC	
NGC 346	2-4	70	-1.3 ± 0.1
		LMC	
LH 6[a]		85:	-1.7 ± 0.4
LH 9	1-5	55	-1.4 ± 0.2
LH 10	0-3	90	-1.1 ± 0.1
LH 38[a]		85:	-1.6 ± 0.2
LH 47/48[b]	2-3	50	-1.3 ± 0.2
LH 58	2-4	50	-1.4 ± 0.2
LH 73[a]		65:	-1.3 ± 0.4
LH 83[a]		50:	-1.3 ± 0.5
LH 114[a]		50:	-1.0 ± 0.1
LH 117/118	1-3	100	-1.6 ± 0.2
		Milky Way	
NGC 6823	2-7	40	-1.3 ± 0.4
NGC 6871	2-5	40	-0.9 ± 0.4
NGC 6913	4-6	40	-1.1 ± 0.6
Berkeley 86	2-3	40	-1.7 ± 0.4
NGC 7235	(6-11)	15	(-2.0)
NGC 7380	2	65	-1.7 ± 0.3
Cep OB5	2-4	30	-2.1 ± 0.6
IC 1805	1-3	100	-1.3 ± 0.2
NGC 1893	2-3	65	-1.6 ± 0.3
NGC 2244	1-3	70	-0.8 ± 0.3
NGC 6611	1-5	75	-0.7 ± 0.2
Cyg OB2	1-4	110	-0.9 ± 0.2
Tr 14/16	0-3	>120	-1.0 ± 0.2

[a]Oey (1996)

[b]Oey & Massey (1995)

quence of the hierarchical (fractal) structure of interstellar gas, similar to the proposal by Larson (1992). In Elmegreen's study, he finds that the derived mass function for stars is largely independent of the assumed physical parameters of the clouds, although slightly shallower IMFs may result in regions with lots of star formation activity; this goes in the same sense as predicted by Larson (1985, 1986).

3.2. The Star Formation History of OB Associations

The picture that emerges about the history of star formation in these OB associations is not as clear. The data are consistent with most of the massive stars having been formed in a short time: $\Delta\tau < 1$–2 Myr. (An occasional 5–10 Myr old red supergiant with mass $\sim 15\ \mathcal{M}_\odot$ is sometimes seen along with 2 Myr old O-type stars, suggesting that not *all* massive stars formed together; see discussion in Massey et al. 1989, 1995b.) However, when do the high mass stars form relative to those of intermediate mass?

In the traditional picture, intermediate mass stars should form first, and star formation should grind to a halt once the most massive stars form (Herbig 1961; see also Iben & Talbot 1966). Herbst & Miller's (1982) analysis of NGC 3293 established the paradigm suggested by Herbig's (1962) study of the Pleadies: that since the contraction times of the lower-mass stars seen on the main-sequence were much longer than the main-sequence lifetimes of the higher mass stars seen on the main sequence, the lower mass stars must have formed first. In the case of NGC 3293, evolutionary isochrones suggested an age of 6 Myr for the highest mass stars, while the contraction times of the lower-mass stars was 25 Myr.

Of the clusters listed in Table 3, NGC 6611 was the first, and most completely, studied in terms of the ages of the intermediate-mass stars (Hillenbrand et al. 1993). It challenges this paradigm directly, finding that "...the formation of O stars neither ushered in nor concluded the star-formation process in this young cluster." In NGC 6611 the highest mass stars are approximately 2 Myr old. But, there are pre main-sequence stars whose ages are spread from 0.25 Myr to > 1 Myr. Similar evidence was found by Massey et al. (1995b) for the majority of the Galactic associations listed above: although there are high mass stars with ages of approximately 2 Myr, much younger pre-main-sequence stars of intermediate mass exist. Clearly star-formation doesn't necessarily stop when the high mass stars form.

Why the difference between NGC 6611 and NGC 3293? It is clear that in the case of NGC 3293, the cluster is old enough for many of the highest mass stars to have exploded as supernova: after all, we can see from Table 1 that the highest-mass stars live on the order of 2–3 Myr. Unlike stars of lower mass, the main-sequence life-time is nearly independent of mass at the high end of the main-sequence (see Table 1). At 6 Myr, all stars with masses above, say, $30\ \mathcal{M}_\odot$ will have dissipated as supernova. But, in the case of NGC 6611, where the high-mass stars are perhaps only 2 Myr old, it is not clear that *any* supernova will have gone off.

