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MASSIVE STARS IN THE FIELD AND ASSOCIATIONS OF THE MAGELLANIC CLOUDS: THE UPPER MASS LIMIT, THE INITIAL MASS FUNCTION, AND A CRITICAL TEST OF MAIN-SEQUENCE STELLAR EVOLUTIONARY THEORY

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ABSTRACT

We investigate the massive star population of the Magellanic Clouds with an emphasis on the field population, which we define as stars located further from any OB association than massive stars are likely to travel during their short lifetimes. The field stars must have been born as part of more modest star-forming events than those that have populated the large OB associations found throughout the Clouds. We use new and existing data to answer the following questions: Does the field produce stars as massive as those found in associations? Is the initial mass function (IMF) of these field massive stars the same as those of large OB complexes? How well do the Geneva low-metallicity evolutionary models reproduce what is seen in the field population, with its mixed ages?

To address these issues we begin by updating existing catalogs of LMC and SMC members with our own new spectral types and derive H-R diagrams (HRDs) of 1584 LMC and 512 SMC stars. We use new photometry and spectroscopy of selected regions in order to determine the incompleteness corrections of the catalogs as a function of mass and find that we can reliably correct the number of stars in our HRDs down to 25 $M_\odot$. Using these data, we derive distance moduli for the Clouds via spectroscopic parallax, finding values of 18.4 ± 0.1 and 19.1 ± 0.3 for the LMC and SMC. The average reddening of the field stars is small: $E(B-V) = 0.13$ (LMC) and 0.09 (SMC), with little spread.

We find that the field *does* produce stars as massive as any found in associations, with stars as massive as 85 $M_\odot$ present in the HRD even when safeguards against the inclusion of runaway stars are included. However, such massive stars are much less likely to be produced in the field (relative to lower mass stars) than in large OB complexes: the slope of the IMF of the field stars is very steep, $\Gamma = -4.1 \pm 0.2$ (LMC) and $\Gamma = -3.7 \pm 0.5$ (SMC). These may be compared with $\Gamma = -1.3 \pm 0.3$, which we rederive for the Magellanic Cloud associations. (We compare our association IMFs with the somewhat different results recently derived by Hill et al. and demonstrate that the latter suffer from systematic effects due to the lack of spectroscopy.) Our reanalysis of the Garmany et al. data reveals that the Galactic field population has a similarly steep slope, with $\Gamma = -3.4 \pm 1.3$, compared to $\Gamma = -1.5 \pm 0.2$ for the entire Galactic sample. We do not see any difference in the IMFs of associations in the Milky Way, LMC, and SMC.

We find that the low metallicity evolutionary tracks and isochrones do an excellent job of reproducing the distribution of stars in the HRD at higher masses, and in particular match the width of the main-sequence well. There may or may not be an absence of massive stars with ages less than 2 Myr in the Magellanic Clouds, as others have found for Galactic stars; our reddening data renders unlikely the suggestion that such an absence (if real) would be due to the length of time it takes for a massive star to emerge. There is an increasing discrepancy between the theoretical ZAMS and the blue edge of the main-sequence at lower luminosities; this may reflect a metallicity dependence for the intrinsic colors of stars of early B and later beyond that predicted by model atmospheres, or it may be that the low metallicity ZAMS is misplaced to higher temperatures. Finally, we use the relative number of *field* main-sequence and Wolf-Rayet stars to provide a selection-free determination of what mass progenitors become WR stars in the Magellanic Clouds. Our data suggest that stars with initial masses >30 $M_\odot$ evolve to a WR phase in the LMC; while the statistics are considerably less certain for the SMC, they are consistent with this limit being modestly higher there, possibly 50 $M_\odot$, in qualitative agreement with modern evolutionary calculations.

Subject headings: Magellanic Clouds — open clusters and associations: general — stars: early-type — stars: evolution — stars: luminosity function, mass function

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1. INTRODUCTION

1.1. *Background and Goals*

The Magellanic Clouds are rich in massive stars, and their low reddening and proximity make them ideally suited to studies of star formation. Our previous work has probed the stellar content of several OB associations in both the LMC and the SMC, and we have begun the critical comparison with the content of OB associations of the Milky Way (see summary in Massey 1993). This latter comparison allows us to understand the role that metallicity plays in the star-formation process, given the 10-fold progression in z between these three systems ($z = 0.02$ in the solar neighborhood, $z = 0.008$ for the LMC, and $z = 0.002$ for the SMC, according to Lequeux et al. 1979). To date we have found only a modest variation in the slope of the initial mass function (IMF), with the differences found within the LMC as large as those found between the Milky Way and the Clouds. We have also found that the mass of the highest mass stars is similar in all three galaxies, contrary to the predictions of star formation models in which radiation pressure on grains is expected to be the limiting factor in how massive a star may form. Stars of masses $\approx 85\ M_{\odot}$ are found in the OB associations of the Milky Way, LMC, and SMC. Shields & Tinsley (1976) suggest that the upper mass limit should scale as $1/z^{1/2}$; we should then expect to see a factor of 3 change in the mass of highest mass stars found in the Milky Way and the SMC. (See also Wolfire & Cassinelli 1987.)

In the present paper, we turn our attention from the densely populated OB associations to the isolated field stars of the Magellanic Clouds. Although we are used to thinking of star formation (and particularly the formation of massive stars) as taking place in giant "stellar nurseries" (factories?), a significant fraction of massive stars are found among the *field* population in our own Galaxy. Garmany, Conti, & Chiosi (1982) compiled a volume-limited sample of O stars within 3 kpc of the Sun in which $\approx 50\%$ appear to be well away from any OB association (see their Fig. 2). We will argue later that this is an overestimate (with OB association stars underrepresented), but the evidence still suggests that a significant fraction of massive stars are born outside clusters and associations in the Galaxy. This is consistent with our (scant) knowledge of the massive star population of the nearby spiral galaxy M33, when small H II regions and some Wolf-Rayet and O stars are known to exist in relative isolation, far from any large OB complexes (Madore 1978; Massey & Conti 1983; Oey & Massey 1994). Similarly, some Hubble-Sandage variables in M31 are isolated (Gallagher, Kenyon, & Hege 1981; Kenyon & Gallagher 1985). Given that the lifetimes of these stars are quite short ($< 10^7$ yr), individual stars might be expected to have drifted by at most a few tens of parsecs from their birthplace, and hence must have been born during modest star-forming events.

The goals of the present study are to answer the following three questions:

1. Are equally massive stars born in the field as in OB associations? Short-lived, massive stars currently found in the field must not have been formed in the large, prodigious star-forming events that we have come to associate with massive stars. Larson (1982, 1986) has suggested that low- and high-mass stars may form under different conditions, and one might imagine that the maximum stellar mass may depend upon the mass of the parent molecular cloud. Hunter & Massey (1990) investigated this issue in the Galaxy by characterizing the stellar content of Galactic H II regions chosen for their small size—typically an order of magnitude smaller than giant regions. They found that the highest mass stars present in small H II regions are in fact similar in mass to those found in large complexes, and hence that the production of high-mass stars does not require extraordinary parent clouds.

2. Is the mass function for stars born in the field the same as for stars born in richer environments? We have found similar IMF slopes for the OB associations of the Galaxy, LMC, and SMC, suggesting that metallicity is not the key player in the local processes that form stars. If the IMF for stars in the field is similar to that of the OB associations, then we might wonder if there are *any* physical properties that affect star formation.

3. How good are current evolutionary models and isochrones? Because the field population will contain stars of all ages, we can make a critical test of stellar models by determining if the distribution of stars in the H-R diagram (HRD) matches that expected from the isochrones within a given mass ranges. Meynet, Mermilliod, & Maeder (1993) have made comparisons of the higher metallicity ($z = 0.02$) models to Galactic clusters, but by using the field stars of the LMC and SMC we can provide a test of the width of the main sequence, which studies of a coeval population, such as that of a cluster, cannot. This test is of particular interest as early analysis of H-R diagrams in the Clouds found that stars occupy a considerably broader region of the H-R diagram than expected by the then-current evolutionary models: the so-called "main-sequence widening" problem (see discussion in Meylan & Maeder 1982). Our study comes at a time when new evolutionary models computed for metallicities appropriate to the LMC and the SMC have recently become available (Schaller et al. 1992; Schaerer et al. 1993).

An additional question we would like to be able to answer, but cannot fully address as yet, is how metallicity affects post-main-sequence massive star evolution. We expect that metallicity might play a strong role given its expected influence on mass loss. Thus one might expect to see a substantial trend in the ratio of O to WR stars as a function of metallicity (Maeder, Lequeux, & Azzopardi 1980). It has been known since the work of Azzopardi & Breysacher (1979) that the number of WR stars in the LMC and the SMC do not scale with galaxian mass, and this has sometimes been taken as evidence of the possible effect metallicity plays on the evolutionary history of massive stars: at lower metallicities (such as in the SMC), mass-loss rates are expected to be lower and hence a massive star would have a harder time peeling off its outer layers to become a Wof-Rayet star. (See discussion in Massey 1985.) However, differing numbers of potential progenitors (due to either differing star formation rates or IMF slopes) could equally well explain the discrepancy, as well as the more subtle effect of differing binary frequencies or distribution of rotational velocities. The population of Wolf-Rayet stars is relatively well known in both the LMC and the SMC; what we do not know is the relative unevolved (O and early B) star populations. The present work addresses this for the field stars, but additional work on the OB associations stars are needed in both Clouds to address this question adequately.

The analysis involved in this study will also lead us to a new determination of the distance to the LMC and SMC, and provide direct information on the average reddening of their early-type stars.

1.2. *Methods and Problems*

In order to answer the three questions posed above we must derive H-R diagrams for the LMC and SMC in which the placement of stars is sufficiently accurate to be able to judge masses and ages, and which are relatively free of contamination by OB association stars. We will begin with the extensive published catalogs of stars with photometry in the two Clouds, and supplement published spectral types with our own slit spectroscopy. We will then use the location and size of the cataloged OB associations in the Clouds to restrict the sample to pure field stars.

In undertaking this analysis there are two observational problems that should be stressed. The first of these is that the visually brightest stars in the Clouds are not the most luminous (and massive), owing to the strong temperature dependence of the bolometric correction. Thus in any (V) magnitude-limited sample the incompleteness for a given luminosity (or mass) will depend strongly on the color (i.e., effective temperature). This effect was well recognized by Rousseau et al. (1978) in their extensive compilation of 1822 LMC members; they note that their work is not complete for the hottest stars and that this means that "the greatest care must be taken in interpreting the H-R diagram which can be derived from the present data." Unfortunately, this proviso has been overlooked by those who have proceeded to derive IMFs from these data.