It is natural, then, to ask if it is the *formation* of the high-mass stars—with their strong stellar winds—that disrupts the formation of lower-mass stars, or if it requires (in some environments) something more. Abbott (1982) finds that the highest mass stars ($\sim 100\ \mathcal{M}_\odot$) may contribute something like 2×10^{51} ergs via their stellar winds over their main-sequence lifetimes. He also notes that although the WR phase lasts for only 10% as long as the main-sequence lifetime, the energy contribution to the ISM is actually greater during the WR phase than during core H-burning for stars with masses 30–60 $\mathcal{M}_\odot$: about 5×10^{50} ergs. These numbers can be compared to the 10^{51} ergs imparted by each Type II supernova.

Thus, in clusters rich enough to have contained many very massive stars when they formed, we would expect that the energy input via stellar winds will play the most significant role in disturbing the gas. As we will see in Sec. 3.3, there was more than enough energy via stellar winds to halt star formation in the core of the R136 "supercluster" once a large number of extremely massive stars formed 1 Myr ago. However, in less rich clusters, one might not expect the gas to be disrupted until a massive star reaches the WR stage (2.5–3 Myr), and in others, not until the most massive stars become supernova (>3 Myr).

3.3. Recent Results on R136: A Cluster of O3 Stars

R136 is the object at the heart of the 30 Doradus nebula in the LMC: Once thought to be a single supermassive object, *HST* images revealed that it is something far more interesting: today R136 is considered the prototype of a "super star cluster", possibly the young analogs of globular clusters.

The massive star content of the R136 cluster had been studied from optical photometry with *HST* images (Hunter et al. 1995, 1996), which established that the IMF of stars of intermediate mass (2.8–15 $\mathcal{M}_\odot$) had an IMF slope that was only slightly shallower ($\Gamma\sim-1$) than Salpeter. However, even with the addition of ultraviolet photometry (Hunter et al. 1997), little could be said about the more massive and luminous stars in the R136 cluster for all the reasons discussed in Sec. 2.2.

However, spectroscopy of individual stars *is* possible using *HST*, even in the dense core of the R136 cluster. Deidre Hunter and I selected the best candidate stars from the UV photometry of Hunter et al. (1997), with the additional restriction that no other stars of comparable brightness be within the 0.13" radius of the FOS aperture. All together we obtained optical classification spectra of 65 of the bluest, most luminous stars (Massey & Hunter 1997).

What did we find? Star after star was of type O3, the hottest and shortest lived phase that only the most luminous and massive stars undergo. Altogether we confirmed 39 stars of this type—considerably more than were presently known in the rest of the LMC, the SMC, or the Milky Way combined.

Does the preponderance of very high mass stars indicate a top-heavy IMF? No! The derived IMF is in fact indistinguishable from a Salpeter slope, irrespective of whether the "hot" or "cool" temperature scale is used for the O3 stars: we find $\Gamma=-1.3\pm0.1$ to -1.4 ± 0.1, depending upon what temperature scale is adopted. Since our spectral types extended down only to 50 $\mathcal{M}_\odot$ or so, below which we were reliant on the photometry, we also considered this question from the other direction: given the number of stars of intermediate mass (2.8–15 $\mathcal{M}_\odot$) in successive annuli found by Hunter et al. (1996), how many very massive stars (50–120 $\mathcal{M}_\odot$) would we predict based upon extrapolating the intermediate-mass IMF? The numbers agreed well with what we found.

Thus we can conclude that the same IMF slope ($\Gamma \sim -1.3$) holds over the range 2.8 $\mathcal{M}_{\odot}$ to $> 100\mathcal{M}_{\odot}$ in R136, and over stellar densities ranging from that of a typical OB association to that of R136, a cluster comparable in stellar density to that of a globular cluster.

If the IMF is normal, then why are there so many stars of extremely high mass in R136? We believe that the answer is simply a combination of youth and richness: the presence of so many O3 stars establishes that the age of the core of the cluster is <1–2 Myr, since stars do not last long in the O3 phase. (This is a little longer than what one might infer from Table 1 as the ZAMS is shifted a little bit to higher temperatures for the lower metallicity of the LMC.) The richness of R136, combined with this youth, then results in there actually being such high mass stars—in a less rich cluster, one would run out of stars before such masses are reached at the top of the IMF. The mass of the highest mass star in R136 can only be estimated as $\sim 150\mathcal{M}_{\odot}$ by extrapolating the mass-luminosity relation, as evolutionary tracks are not available yet for these high luminosities.

This does suggest that the concept of upper-mass cutoff to the IMF may not even be valid. At what mass, consistent with age, does the IMF predict only one star? For a very rich, very young cluster like R136, the answer will be different than for the far sparser OB associations listed in Table 3, even for the ones young enough (and rich enough) to contain an O3 star or two (Tr 14/16, LH 10, LH 117, Cyg OB2, NGC 346). I would suggest that we have not yet found observational evidence for an actual cut-off to the IMF—instead, we simply run out of stars.

Finally, the story of star formation in the R136 cluster also casts some light on the issues raised in Sec. 3.2. Some of the intermediate-mass stars formed at least 4–5 Myr ago, as this is the time required to reach the main-sequence for the lowest mass main-sequence stars detected by Hunter et al. (1996). However, they also found stars that fell along the pre-main-sequence tracks, suggesting a range of younger ages, possibly as young as 1–2 Myr. With the age of the massive stars placed at ~1 Myr, a simple story emerges: within the central cluster, intermediate mass stars began forming 4–5 Myr ago, and star-formation ended there with the formation of the high mass stars 1 Myr ago. The many high mass stars, with their strong stellar winds, would quickly disperse the gas cloud from the immediate vicinity of the cluster, as has been observed.

4. Results for the Massive Stars in the Field

There are large, "monster" star-forming events that lead to "super star clusters" such as R136. There are smaller, more common star-forming events that are responsible for the many OB associations we see today. Could there be even more modest events that produce even fewer stars?

It has long been known that some very massive stars are not now currently found in OB associations. I first became aware of this myself when Peter Conti and I were identifying Wolf-Rayet stars in the nearby spiral galaxy M 33. While most WR stars were located where one would expect, some were far from any obvious OB association or spiral arms (Massey & Conti 1983). Madore (1978) had noted that not all star-formation took place in the arms, a fact that "subsequently led to some surprise each time it has been rediscovered."

In the case of WR stars, the oldest extant examples of massive stars, one might suppose that they had perhaps traveled far from their birth places, perhaps ejected as one of a pair of stars underwent a supernova, causing a classic "runaway". But, the statistics for O stars seemed surprising. For instance, in Fig. 2 of Garmany et al. (1982) it would appear that about half of the known O-type stars within 3 kpc of the Sun belonged to associations, and the other half did not. Could they all be run-aways?

Massey et al. (1995c) investigated this "field" population for the Magellanic Clouds, using what they argue is a highly conservative definition: to be counted as a "field star" an OB star would have to be at least 2' away from the boundary of any known OB association determined by Lucke & Hodge (1970) for the LMC or by Hodge (1985) for the SMC. This is the maximum that a star would travel in 10 Myr at a "typical" birth speed of 3 km s^{-1}. What they found was that the most massive stars in the field were as massive as any found in OB associations. Furthermore, when they restricted their sample to the "extreme field stars"—stars that were at least 15' from the boundaries of any OB association (100 km s^{-1} for a more representative 3 Myr time), the same was true: very massive stars are found in highly isolated regions.

There are even a number of O3 stars that exist in the field, which is hard to understand unless they were born there, given how short lived ($<$ 2 Myr) the O3 stage must be. We illustrate four of the LMC's O3 stars in Fig 4: two that are quite isolated, Sk $-67°$ 22 and Sk $-67°$ 274, and two that are members of OB associations, Sk $-67°$ 211 in the LH 82 and Sk $-71°$ 51 in LH 110. Underneath the nebulosity in the latter two are a dozen bright, blue stars. If the first two came from clusters, then where are the clusters? Runaway stars are believed to have been flung out of a cluster if its companion sheds enough mass in a supernova explosion—but this won't happen for $>$ 3 Myr, and these O3 stars are $<$ 2 Myr old. But, even if these somehow ran away at birth, they could only have covered 100 pc at a "runaway" velocity of 50 km s^{-1}. Yet no collection of young stars are visible within this distance. One is tempted to conclude that these O3 stars formed pretty much where we see them today, as part of very modest star-forming events, ones that perhaps produce "a single O star plus some change", as Jay Gallagher aptly put it.