We illustrate this point in Figure 1, where we show the $z = 0.001$ (SMC-like) evolutionary tracks from Schaller et al. (1992). As stars evolve away from the zero-age main-sequence (ZAMS) (inferred from the beginning of the tracks on the left), they do so at nearly constant bolometric luminosity. However, because the bolometric correction is a steep function of effective temperature, particularly for the hotter stars, a star evolving to the right will become visually brighter as more of its energy shifts from the ultraviolet to the visible. We can demonstrate this by sketching in the locus that would be occupied by stars between $V = 11.0$ and $V = 14.0$ magnitudes stars of different effective temperatures. Thus if a catalog was complete to $V = 14$ mag, all of the B8 supergiants (log $T_{\rm eff} \approx 4.1$) would have been detected as faint as $M_{\rm bol} \approx -6$ or a mass of 10 $M_\odot$. However, near the ZAMS completeness to $V = 14.0$ would imply completeness in luminosity only to $M_{\rm bol} = -10.5$ or a mass of 60 $M_\odot$! In Figure 1*b* we have populated this H-R diagram with stars drawn from the Azzopardi & Vigneau (1982) catalog of the SMC stars for the purposes of illustration; the actual derivation of this diagram will occupy most of § 2. (We note in passing that this diagram reveals that the photometry is likely complete to $V = 14$ but not much beyond that.) We can see clearly from this diagram that the visually brightest stars in the SMC are the 25–40 $M_\odot$ A-type supergiants (log $T_{\rm eff} \approx 4.0$), which are 10th to 11th magnitude. However, the most massive stars present are considerably fainter: the 60 $M_\odot$ stars near the ZAMS are closer to $V = 14$, with a handful of slightly more evolved 60–85 $M_\odot$ stars perhaps a magnitude brighter.

We will deal with this first problem by using information from several test fields in which we push deeper in order to understand the correction for incompleteness. We will find that in practice we can reliably correct the LMC and SMC catalogs down to 25 $M_\odot$.

The second problem that we encounter is the need for spectroscopy to place the hotter stars (log $T_{\rm eff} > 4.4$) in the H-R diagram. For stars this hot, the broad-band colors are degenerate. For example, there is a difference of only 0.02 mag in $(B-V)_0$ between a B0 V and an O5 V star (FitzGerald 1970); earlier than O5 there are no comparable data, but $(B-V)_0$ should change by perhaps an additional 0.01 mag in extrapolating to O3 V. Yet in this same range log $T_{\rm eff}$ increases from 4.46 to 4.69 (Chlebowski & Garmany 1991). Even using $(U-B)_0$ colors, we find a change of less than 0.10 mag over the range O9–O3 (log $T_{\rm eff} = 4.55$–4.69). Yet differing color calibrations differ by a significant amount; i.e., there are 0.02 mag differences in the color to spectral type calibration of O stars given by Schmidt-Kaler (1982) and FitzGerald (1970). The crucial effect of this degeneracy is not so much in placing the stars in log $T_{\rm eff}$, however, but rather in the resulting uncertainty of the bolometric magnitude and hence the conversion to luminosity and mass. For instance, in the same 0.03 mag range in $(B-V)_0$ from B0 to O3, the bolometric correction changes from -3.0 to -4.4. For two stars of equal M_V this difference in the bolometric correction then corresponds to a factor of nearly 3 in mass (i.e., 30–85 $M_\odot$ for $M_V \approx -6.0$). In practice the situation is confused by reddening, and by the fact that the mapping from $(U-B)_0$ or $(B-V)_0$ (or a reddening-free index) to log $T_{\rm eff}$ is actually dependent upon log g as well: supergiants and main-sequence stars of the same color do not have the same effective temperatures. As an example of this we note that both an O9 I star and an O7 V star have identical values of the reddening-index $Q = (U-B) - 0.72(B-V) = -0.94$: Does a star with this value of Q have log $T_{\rm eff} = 4.51$ and a BC $= -3.1$, or does it have log $T_{\rm eff} = 4.60$ and a BC $= -3.8$? Since the "standard" M_V's of stars classified as O7 V and O9 I overlap (see Table 1-21 in Conti 1988), we cannot tell its placement within the HRD to better than 1.5 mag given photometry alone, and hence whether its mass is 25 $M_\odot$ or 60 $M_\odot$. *Spectroscopy, however, resolves the effective temperature neatly, and moderate (2–3 Å) resolution allows us to readily distinguish between such stars.*

We illustrate in Figure 1*c* this pressing need for spectroscopy and the resulting errors that result from (1) an error of one spectral subtype, and (2) an error of one spectral luminosity class. The error bars are slanted as an error in effective temperature translates to an error in luminosity via the bolometric correction. Note, however, that an error in one luminosity class is nearly parallel to the error produced in being off by one spectral subtype. For comparison, we show in Figure 1*d* the resulting errors in the HRD due to finite errors in photometry, where the error bars correspond to an uncertainty of 0.02 mag in $(B-V)_0$ and $(U-B)_0$. We emphasize, though, that this error can mostly mainifest itself as a systematic problem via the calibration equations, and we will demonstrate in § 3.3.3 that reliance upon photometry alone can bias derived IMFs. For our work, we obtained classification spectroscopy for the stars whose photometry indicates a reddening-free index $Q < -0.7$, roughly corresponding to spectral type B1.5 and earlier. (At later types photometric errors are not as important, and run essentially parallel to the mass tracks by A0.)

A third problem—but about which we can do little—is the fact that some number of the stars in the HRD will be binaries. If the magnitude difference is large between the two components, then our photometry detects and spectroscopy places the brighter (in V) component but ignores the other. If the magnitude difference is small, then our photometry and spectroscopy is of a combined system. (We detect this spectroscopically for two stars described in the Appendix.) Observationally there is little we can do to account for the bias that these binaries will introduce into the HRD. However, most of the comparisons we will make throughout this paper are differential—the field compared to association, the Magellanic Clouds compared to the Milky Way. To whatever extent the binary frequency and distribution of mass ratios are the

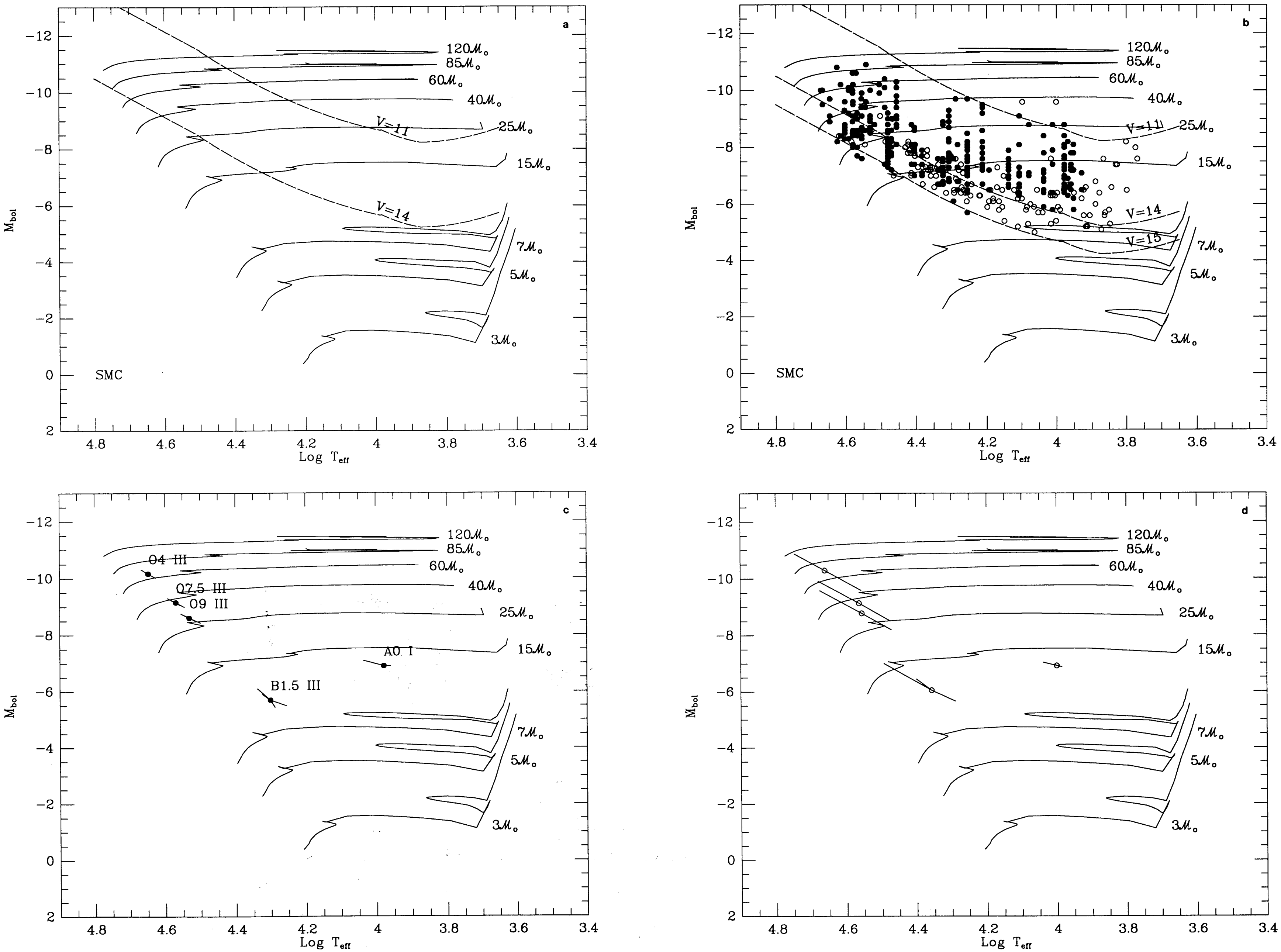

FIG. 1.—The theoretical evolutionary tracks of Schaller et al. (1992) are shown as solid lines, marked with the (initial) mass. (*a*) The two dashed lines show the location of stars with $V = 11$ and $V = 14$, given a modest reddening $[E(B-V) = 0.09]$ for stars at the distance of the SMC. (*b*) The stars from the Azzopardi & Vigneau (1982) catalog of SMC stars are added; solid points are stars placed on the basis of their spectra while open circles denote stars with only photometry. (*c*) The uncertainty in placing a star in the H-R diagram is shown for an error of one spectral type or one luminosity class. (*d*) The uncertainty in placing a star in the H-R diagram is shown for an error of 0.02 in $(B-V)_0$ and $(U-B)_0$.

same among these systems, the comparisons should remain unaffected by our inability to detect or correct for duplicity.