If we accept that star formation comes in a variety of sizes, and then assume that these "field stars" are simply the result of small star-forming events, individually producing perhaps only a few stars each, what can we deduce about the initial mass functions of these small events? Massey et al. (1995c) found that the distribution of the field stars in the HRD was in excellent accord with the prediction of stellar evolutionary theory at each mass range if one assumed "steady-state" star-formation; i.e., that the number of 30-40 $\mathcal{M}_\odot$ stars between the 2 and 3 Myr isochrones were the same as between the 3 and 4 Myr isochrones, and that this continued until one reached the appropriate τ_{ms}. In addition to substantiating excellent agreement between the evolutionary models of the Geneva group (i.e., Schaerer et al. [1993] and Schaller et al. [1992]), this also suggested that one could use the simple assumptions of Equ. 3 to derive an IMF of these "field stars"—much as Garmany et al. (1982) had done for their

Figure 4. This figure illustrates four O3 stars in the LMC; two that are "field" stars according to Massey et al. (1995c), and two that are members of OB associations. Each image covers 15 arcmin (220 pc) on a side, and were constructed from STScI's Digitized Sky Survey material. Sk -71° 51 is buried within the nebulosity of LH 110.

mixed set of association and field stars. The results were surprising. Despite the presence of high mass stars in the field, the actual IMF slope was quite steep, with $\Gamma = -4.1 \pm 0.2$ (LMC) and $\Gamma = -3.7 \pm 0.5$ (SMC). They then reanalyzed the Garmany et al. (1982) sample, this time restricting the study to only stars outside of OB associations, and found a similar result, albeit with large uncertainties: $\Gamma = -3.4 \pm 1.3$.

Where do these "field stars" come from? In their study of Galactic OB associations, Massey et al. (1995b) found a range of a factor of 40 in the surface density of stars with masses $> 10\mathcal{M}_\odot$ (or a factor of 12 if Cyg OB2 is ignored): from 0.8×10^4 kpc^{-2} to 30×10^4 kpc^{-2}, with an "average" value of about 2×10^4 kpc^{-2}. The surface density of massive stars in the OB associations of the LMC and SMC is quite similar to this. In these units, the density of R136 is of order $> 630 \times 10^4$ kpc^{-2}, or more than two orders of magnitude greater than that of a typical OB association. If we accept, then, that there is a *continuum* of star-formation events—big to small—then how far down would we have to go to explain the number of field stars we see in the Magellanic Clouds? The answer depends upon what assumptions we make about the length of time over which star formation takes place in OB associations. If it is 1 Myr than the "star formation surface density" in the field is about 1.5 orders of magnitude less than that in OB associations; if it is 0.1 Myr, then the star formation surface density of field stars is down 2.5 orders of magnitudes.

Elmegreen (1997) offers a further suggestion to explain the formation of these field stars: that perhaps these form in non-fractal clouds, like the Taurus clouds. These are filaments with regularly spaced globules, and because of the regular (rather than hierarchical) structure, stars of a single mass will form preferentially, making the IMF particularly steep.

5. Extension into the Local Group: Indirect Evidence

What do we know of the massive star populations of galaxies beyond the Magellanic Clouds? In these proceedings, Rob Kennicutt is reviewing what we know primarily from integrated properties. In terms of resolved stellar content, we have mainly indirect evidence that bears on the massive star IMF.