Finally, we note that when discussing massive star evolution there is often some nomenclature confusion. Throughout this paper we will use "main sequence" in its physical, rather than morphological, sense: the main sequence will refer to the part of the H-R diagram where stars are in the core–H-burning phase and are supposed to spend the majority (90%) of their lives. (This is roughly to the left of the first kink in each of the tracks in Fig. 1.) For the hotter stars, however, that means that stars of luminosity classes, V, III, and even I are "main-sequence" objects—i.e., thanks to historical usage we can speak of a "main-sequence supergiant" when referring to early-type stars (i.e., an O star of luminosity class "I" is still core–H burning). For the cooler stars this confusion does not exist: the only M-type stars on the main sequence are luminosity class "V," while all M supergiants are He-burning, evolved (non–main-sequence) stars. Whether or not the current evolutionary models show the correct width of the main-sequence is one of the "critical tests" referred to in the title.

2. CONSTRUCTING THE H-R DIAGRAMS

2.1. *Field Samples LMC/SMC*

Identification of probable members of both the Large and Small Magellanic Clouds have historically relied upon objective prism studies to detect OB and supergiant stars. Subsequent studies have added photometry and classification spectra.

For the LMC the primary source is the Sanduleak (1969) objective-prism survey, which identified 1272 problem members. Additional candidate members were found primarily by Brunet et al. (1975). These stars were compiled in the Rousseau et al. (1978) catalog of 1822 members, which included all the *UBV* and slit spectroscopy that was then available.

The Rousseau et al. (1978) catalog was updated by Fitzpatrick & Garmany (1990) to include subsequently published photometry (Isserstedt 1979, 1982) and spectroscopy (Crampton 1979; Walborn 1977; Conti, Garmany, & Massey 1986; Fitzpatrick 1988), and their machine-readable catalog is the starting point of the present work. It contains 1584 non-WR stars with complete *UBV* photometry, of which 430 stars have slit classification spectra. To this we will add 112 new spectral types, seven of which overlap with stars previously classified.

For the SMC the primary source is the Azzopardi & Vigneau (1975) catalog of 506 members. This list was updated and expanded by Azzopardi & Vigneau (1982) to include 524 stars, and includes all the *UBV* and spectroscopy available at that time. We began with a machine-readable version of this file made available through the NASA Astronomical Data Center, and updated it to include the literature-based and new spectral types given by Garmany, Conti, & Massey (1987). (Several omissions in their Table 1 have also been corrected.) There are 512 non-WR stars with *UBV* photometry, of which 285 have MK classification. (Of these stars, 95 were classified by Azzopardi 1987 on the basis of objective-prism work.) To this we will add 78 new spectral types, eight of which overlap stars previously classified.

2.2. *Incompleteness Fields*

In order to understand the limitations of the existing catalogs, and to determine incompleteness correction factors as a function of mass, we selected two regions in the LMC and one region in the SMC to probe considerably deeper than what was available in the literature. We had originally planned to obtain *UBV* wide-field CCD images of all three fields in order to select the bluest objects for spectroscopy, in much the same way we were using the photometry available in the literature. Technical problems prevented obtaining the images for the SMC, but we were able to use other material for selecting stars for spectroscopy.

2.2.1. *LMC*

B. Weller generously agreed to obtain *UBV* images with the 0.6 m Curtis Schmidt at CTIO of two test fields in the LMC. The exposures were made with a Metachrome-coated Thomson 1024 × 1024 CCD, with a scale of 1".8 pixel^{-1} (31' on a side). Pointing differences between the *U*, *B*, and *V* exposures reduced the effective field size to roughly 20' × 20'. The two fields selected were centered on $\alpha_{1975} = 05^h30^m.8$ and $\delta_{1975} = -71°04'$ (field 1) and $\alpha_{1975} = 05^h43^m.6$ and $\delta_{1975} = -67°56'$ (field 2). Contained within field 1 were the two "certain" OB associations LH 66 and 69, and within field 2 the "doubtful" OB association LH 114 (Lucke & Hodge 1970; Lucke 1972). The exposure sequence consisted of "short" and "long" exposures in each filter; integration times were 60 and 300 s in *U*, 40 and 240 s in *B*, and 20 and 120 s in *V*. The data were obtained during 1990 November.

Digital aperture photometry was run on all images, and a catalog of more than 2000 stars (1037 from field 1 and 1116 from field 2) was produced, after eliminating stars that were too crowded to have reliable photometry. Of these, 18 are in the updated Rousseau et al. (1978) catalog and could be used for converting the instrumental magnitudes to the standard *UBV* system. The fits of this transformation showed typical residuals of 0.02 mag in each filter, comparable to the photoelectric accuracy of the stars used as local standards. Of these, approximately 120 (field 1) and 30 (field 2) were "bright" ($V < 16$) and "blue" ($Q < -0.70$) enough for spectroscopy to be desirable; 16 of these were in the Rousseau et al. (1978) catalog and had either published or new spectroscopy. We successfully obtained spectra for most of the others.

2.2.2. *SMC*

In lieu of comparable photometry for the SMC, we instead selected candidate hot stars from a 4 m "grism" plate obtained by D. Crampton (in collaboration with A. Cowley, J. Hutchings, P. Conti, and P. M.) in 1982 November. The exposures were 15 minutes on baked IIIa-J emulsions, and were obtained with the "blue grism" (1600 Å mm^{-1}). The field examined for candidates measured 20' (NS) × 50' (EW), and was roughly centered on $\alpha_{1975} = 00^h58^m.9$ and $\delta_{1975} = -72°31'$. All in all, there were 90 candidate blue stars, of which 17 appeared in the Azzopardi & Vignau (1982) catalog. As described in the next section, we obtained spectra of roughly half of these.

In order to obtain photometry for these stars, we performed digital aperture measurements on the STScI scan of a 4 minute *V* plate that was obtained in support of the STScI Guide Star program. (The scan was kindly made available by M. Postman.) Conversion from the instrumental "photographic" magnitudes to standard *V* was obtained from the 17 Azzopardi & Vigneau (1982) stars, supplemented with a few stars from Massey et al. (1989b). (The outskirts of NGC 346 appear on the STScI scan but were not actually within the grism field.) The conversion spanned the range $V = 11.8$ to $V = 14.0$, with a second-order relation providing an rms residual of 0.14 mag; transformations for fainter or brighter stars therefore relied upon extrapolation. There was no color information, but the low and uniform reddening of the SMC permitted us to use a uniform correction (§ 2.4) to the V magnitudes.

2.3. *New Spectroscopy*

It was our goal to obtain reliable spectral types for all of the catalog stars with intrinsically blue colors, in addition to stars from our incompleteness fields. In all we succeeded in classifying 292 stars.

The spectra on which these classifications were based were obtained over the course of eight Magellanic Cloud observing seasons. Most observations were obtained with the CTIO 4 m telescope and various CCDs in the " blue air-Schmidt " camera between 1988 October and 1992 December. A few earlier spectra were obtained with the 2D-Frutti photoncounting system on both the CITO 4 m (1985 October) and 1.0 m (1988 January) telescopes. For all of these spectra the resolution was ~3 Å, and included at least the region from Si IV λ4089 to He II λ4686. Typically we aimed for a SNR of 75 per spectral resolution element (3 pixels).

The spectra were classified independently by two of the authors (C. C. L. and K. D. E.), usually with agreement to within one subtype. Larger differences were resolved by mutual inspection of the spectra. The classification precepts of the Walborn & Fitzpatrick (1990) atlas were used throughout.

These new spectral types are given in Tables 1–4. Of these 292 stars with new spectral types, 112 are LMC stars in the updated Rousseau catalog (Table 1), and 78 are SMC stars from the Azzopardi & Vigneau (1982) catalog (Table 2). An additional 67 stars were classified from the LMC incompleteness field (Table 3), and an additional 35 from the SMC incompleteness field (Table 4).

We list in Table 5 the 15 of these stars for which classification also exists in the literature. We see good agreement (roughly 1 subtype) for the LMC stars, with somewhat greater differences apparently for the SMC stars. For the latter, one must keep in mind that the primary source of spectroscopy has been the Azzopardi (1987) study, which relied upon objective-prism plates for producing MK types; although generally successful (as judged against our higher dispersion slit spectra), there are occasional significant disagreements (e.g., AV 476).

Our study has detected a number of previously unknown stars of very early type (O3–O4); we illustrate these and discuss them separately in the Appendix. (Not all of these are necessarily found in the pure field, of course.) In addition, we find a seemingly large number of Be stars listed in Tables 1–4. This is to be expected given our selection of stars for spectroscopy based upon very negative Q's. Be stars have abnormally blue $U-B$ colors since the U photometry is affected by Balmer continuum emission; Garmany & Humphreys (1985) chanced

TABLE 1

New Spectral Types in the LMC

Star	Type	Star	Type	Star	Type
Sk −65 2	B1 V	Sk −69 36	B0.5 III	B I 98	B1: V
Sk −65 17A	O7 III	Sk −69 50	O7 If	B I 103	B2 Ia
Sk −65 38	B1.5 III	Sk −69 119	B0.5 V	B I 104	O7: V
Sk −65 61	O7.5 V	Sk −69 216A	O6 V((f))	B I 105	B0.5 V
Sk −66 14	B1 Ib	Sk −69 217	O6.5 Iaf	B I 110	B1 III
Sk −66 18	O6 V((f))	Sk −69 218	O8 Ia	B I 121	B1 V
Sk −66 20	O6 III	Sk −69 229	O8.5 III	B I 128	O9 V
Sk −66 26	B1 III	Sk −69 238	O6.5 V((f))	B I 130	O8.5 V((f))
Sk −66 32	O6 III(f)	Sk −69 269	B1 III:	B I 142	B1 Ib
Sk −66 110	O9 I	Sk −69 300	B1 Ib	B I 145	O7 V
Sk −66 138	O9.5 III	Sk −69 303	O9 V	B I 164	O8 V((f))
Sk −67 38A	O8.5 V	Sk −70 13	O9 V	B I 178	B1 III
Sk −67 52	B0 I	Sk −70 33	O9.5 III	B I 184	B0.5 V
Sk −67 83	B0.5 III	Sk −70 79	B0 III	B I 185	O9 V
Sk −67 85	B0 III	Sk −70 109	B1 Ib	B I 187	B0 III
Sk −67 118	O7 V	Sk −71 8A	B1 Ib	B I 188	B0 Ia
Sk −67 119	O7 III(f)	Sk −71 13	Be (Fe II)	B I 189	O8 V((f))
Sk −67 158	B1 III	Sk −71 33	O9 Ve (Fe II)	B I 190	O8 V((f))
Sk −67 173	B0 Ia	Sk −71 34	WN3+O7.5	B I 216	B0.5 V
Sk −67 174	O8 V	Sk −71 35	B1.5 V	B I 217	B2 III
Sk −67 216	B0.5 V	Sk −71 39	O8 II(f)	B I 220	B0.5 III
Sk −67 241	B1 III	Sk −71 41	O8.5 I	B I 223	O9 V
Sk −67 250	O7.5 II(f)	Sk −71 42	B1 Iab	B I 228	B0.2 III
Sk −67 257	O8 III	Sk −71 44	B0 Ia	B I 235	Be
Sk −67 261	O8.5 III	Sk −71 45	O4 If	B I 237	O3 V
Sk −68 23A	B1 III	Sk −71 46	O4 If	B I 238	B1 Ib
Sk −68 47	O8 I	B I 7	B1 III	B I 243	B0.5 V
Sk −68 84A	B1 Ib	B I 24	O7 III	B I 244	Be(Fe II?)
Sk −68 95	B1 III	B I 28	B1 V	B I 245	O8.5 V
Sk −68 105	B0 Ia	B I 50	B1 Ib	B I 246	O9 V
Sk −68 117	B1 III	B I 53	B0 III	B I 249	O9.5 V
Sk −68 119	O9 V	B I 64	Oe	B I 253	O3 V
Sk −68 124	B0.5 V	B I 66	B1 V	B I 254	O8 V
Sk −68 127	B0.5 Ia	B I 78	B1 Ib	B I 260	O7 V((f))
Sk −68 132	B0 V	B I 81	Be(Fe II?)	B I 265	O6 V
Sk −68 134	B1 Ib	B I 85	O8 V	B I 267	B0 III
Sk −68 150	B2 II	B I 92	B3 V		
Sk −68 161	B1 III	B I 94	O7 V((f))		

TABLE 2

New Spectral Types in the SMC

Star	Type	Star	Type
AV 4	Be	AV 302	O8.5 V
AV 7	O7 V	AV 306	O9 V((f))
AV 9A	B2 V	AV 309	Be (Fe II)
AV 24	B1.5 V	AV 326	O9 V
AV 28	B1 I	AV 328	B0 V
AV 50	B0.5 III	AV 330	Be
AV 51	B0.5 V	AV 334	O8.5 V
AV 66	B0.5 V	AV 335	B1 V
AV 73	O8.5 V	AV 346	B2 V
AV 80	O7 III	AV 350	B1.5 V
AV 83	O7.5 Iaf+	AV 351	B0 V
AV 84	B1 V	AV 354	B1.5 V
AV 93	B1 Ib	AV 358	Be
AV 114	O7.5 V	AV 358A	B1 V
AV 124	O7.5 Ve	AV 361	B0 III
AV 128	O7 V	AV 363	B0.5 III
AV 130	B0.5 V	AV 364	B0.5 V
AV 144	Be	AV 376	Be (Fe II)
AV 148	O8.5 V	AV 386	B1.5 V
AV 189	O9 V	AV 395	B1 III
AV 196	B0.5 III	AV 402	Be
AV 201	B1 V	AV 409	Be
AV 203	B2.5 V	AV 423	O9.5 V
AV 208	O8 V	AV 425	B2.5 V
AV 209	B1 V	AV 427	B2.5 V
AV 214	B1 III	AV 435	O4 V
AV 217	B1 III	AV 436	O7.5 Ve
AV 233	B1 V	AV 456	O9.5 V
AV 236	B0 III	AV 457	Be (Fe II)
AV 255	O8 V	AV 458	Be
AV 259	B0.5 V	AV 460	Be (Fe II)
AV 261	O8.5 I	AV 464	O7.5 V
AV 266	B1 III	AV 471	B0.5 V
AV 268	B2.5 V	AV 476	O6.5 V
AV 271	B1.5 III	AV 480	Oe
AV 274	O7 Ve	AV 489	O8.5 V
AV 281	O7 III:	AV 491	O7.5 III:
AV 285	O8 V	AV 503	Be(Fe II?)
AV 301	B1 V	AV 506	B1 V

TABLE 3

SPECTRAL TYPES IN THE LMC INCOMPLETENESS FIELDS

Star	α_{1975}	δ_{1975}	V	$B-V$	$U-B$	Q	Type	Other ID
lmc1- 23	05 31 01.2	−71 17 37	14.25	−0.14	−0.94	−0.84	O9 V	B I 185
lmc1- 94	05 32 21.1	−71 14 51	13.67	0.05	−0.69	−0.73	B1 III	
lmc1- 104	05 31 09.4	−71 13 03	14.20	−0.11	−0.91	−0.83	O9.5 Vdbl?	
lmc1- 170	05 31 27.6	−71 10 06	14.14	−0.17	−1.01	−0.88	O8.5 V	
lmc1- 212	05 32 19.9	−71 09 17	14.73	−0.18	−0.87	−0.74	B1 III	
lmc1- 226	05 31 31.9	−71 09 25	15.34	−0.17	−0.87	−0.74	B1 V	
lmc1- 233	05 30 06.5	−71 09 23	13.77	−0.02	−0.84	−0.83	B0 V	
lmc1- 246	05 30 07.5	−71 09 45	15.04	0.06	−0.71	−0.76	B1: V	
lmc1- 249	05 30 15.4	−71 09 47	14.95	−0.10	−0.82	−0.74	B1 V	
lmc1- 263	05 30 13.2	−71 08 01	14.43	−0.05	−0.91	−0.88	O8 III(f)	
lmc1- 289	05 31 08.0	−71 08 49	14.22	0.11	−0.81	−0.90	O7.5 Ve	
lmc1- 294	05 30 24.5	−71 08 50	12.97	−0.09	−0.94	−0.88	B1.5 V	Sk −71 35
lmc1- 295	05 31 13.9	−71 08 53	14.90	0.16	−0.73	−0.85	O7 V	
lmc1- 390	05 30 42.8	−71 06 29	14.34	−0.16	−0.96	−0.84	O8 V	
lmc1- 402	05 30 39.6	−71 06 41	15.05	−0.06	−0.80	−0.76	B0 V	
lmc1- 408	05 29 33.3	−71 06 43	15.94	−0.21	−0.92	−0.77	B1 V	
lmc1- 413	05 30 31.7	−71 06 48	14.54	−0.17	−0.94	−0.82	B0.5 V	
lmc1- 415	05 30 42.8	−71 06 50	15.70	−0.03	−0.81	−0.79	Be	
lmc1- 436	05 29 43.8	−71 05 09	14.99	−0.16	−0.87	−0.76	B1 III	
lmc1- 437	05 31 00.3	−71 05 13	12.82	−0.07	−0.97	−0.91	O8.5 I	Sk −71 41
lmc1- 466	05 31 14.5	−71 05 42	13.38	−0.09	−0.95	−0.89	B0 III	B I 187
lmc1- 485	05 31 32.2	−71 04 00	14.48	0.02	−0.93	−0.94	B0: Ve (Fe II)	
lmc1- 493	05 30 31.3	−71 04 01	14.61	−0.06	−0.87	−0.83	O9 V	
lmc1- 504	05 31 39.0	−71 04 16	14.54	0.17	−0.81	−0.94	O7: Ve	
lmc1- 507	05 30 11.6	−71 04 13	14.17	−0.22	−1.04	−0.88	O8 V	
lmc1- 538	05 31 16.4	−71 04 33	12.91	−0.12	−0.98	−0.89	B0 Ia	B I 188
lmc1- 541	05 30 56.6	−71 04 35	14.52	−0.14	−0.99	−0.88	O9 V	
lmc1- 546	05 31 43.7	−71 04 43	13.88	0.05	−0.90	−0.93	O6 Ve	
lmc1- 548	05 30 11.6	−71 04 39	14.82	0.04	−0.68	−0.70	B1.5 V	
lmc1- 550	05 31 31.8	−71 04 44	12.90	−0.16	−1.01	−0.89	O5 V	
lmc1- 552	05 31 35.5	−71 04 45	11.60	−0.14	−1.00	−0.90	O4 If	Sk −71 45
lmc1- 555	05 30 31.6	−71 04 45	13.98	−0.04	−0.89	−0.86	B1 III	
lmc1- 556	05 31 41.5	−71 04 49	14.64	0.02	−0.88	−0.90	O8 V	
lmc1- 559	05 31 07.7	−71 04 49	11.19	−0.04	−0.85	−0.82	B1 Iab	Sk −71 42
lmc1- 572	05 31 29.7	−71 03 02	15.85	0.02	−0.80	−0.81	B1 III	
lmc1- 576	05 31 33.9	−71 03 04	14.53	−0.10	−0.94	−0.86	O7.5 V	
lmc1- 584	05 30 44.7	−71 03 04	14.62	−0.18	−1.07	−0.94	O7 V	
lmc1- 595	05 32 09.6	−71 03 20	13.27	−0.01	−0.89	−0.88	O4 If	Sk −71 46
lmc1- 598	05 31 17.6	−71 03 19	13.45	−0.06	−0.94	−0.90	O8 V((f))	B I 189
lmc1- 620	05 31 07.0	−71 03 36	13.34	0.01	−0.83	−0.85	B0 Ia	
lmc1- 642	05 31 10.6	−71 02 04	14.32	−0.17	−1.04	−0.92	O8 V	
lmc1- 662	05 30 50.5	−71 02 18	13.84	−0.08	−0.96	−0.91	B0.5 V	B I 184
lmc1- 702	05 30 54.9	−71 02 53	14.87	−0.06	−0.90	−0.86	B1 Ve	
lmc1- 703	05 32 28.8	−71 01 02	14.67	−0.20	−1.03	−0.88	O9 V	
lmc1- 713	05 31 04.4	−71 01 06	15.87	−0.17	−0.87	−0.75	B1 III	
lmc1- 715	05 30 56.6	−71 01 07	14.70	−0.14	−0.88	−0.78	B0.5 V	
lmc1- 716	05 31 30.5	−71 01 10	14.32	−0.11	−0.98	−0.90	O8 III((f))	
lmc1- 719	05 31 28.1	−71 01 11	13.04	−0.11	−0.99	−0.91	B0 Ia	Sk −71 44
lmc1- 732	05 31 17.4	−71 01 21	14.60	−0.17	−0.90	−0.78	B1 III	
lmc1- 733	05 30 59.6	−71 01 21	12.54	−0.13	−0.98	−0.89	O8 II(f)	Sk −71 39
lmc1- 756	05 30 58.8	−71 01 50	14.60	−0.19	−0.92	−0.78	B1 III	
lmc1- 838	05 29 52.8	−70 59 09	13.45	−0.21	−1.02	−0.87	WN3+O7.5	Sk −71 34
lmc1- 846	05 28 40.1	−70 59 13	13.82	−0.04	−0.97	−0.94	O9 Ve (Fe II)	Sk −71 33
lmc1- 859	05 30 47.8	−70 59 26	13.68	−0.16	−0.97	−0.86	B0.5 III	
lmc1- 929	05 29 57.4	−70 58 54	14.33	−0.20	−0.98	−0.84	O8.5 V	
lmc1- 930	05 32 09.5	−70 57 05	13.98	−0.15	−0.98	−0.87	B0.5 V	
lmc1- 938	05 32 00.5	−70 57 16	15.06	−0.15	−0.91	−0.80	B1 V	
lmc1-1015	05 31 53.5	−70 56 31	14.46	−0.17	−0.84	−0.72	B1 V	
lmc1-1027	05 31 43.7	−70 56 50	14.86	−0.14	−0.87	−0.76	B1 V	
lmc1-1044	05 30 03.9	−70 55 46	13.86	−0.19	−0.98	−0.84	B0 V	
lmc1-1052	05 30 41.3	−70 54 12	14.97	−0.16	−0.91	−0.79	B0.5 V	
lmc2- 8	05 41 35.7	−68 10 06	15.54	−0.01	−0.75	−0.74	B0.5 Ve:	
lmc2- 107	05 44 31.7	−68 05 54	15.03	−0.16	−1.01	−0.90	Be	
lmc2- 195	05 43 17.1	−68 02 47	15.42	−0.05	−0.98	−0.94	Be (Fe II)	
lmc2- 277	05 42 17.3	−68 01 13	13.82	−0.10	−0.82	−0.75	B2 II	Sk −68 150
lmc2- 341	05 44 18.5	−67 59 05	14.94	−0.22	−1.00	−0.84	B0.5 V	
lmc2- 486	05 44 52.2	−67 54 54	15.56	−0.16	−0.83	−0.71	B1 V	
lmc2- 488	05 42 18.1	−67 54 49	14.35	−0.08	−0.86	−0.80	Be	
lmc2- 627	05 43 15.4	−67 51 49	15.34	−0.10	−0.94	−0.86	B1.5: V	
lmc2- 675	05 43 18.1	−67 50 55	13.66	−0.25	−1.14	−0.96	O3 If + O	