There are of course some things that we *do* know directly. OB associations had been delineated in the Local Group galaxies M 31 (van den Bergh 1964), M 33 (Humphreys & Sandage 1980), NGC 6822 (Hodge 1977) and IC 1613 (Hodge 1978), demonstrating that these systems have substantial massive star populations, in accord with what had been inferred from Hα surveys. Photographic spectroscopy by Humphreys (1979, 1980a, 1980b) provided direct confirmation of some B and A supergiants in all four of these galaxies. The first unambiguous classification spectra of O-type stars beyond the Milky Way were obtained by Massey, Armandroff, & Conti (1986) followed by Humphreys, Massey, & Freedman (1990). Use of automatic plate measuring engines allowed extensive photographic catalogs of blue and red stars to be constructed for M31 (Berkhuijsen et al. 1988; Berkhuijsen & Humphreys 1989) and M33 (Ivanov, Freedman, & Madore 1993). Follow-up spectroscopy for some of the visually brightest of these stars were reported by Herrero et al. (1994). CCD surveys of selected regions in M31 (Massey, Armandroff, & Conti 1986; Odewahn 1987;

Hodge & Lee 1988; Hodge, Lee, & Mateo 1988), NGC 6822 (Wilson 1992), M33 (Wilson, Madore, & Freedman 1990; Wilson 1991; Regan & Wilson 1993) and IC 1613 (Freedman 1988) followed, but until recently the very limited spectroscopy available reduced all statements about the massive star populations to inference for the reasons demonstrated in Sec. 2.2.

In a recent study, my collaborators and I combined extensive optical spectroscopy with *UBV* photometry to give a direct measure of the relative massive star populations of M 31, M 33, and NGC 6822 (Massey et al. 1995a). We confirmed that stars with high luminosity ($M_{bol} \sim -11$) and mass ($> 80\mathcal{M}_\odot$) exist in M 31 and M 33, but are lacking in NGC 6822, where the most luminous star present has $M_{bol} = -10$ and a mass of 60 $\mathcal{M}_\odot$. The richest OB associations of M 31 are comparable to those known in the Milky Way and Magellanic Clouds, while those of M 33 appear to be even more extreme. Those in NGC 6822, on the other hand, are quite puny. We also confirmed that the Hα flux was not an incredibly reliable indicator of the massive star populations: as noted by Massey, Armandroff, & Conti (1986), two of the richest OB associations in M 31 show a similar surface density of massive, hot stars (OB 48 and NGC 206), but very different Hα fluxes, providing an important warning about the overinterpretation of integrated properties. A recent paper by Oey & Kennicutt (1997) shows that indeed as much as 50% of the ionizing radiation escapes from some of the Magellanic Cloud OB associations, while for others all of it is captured by the surrounding gas.

HST has now made it possible to distinguish multiple objects from single stars in these galaxies, at least with comparable resolution to what ground-based observations achieved in the Magellanic Clouds (0.3 pc). This ten-fold improvement in resolution for the more distant members of the Local Group allows an important check on how well we really have been doing in these galaxies. Fig 5 shows three "state-of-the-art" U-band images of a section of an OB association in M 33. The top image was taken with the Canada-France-Hawii Telescope in sub-arcsecond conditions by Wendy Freedman and Barry Madore using a photographic plate. The middle image was taken with an RCA CCD on the Kitt Peak 2.1-m, also in sub-arcsecond conditions, and is typical of the data used in the photometry by Massey et al. (1995a). The bottom panel shows a WFPC2 image of the same field. One of the comforting things that this comparison shows is that stars that we concluded were probably single from good, ground-based imaging generally really *are* single, at least as judged from *HST*. This is consistent with a point that Rolf Kudritzki has emphasized (cf. Kudritzki et al. 1996): there will generally be at least spectral indications if a star is composite. Star OB88-6 (on the right) doesn't look quite stellar in either of the ground based images, and it is nicely separated by the WFPC2 image; similarly a star like OB89-4 (at lower left) is a real mess in all three images. The star OB88-10, though, which appears stellar on both the photographic plate and the ground-based CCD image is still nicely single even on the WFPC2 image. (It is a B0.5 Ia; see Table 4 of Massey et al. 1995a). OB89-7 is a case where we might have fooled ourselves: a close examination of the WFPC2 frame reveals it is a very close double; no spectroscopy is available for this star.