TABLE 3—*Continued*

Star	α_{1975}	δ_{1975}	V	$B-V$	$U-B$	Q	Type	Other ID
lmc2- 682	05 43 20.7	−67 51 02	12.88	−0.14	−1.06	−0.96	O7.5 II(f)	Sk −67 250
lmc2- 696	05 43 23.3	−67 51 12	15.02	−0.19	−0.94	−0.80	B1 V	
lmc2- 702	05 43 17.4	−67 51 18	15.37	−0.08	−0.95	−0.89	B0 V	
lmc2- 703	05 43 13.4	−67 51 18	14.11	−0.23	−1.03	−0.86	O9.5 V	
lmc2- 716	05 43 16.3	−67 49 31	15.63	−0.20	−0.95	−0.81	B0 V	
lmc2- 718	05 43 40.2	−67 49 33	14.91	−0.18	−0.91	−0.78	B0.5 V	
lmc2- 724	05 43 29.8	−67 49 42	14.48	−0.23	−1.02	−0.85	B0 V	
lmc2- 733	05 43 29.0	−67 49 51	15.46	−0.23	−1.04	−0.88	B0 V	
lmc2- 747	05 43 11.7	−67 50 15	15.40	−0.19	−0.91	−0.77	B1 V	
lmc2- 750	05 42 46.7	−67 50 18	15.34	−0.01	−0.85	−0.85	Be (Fe II)	
lmc2- 755	05 43 21.0	−67 50 21	14.37	−0.22	−1.06	−0.90	O8 V	
lmc2- 759	05 43 25.7	−67 48 28	14.99	−0.19	−0.84	−0.70	B0.5 V	
lmc2-1008	05 41 49.8	−67 42 22	15.50	−0.02	−0.99	−0.98	Be (Fe II)	

upon Be stars in the Magellanic Clouds for the same reason. A number of these stars have Fe II in emission in their spectrum, a spectroscopic characteristic of "extreme Be" stars (Schild 1966); we will also discuss these further in the Appendix.

TABLE 4

SPECTRAL TYPES IN THE SMC INCOMPLETENESS FIELDS

Star	α_{1975}	δ_{1975}	V	Type	Other ID
smc- 1	00 56 42.0	−72 36 51	14.5	O9 V	AV189
smc- 2	00 56 53.7	−72 23 33	15.3	Be	
smc- 3	00 56 56.6	−72 23 02	15.3	B1 III	
smc- 4	00 57 03.4	−72 34 00	13.9	B0.5 III	AV196
smc- 5	00 57 30.6	−72 23 56	15.2	B0 V	
smc- 6	00 57 32.3	−72 25 52	15.5	O8.5 V	
smc- 7	00 57 38.0	−72 25 58	15.4	B0.5 V	
smc- 8	00 57 41.2	−72 25 23	15.0	B0.2 V	
smc- 9	00 57 42.9	−72 27 28	14.9	B1 V	
smc-10	00 57 48.2	−72 27 09	15.0	B1.5: V	
smc-11	00 58 11.7	−72 26 52	14.6	B1 III	AV217
smc-12	00 58 18.3	−72 36 40	15.0	B1 Ve	
smc-13	00 58 26.4	−72 30 27	15.1	B8 Ve	
smc-14	00 59 40.5	−72 27 14	14.7	Be	
smc-15	00 59 53.2	−72 32 51	15.3	Be (Fe II)	
smc-16	00 59 54.6	−72 33 43	14.3	O8.5 V	
smc-17	01 00 09.7	−72 36 58	14.1	O8.5 I	AV261
smc-18	01 00 15.3	−72 27 25	15.5	Be	
smc-19	01 00 20.4	−72 34 20	12.6	B1 III	AV266
smc-20	01 00 23.4	−72 27 00	15.2	B1.5 V	
smc-21	01 01 12.1	−72 34 38	14.7	O8 V((f))	
smc-22	01 01 22.9	−72 34 24	15.5	B0.5 V	
smc-23	01 01 27.0	−72 25 09	15.2	B1.5 V	
smc-24	01 01 30.3	−72 29 46	14.4	O8.5 V	AV302
smc-25	01 01 32.8	−72 34 18	15.7	B1: V	
smc-26	01 01 38.0	−72 34 12	15.8	B0: V	
smc-27	01 01 51.6	−72 32 38	13.4	Be (Fe II)	AV309
smc-28	01 01 57.7	−72 31 09	14.3	O9.5 V	
smc-29	01 01 57.7	−72 31 09	14.7	O8.5 V	
smc-30	01 02 05.1	−72 25 25	14.9	B1.5 V	
smc-31	01 02 10.6	−72 37 18	14.5	B0 Ve	
smc-32	01 02 20.2	−72 25 48	15.5	B1.5 V	
smc-33	01 02 21.1	−72 33 56	13.9	O9 V	AV326
smc-34	01 02 25.8	−72 32 45	14.6	O8.5 V	
smc-35	01 02 31.7	−72 33 28	13.7	B0 V	AV328
smc-36	01 02 40.6	−72 36 25	14.9	B1 V	AV335
smc-37	01 02 41.5	−72 25 57	15.2	O8 V	
smc-38	01 02 54.9	−72 34 28	15.2	B0.5 V	
smc-39	01 03 05.5	−72 26 05	15.4	B1.5 V	
smc-40	01 03 08.3	−72 35 30	13.6	B0.5 V	
smc-41	01 03 23.4	−72 28 52	13.7	B2 V	AV346
smc-42	01 03 25.7	−72 25 06	15.1	B1.5 V	
smc-43	01 03 30.3	−72 23 59	15.4	Be	
smc-44	01 03 46.0	−72 33 43	14.7	B1.5 V	AV354
smc-45	01 04 06.7	−72 30 28	15.4	B1.5 V	
smc-46	01 04 18.0	−72 31 57	14.3	B1 V	
smc-47	01 04 20.0	−72 31 19	13.5	Be (Fe II)	AV376
smc-48	01 04 23.9	−72 31 39	14.4	B0.5 V	

2.4. *Reddening and Distance Moduli of the Clouds*

With this database we are in a unique position to determine the reddening in both Clouds, and, more importantly, to redetermine the distance moduli to the Clouds. Both of these are necessary to achieving the goals of this paper: with a knowledge of the normal range in reddening within the Clouds, we can better remove stars from our sample whose photometry may be suspect. In addition, if we are to make valid comparisons between the HRDs of the Clouds with those of clusters and associations within our own Milky Way, it is important for the distances to be on the same system. Of course, we are cognizant of the fact that there is broader interest in knowing the distances of our nearest galaxian neighbors.

We begin by first rejecting all stars whose photometry disagrees significantly with that expected on the basis of its spectral type, as a self-consistency check. We do this by constructing a reddening-free parameter $Q = (U-B) - 0.72(B-V)$, and comparing it to the Q expected as a function of spectral type determined from the intrinsic colors of FitzGerald (1970). After we remove stars with significant discrepancies, we are left with 414 LMC stars and 179 SMC stars. In analyzing these data we have left in the few stars with physically impossible color excesses $[E(B-V) < 0.0]$ so as not to bias the average, but excluding these makes no difference.

TABLE 5

COMPARISONS OF NEW AND PREVIOUS SPECTRAL TYPES

Star	New	Literature Type	Literature Reference
		LMC	
Sk −67 85	B0 III	O9 III	Conti et al. 1986
Sk −67 250	O7.5 II(f)	O9 III	Conti et al. 1986
Sk −67 174	O8 V	O9 II:	Fitzpatrick 1988
Sk −70 13	O9 V	O9 III:	Fitzpatrick 1988
Sk −71 34	WN3 + O7.5	WN3 + O	Conti et al. 1986
Sk −71 42	B1 Iab	B2 Ia	Fitzpatrick 1988
Sk −71 45	O4 If	O4-5 III(f)	Walborn 1977
		SMC	
AV214	B1 III	B3 Ib	Azzopardi 1987
AV266	B1 III	B5 Iab	Humphreys 1983
AV268	B2.5 V	B0 Ia	Azzopardi 1987
AV271	B1.5 III	B0 Iab	Azzopardi 1987
AV328	B0 V	B1 Iab	Azzopardi 1987
AV423	O9.5 V	B0 Ia	Azzopardi 1987
AV471	B0.5 V	B2 Ib	Azzopardi 1987
AV476	O6.5 V	B2 II	Azzopardi 1987

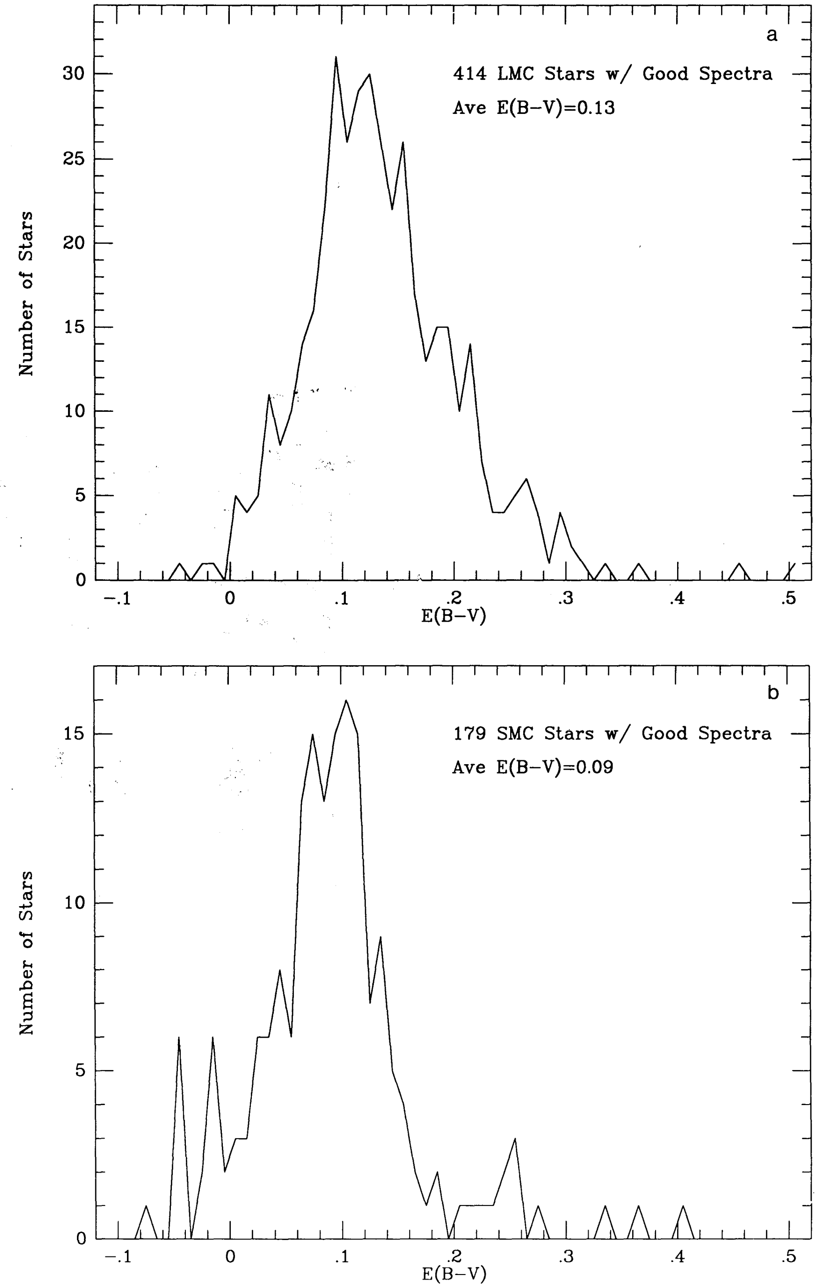

FIG. 2.—The number of stars with various amount of reddenings (inferred from their photometry and spectral types) are shown (*a*) for the LMC, and (*b*) for the SMC. Only stars whose reddening-free index Q agreed with that expected for its spectral type were included in these diagrams.

We show the number of stars as a function of $E(B-V)$ in Figure 2. For the LMC we find that the mean reddening is $E(B-V) = 0.13$, and, for the SMC, $E(B-V) = 0.09$, values which should be of little surprise to anyone. (See, for example, Tables I and II in McNamara & Feltz 1980, and Figure 4*a* in Caplan & Deharveng 1986.) However, what is new to this study is the fact that the spread in reddening for these (mostly field) stars is uniformly low in both the Large and the Small Clouds; we derive formal standard deviations of 0.07 mag for each galaxy. Over 50% of the stars in our LMC sample are found to have reddenings $0.09 < E(B-V) < 0.17$, while over 50% of the SMC sample are found to have reddenings $0.05 < E(B-V) < 0.12$.[4]

We can examine the reddening further by restricting the sample to the O-type and earliest B-type stars ($\log T_{\rm eff} > 4.40$). If we do this we obtain very similar results: for the LMC, the

[4] We note, for comparison, the study of LMC H II regions by Caplan & Deharveng (1986), who find reddenings corresponding to $E(B-V) \approx 0.08$–0.3. This is consistent with what we find here, with some stars in H II regions having higher reddenings that our sample, which is predominantly drawn from the field.

mean $E(B-V) = 0.14$ (213 stars), and for the SMC, the mean $E(B-V) = 0.10$ (104 stars). Thus we find that the earliest type stars do not show significantly higher reddenings than those of later type supergiants. This argues that even the younger massive stars have reddenings similar to their older siblings, which suggests therefore that massive stars may emerge from their cocoons more rapidly than sometimes supposed. We will return to this point shortly (§ 3.2).

We can now compute the distance moduli to the Clouds using the classical method of "spectroscopic parallax." (To the best of our knowledge this was first tried for the LMC using early-type stars by Crampton 1977.) We restrict ourselves to the earliest type stars (in practice, B0.5 and earlier); later than this, increasing errors are introduced with slight misclassification as to luminosity class. We adopt the calibration of M_V with spectral type from Conti et al. (1983) for the O-type stars, and from Humphreys & McElroy (1984) for the early B-type stars, and for each star in our sample compute $V - A_V - M_V$, where $A_V = 3.1E(B-V)$, determined for each star individually using its observed $B-V$, and the intrinsic $(B-V)_0$ expected for its spectral type based upon FitzGerald (1970). In analyzing our data, we have also restricted our samples to only those stars meeting the strict criteria that their color excesses be within 1 σ of the means determined above, and also that the individually derived spectroscopic distance moduli be within 1 mag of the "canonical" values of 18.6 and 19.1 for the LMC and SMC, respectively. This latter has the effect of safeguarding against stars misclassified as to luminosity class.

When we do this we find a distance modulus to the LMC of 18.41 ± 0.04 (s.d.m.) based upon 132 stars of all luminosity classes, and a distance modulus to the SMC of 19.09 ± 0.06 based upon 68 stars. We have attempted to refine these numbers by restricting ourselves to stars only of luminosity class III and V or just V, but arrive with similar numbers in each case (see Table 6). In practice, we find that exactly what rejection criteria we use changes the derived distance moduli by at most 0.1 for the LMC, and by 0.2 for the SMC; i.e., the values quoted here are stable to within a few times the formal standard deviations of the mean, as expected. We adopt values for the true distance moduli of the LMC and SMC of 18.4 and 19.1, respectively.

These numbers are essentially identical to those we have previously derived using spectral types for stars in the Clouds. Conti, Garmany, & Massey (1986) find a distance modulus of 18.3 ± 0.4 for the LMC, while Garmany, Conti, & Massey (1987) find a value of 19.1 ± 0.1 for the SMC, using the data then available, and applying fewer restrictions to their samples. We note that using this method implicitly assumes that the calibration of M_V as a function of spectral type is the same in the Clouds as it is the Galaxy.

TABLE 6

DERIVED DISTANCES

Includes	Number of Stars	Distance Modulus
	LMC	
V, III, I......	132	18.41 ± 0.04
V, III	68	18.33 ± 0.05
V............	29	18.26 ± 0.09
	SMC	
V, III, I......	68	19.09 ± 0.06
V, III	40	19.06 ± 0.07
V............	25	19.05 ± 0.10

These distances compare quite favorably with other recent determinations; i.e., those using planetary nebulae and SN 1987A. Jacoby, Walker, & Ciardullo (1990) derive distance moduli of 18.44 ± 0.08 and 19.09 ± 0.3 for the two Clouds based upon the [O III] λ5007 flux of the brightest planetary nebulae. (See their Table 6 for comparison with other determinations.) The zero-point of the planetary nebula distance scale is set by the Cepheid distance of M31, while the O star spectral type–to–M_V calibration is based upon Galactic OB associations whose ages are too young to contain any Cepheids but whose distances have primarily been determined by main-sequence fitting and/or using spectroscopic parallax of B stars. (See discussion in Conti & Alschuler 1971; Walborn 1972; Conti & Burnichon 1975). Subsequent to the Jacoby et al. (1990) discussion, Panagia et al. (1991) derive a value of 18.50 ± 0.13 from SN 1987A for the LMC, also in excellent accord with the value derived here from early-type stars. Given the differences in the various assumptions that have gone into these three determinations, the agreement may be considered fortuitous, if not fantastic, and at least indicates that the derivation of distances to Galactic clusters and associations using spectroscopic parallax of OB stars is consistent with the extragalactic distance scale.

2.5. *Transformation to* log $T_{\rm eff}$ *and* $M_{\rm bol}$

In our studies of Magellanic Cloud associations we have previously described the transformations from photometry to log $T_{\rm eff}$ and bolometric corrections (BCs), but these were derived primarily for luminosity class V objects (see Table IX in Massey et al. 1989b). The field population will contain many stars of luminosity classes I and III as well, and so here we derive new transformation equations.

For stars with MK spectral types, we use the calibration of spectral type with effective temperature and bolometric correction from Chlebowski & Garmany (1991) for the O-type stars and that compiled by Humphreys & McElroy (1984) for later types. The photometry is corrected for extinction using the spectral-type to $(B-V)_0$ calibration of FitzGerald (1970), and assuming $A_V = 3.1E(B-V)$.