However, while such studies can tell us something about what is happening at the very top of the HRD, we are a long ways from being able to deduce

Figure 5. Three U-band images of the OB88/89 association in M 33. All three were "state of the art" in their day, and show the improvement that has taken place over the past 15 years. The CFHT plate (top) was taken in 1981 by Freedman and Madore; the middle image was obtained on the KPNO 2.1-m with an RCA CCD during 1986. Both ground-based images were taken in sub-arcsecond conditions. The bottom image was obtained with the WFPC2 camera on *HST* during 1996.

anything about the IMF. A $25\mathcal{M}_\odot$ ZAMS star (O8 V) will have $M_V \sim -3.8$ despite having an $\mathrm{M}_{bol} \sim -7.5$ (Table 1), and hence be found at $V = 21.3$ in M 33, assuming an average reddening (Table 5 of Massey et al. 1995a). Even a $60\mathcal{M}_\odot$ ZAMS star will be only $V \sim 19.9$. Thus sufficient completeness to construct an IMF is simply not possible, given even unlimited access to 4m class telescopes.

However, there are some intriguing hints about the massive star populations based upon indirect indicators. Massey (1985) noted that although it was difficult to obtain a complete sample of Wolf-Rayet stars in a galaxy even at 1 Mpc, it was at least a lot more practical than studying the unevolved massive star population directly. Complete samples have been identified in selected OB associations of M 31, M 33, NGC 6822, IC 1613 (Armandroff & Massey 1991 and references therein), as well as IC 10 (Massey & Armandroff 1995).

According to stellar evolutionary theory—and this appears to be borne out observationally (Massey 1996)—the mass limit for becoming a Wolf-Rayet star depends on the metallicity. This was inherent in the first suggestion that normal O stars may evolve into a Wolf-Rayet stage (Conti 1976) via mass-loss, and expanded upon by Maeder, Lequeux, & Azzopardi (1980). In a lower metallicity environment, the mass-loss rates should be lower, and hence it should be more difficult to produce a Wolf-Rayet star—only the highest mass stars should make it. In a higher metallicity environment, mass-loss rates will be greater at any given luminosity, and hence a somewhat lower mass (but still massive) star will evolve to the WR stage. Thus we have to be careful in using Wolf-Rayet stars as tracers of the most massive stars, but it can be done.

As a massive star evolves, it should first become a WN-type Wolf-Rayet star, with the products of CNO burning visible at the stellar surface. More evolution and mass-loss will lead to a WC-type Wolf-Rayet star, with the products of triple-α He-burning visible at the surface. Thus with higher mass-loss rates (such as are expected to occur in galaxies with higher metallicities) WCs will be produced at a lower mass than would be possible at lower metallicities. Thus the relative number of WC and WN Wolf-Rayet stars should depend upon metallicity. Historically, this was well known to be true for the Milky Way, the LMC, and the SMC ($z = 0.02$, $z = 0.008$, and $z = 0.002$, respectively, according to Lequeux et al. 1979), where the WC:WN ratio was something like 1:1, 1:3.5, and 1:7 (see Massey & Conti 1983 and references therein). Extension to other galaxies of the Local Group have been difficult, and extended over the past decade, but in Fig. 6 we show the results of this work.

Indeed, we find that there really is an excellent correlation between the observed WC/WN number ratio and the metallicity, with one very notable exception. Massey (1996) argues that the solar neighborhood point might be a little higher than is actually the case, and that the NGC 6822 and SMC points can flop around a little due to small-number statistics and an uncertain metallicity for the former. But, in any event the galaxy IC 10 simply has way too high a proportion of WC stars for its metallicity.

IC 10 is a starburst galaxy, as determined by a comparison of its star formation rate determined from its Hα luminosity with the fuel for continued star formation, its H I mass (Massey & Armandroff 1995). Consistent with this is