In order to deal with the stars for which no spectral types are known, we used the same calibration to derive transformation equations between the intrinsic colors and effective temperatures and bolometric corrections. For the hotter stars, we can use the reddening-free parameter Q directly for these transformations; later than B5 or so, however, Q becomes degenerate with spectral type, and we must estimate the intrinsic colors using the average reddenings found above. In determining which transformation equations to use, we first determine the probable luminosity class of the star using the mean reddening and a relationship between M_V and intrinsic color for the various luminosity classes. We summarize our transformations in Table 7.

Although these transformations differ in detail from those we have used previously, we find that there is relatively little disagreement between the various calibrations. In Figure 3 we show the transformations from Table 7 for $(B-V)_0 < 0.0$. (Nearly all the stars we are concerned with in this paper meet this criterion.) In order to see how well our transformations compare to those from other data, we also show the intrinsic colors, log $T_{\rm eff}$, and BCs from Schmidt-Kaler (1982). In addition, we have included as dashed lines the fits adopted by

TABLE 7

TRANSFORMATION EQUATIONS FOR PHOTOMETRY

$Q < -0.4$:	
$\log T_{\rm eff} = 4.055 + 0.041Q + 0.6514Q^2$	(V)
$\log T_{\rm eff} = 4.342 + 1.105Q + 1.4793Q^2$	(III)
$\log T_{\rm eff} = 3.859 - 0.619Q - 1.4160Q^2 - 1.6405Q^3$	(I)
$(B-V)_0 = -0.013 + 0.325Q$	(III, V)
$(B-V)_0 = -0.016 + 0.342Q$	(I)
$A_V = 3.1E(B-V)$	
$Q \geq -0.4$:	
$(B-V)_0 = (B-V) - \overline{E(B-V)}$	
$\log T_{\rm eff} = 3.981 - 1.177(B-V)_0$	(V, III, I, $(B-V)_0 < 0.0$)
$\log T_{\rm eff} = 3.985 - 0.563(B-V)_0$	(III, I, $0.0 \leq (B-V)_0 < 0.2$)
$\log T_{\rm eff} = 3.914 - 0.230(B-V)_0$	(III, I, $0.2 \leq (B-V)_0 < 2.0$)
$\log T_{\rm eff} = 3.45$	(III, I, $2.0 \leq (B-V)_0$)
$A_V = 3.1\overline{E(B-V)}$	
All:	
$\mathrm{BC} = -1092.40 + 596.295 \log T_{\rm eff} - 81.3982(\log T_{\rm eff})^2$	($\log T_{\rm eff} < 3.650$)
$\mathrm{BC} = -141.04 + 72.763 \log T_{\rm eff} - 9.3752(\log T_{\rm eff})^2$	($3.650 \leq \log T_{\rm eff} < 3.870$)
$\mathrm{BC} = -270.00 + 141.447 \log T_{\rm eff} - 18.5119(\log T_{\rm eff})^2$	($3.870 \leq \log T_{\rm eff} < 3.985$)
$\mathrm{BC} = -72.61 + 39.070 \log T_{\rm eff} - 5.2493(\log T_{\rm eff})^2$	(I, $3.985 \leq \log T_{\rm eff} < 4.500$)
$\mathrm{BC} = -45.36 + 25.891 \log T_{\rm eff} - 3.6662(\log T_{\rm eff})^2$	(III, $3.985 \leq \log T_{\rm eff} < 4.500$)
$\mathrm{BC} = -9.93 + 9.130 \log T_{\rm eff} - 1.6940(\log T_{\rm eff})^2$	(V, $3.985 \leq \log T_{\rm eff} < 4.500$)
$\mathrm{BC} = 22.71 - 4.799 \log T_{\rm eff} - 0.20839(\log T_{\rm eff})^2$	($4.500 \leq \log T_{\rm eff}$)

Flower (1977). Although his relations were derived quite differently than ours, the agreement is good. The only significant differences appear for the B-type giants (luminosity class III); here the Humphreys & McElroy (1984) data used in determining our transformations suggest a somewhat cooler temperature as a function of intrinsic color than that of either Schmidt-Kaler or Flower.

In determining these transformations we have assumed that the Galactic calibration of Q to spectral-type to effective temperature and bolometric correction applies equally well to the Magellanic Clouds. We will note in § 3.2 that model atmosphere predictions are consistent at least with the conversion of Q to effective temperature, and, by implication, bolometric correction. In their study of the SMC cluster NGC 346, Massey et al. (1989b) find a hint that the O stars have slightly redder (0.05 in Q) colors than their Galactic counterparts. We can use the much larger data set available to us here to explore this question further by comparing Q determined from their photometry with the Q expected on the basis of their spectral type, using the Galactic calibration of FitzGerald (1970). We find for the LMC a difference $Q_{\rm ph} - Q_{\rm sp} = 0.015 \pm 0.003$ ($Q < -0.4$). For the SMC there is a suggestion that the colors of star are in fact slightly redder than their Galactic counterparts: we find $Q_{\rm ph} - Q_{\rm sp} = 0.049 \pm 0.006$, in agreement with Massey et al. (1989b). At $Q = -0.4$ this would amount to a potential systematic error of 0.02 in $\log T_{\rm eff}$, or less than 0.1 of a magnitude in the bolometric luminosity.

2.6. *Separating Field Stars from Association Members*

Within the Rousseau et al. (1978) and Azzopardi & Vigneau (1982) catalogs we can expect to find a mixture of OB association members and true field stars. How do we separate these? The OB associations of the LMC have been identified by Lucke & Hodge (1970), while those of the SMC have been identified by Hodge (1985). In order to evaluate the "field status" of each star, we compared its coordinates to the centers and sizes of the cataloged associations; we then determined the relative distance from the nearest association boundary. An inspection of the finding charts given by Lucke (1972) suggests that the listed sizes of the associations are realistic, and that the positions both of the cataloged stars and OB associations are known sufficiently precisely to determine if a star is inside or outside an association boundary.

How do we know that these "field stars," however, were in fact born in the field? We can address this by considering how far a massive star might travel during its short lifetime. Churchwell (1991) suggests that velocities of 3 km s^{-1} relative to the parent molecular cloud may be typical, citing the study of Orion by Zuckerman (1973). In 10 Myr we would expect a star to travel 1×10^{16} km, or 30 pc. At the distance of the LMC, this corresponds to an angular distance of 2′. We will therefore adopt the criterion that for a star to be considered a true field star, it must lie at least 2′ from the nearest association boundary.

We feel that this limit is fairly conservative for a variety of reasons, although we will see in § 3.1 than our results do not change if we make our selection criterion even more stringent. Here we merely note the following: (*a*) The average separation would be less than this by a factor of $2^{1/2}$ due to projection effects. (*b*) The 10 Myr lifetime is considerably longer than the lifetime of the most massive stars (<3 Myr); it is these stars that we are primarily interested in knowing whether or not they occur in the field. Even 25 $M_\odot$ stars (the limit for which we will compute an IMF), have main-sequence lifetimes which are only 7 Myr long. (*c*) In determining the separation from the nearest OB association boundary, we have approximated each association as round, using the *larger* of the two dimensions listed by Lucke & Hodge (1970) and Hodge 1985). (*d*) Finally, we have included even OB associations classified "unlikely" in determining the field status of our stars.

3. RESULTS

In this section we will proceed to answer the questions posed in the introduction: (1) Are stars as massive as those found in associations produced in the field? (2) Are very massive stars produced as commonly in the field as in associations (relative

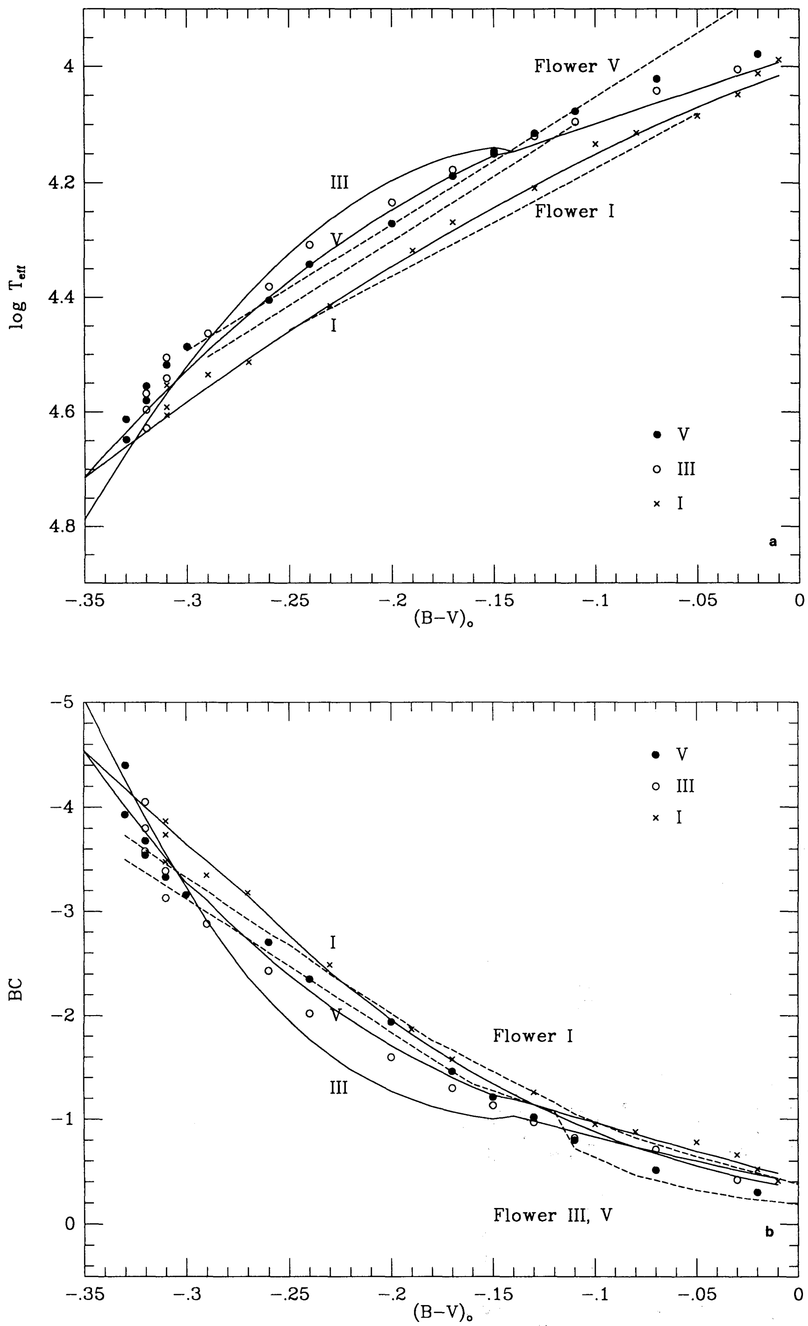

FIG. 3.—(*a*) The adopted transformations (Table 7) for log $T_{\rm eff}$ as a function of $(B-V)_0$ are shown as the three solid lines; these fits come from the calibrations (not shown) of Chlebowski & Garmany (1991), Humphreys & McElroy (1984), and the intrinsic colors of FitzGerald (1970). The points shown are those of a different calibration (Schmidt-Kaler 1982), and were not used in the fits; they are included merely to show what differences exist among the various calibrations. The three dashed lines show the transformations found by Flower (1977), who employed a third calibration. (*b*) Same as (*a*), but for the relationship between the bolometric correction (BC) and $(B-V)_0$.

to stars of lower mass); i.e., is the initial mass function the same for the field stars as for the association stars? In order to answer the second question with any confidence we must first answer the third goal of this paper: (3) How good are the evolutionary models? We can test this by comparing the distribution of our stars within the HRD to that predicted by modern low-metallicity models. Finally, we use our data to investigate the more advanced stages of stellar evolution: what mass stars become WR stars, and does this limit depend upon metallicity?