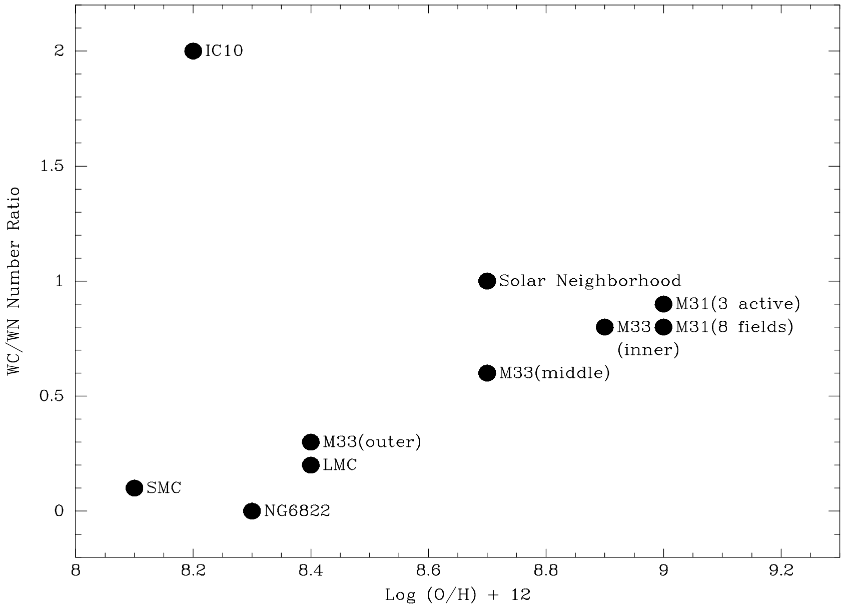

Figure 6. The WC/WN ratio correlates well with metallicity, with the glaring exception of IC 10, a starburst galaxy. This figure is from Massey (1996).

the incredibly high surface density of Wolf-Rayet stars seen: the *global* average is similar to that seen in the richest, most intense OB associations of M 33.

Why would a starburst cause a high fraction of WC stars? One could obtain this if all of the massive stars in IC 10 formed in lock-step; i.e., that the burst occurred with a *very* short time scale ($\Delta\tau < 0.3$ Myr) *and* that we were viewing the galaxy at a very special time. However, while such short time-scales might very well be applicable to the individual *knots* of star-formation seen in "Wolf-Rayet galaxies" Conti & Vacca 1994), it does not seem applicable here: how could star formation take place simultaneously over a scale of several kpc? (See discussion following Massey 1996.) Instead, one is driven to consider the possibility that there are simply a much higher proportion of the highest mass stars present in IC 10: that at least in this particular starburst galaxy the massive star IMF may be much flatter. This does not appear to be the case for "Wolf-Rayet galaxies" in general, as Schaerer (1996) is able to produce the observed WR/O star ratios with normal IMFs and sensibly long star burst times.

6. Conclusions

We have answered the questions posed in the introductions as follows:

1. For massive stars, we can actually determine the initial mass function (IMF) *directly*: the present-day mass function (PDMF) can be determined by counting individual stars as a function of mass; the PDMF can be

converted to an IMF by using the H-R diagram to determine the star-formation history of the region. There are two significant complications to keep in mind: (a) as a massive star evolves at nearly constant bolometric luminosity, its visual luminosity increases as its effective temperature decreases. The visually brightest stars in a mixed-age population (such as what we see when we look at a nearby galaxy) will be A-type supergiants of modest mass ($15\mathcal{M}_\odot$), not the considerably higher mass, more bolometrically luminous stars that may be present. (b) Because of the degeneracy of optical and near-UV colors for high effective temperatures (T_{eff} >30,000°K), spectra are needed to determine a star's effective temperature accurately enough for the bolometric correction to be sufficiently well determined for a meaningful placement in the HRD. Because of both of these effects, luminosity functions are not sensitive to the IMF in mixed-age populations.

2. The IMF appears to be well represented by a Salpeter slope ($\Gamma \sim -1.3$) regardless of metallicity ($z = 0.002$ to $z = 0.02$), and over the range of stellar densities we find from that of sparse OB associations to that of the "super star cluster" R136, where the stellar density is similar to that of a globular cluster. At very low star formation levels found in the field, however, the IMF of massive stars appears to be very steep ($\Gamma \sim -4$).

3. The concept of "upper mass cutoff" may not even be valid. The youngest, richest region studied directly, that of R136, has stars that appear to be as massive as 150 $\mathcal{M}_\odot$—but even there the limit may be *statistical* rather than *physical*. We may simply have not yet encountered a region that is rich enough that it should have produced stars more massive than what nature can actually produce.

4. Star formation in OB associations appears to begin with stars of intermediate mass and lasts for several million years. However, the highest mass stars appear to form over a short time scale ($\Delta\tau$ <1–2 Myr). In some cases star formation of intermediate stars continue (NGC 6611); however, if sufficient number of very high mass stars are formed (R136), the gas is swept away and star formation in the immediate vicinity ends with the formation of the most massive stars.

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