We begin by showing in Figures 4*a* and 5*a* the HRDs of the complete LMC and SMC data sets. We have overlaid the evolutionary tracks of Schaerer et al. (1993) (computed for $z = 0.008$) for the LMC and those of Schaller et al. (1992) (computed for $z = 0.001$) for the SMC. For clarity we have truncated the tracks at the beginning of the WR stage (we have

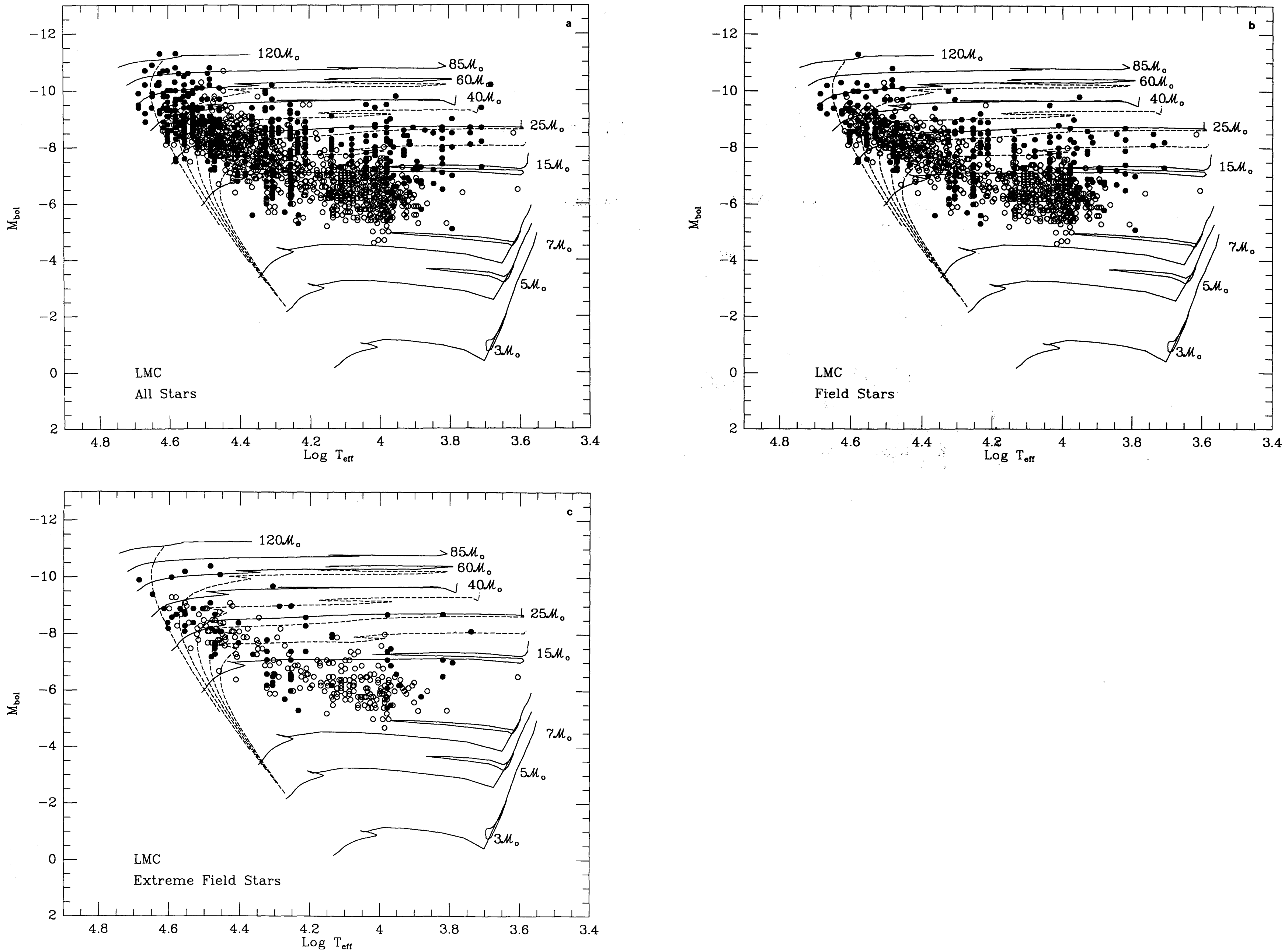

FIG. 4.—The HRDs for the LMC data set. Filled circles have been placed in the diagram based upon their spectral type; open circles have been located based only upon photometry. The $z = 0.008$ evolutionary tracks of Schaerer et al. (1993) are shown as solid lines, labeled by the initial mass. The dashed lines show isochromes at 2 Myr intervals. (*a*) Includes all the stars in the Rousseau et al. (1978) catalog. (*b*) Includes only stars we considered to be true field stars, based upon their location $> 2'$ (30 pc) from the boundary of an OB association or star cloud. (*c*) Includes only stars located $> 20'$ (300 pc) from the nearest border.

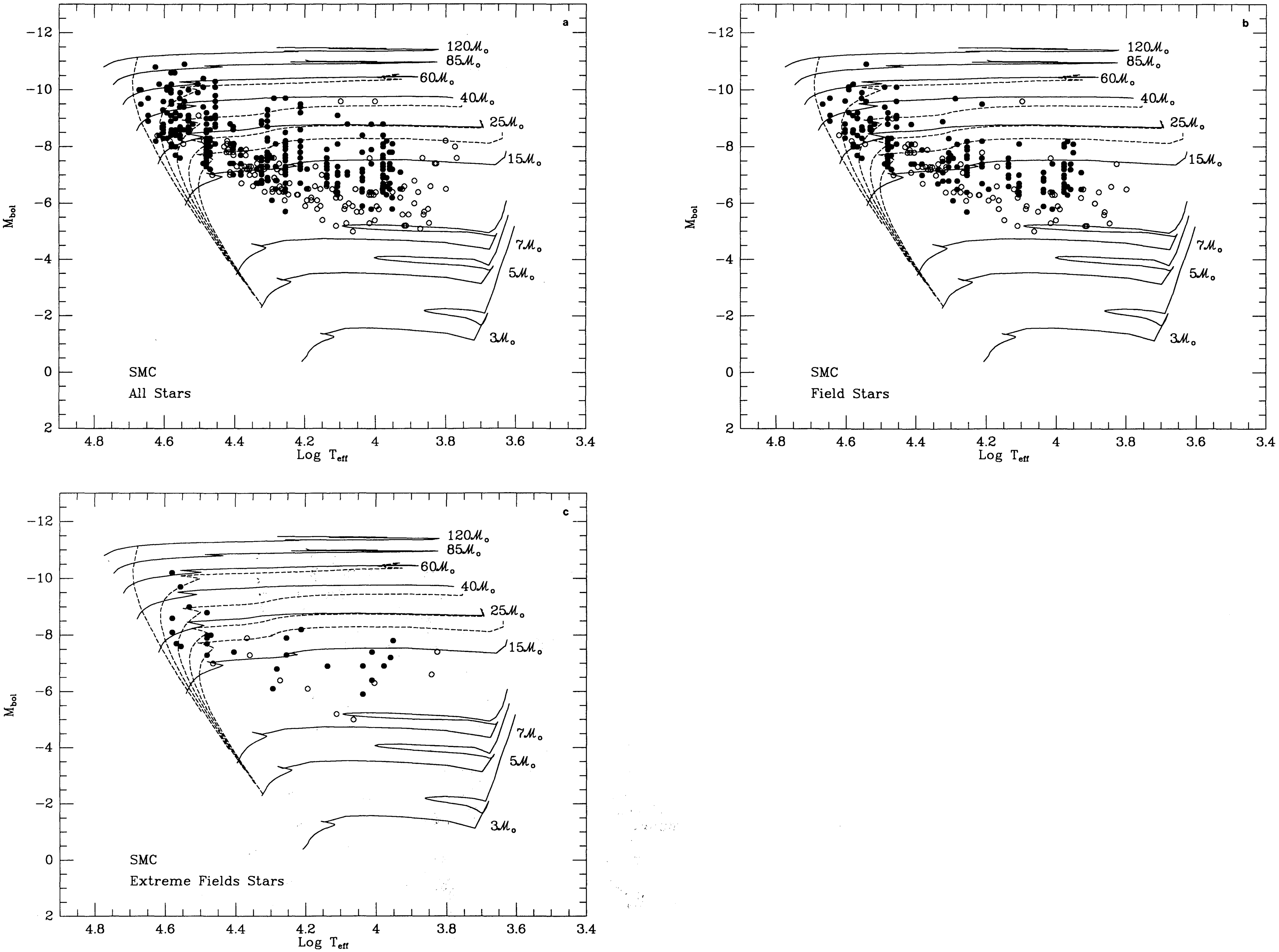

FIG. 5.—The HRDs for the SMC data set. Symbols have the same meaning as in Fig. 4. We show the $z = 0.001$ evolutionary tracks of Schaller et al. (1992). (*a*) Includes all the stars in the modified Azzopardi & Vigneau (1982) catalog. (*b*) Includes only stars we considered to be true field stars, based upon their location $> 2'$ (38 pc) from the boundary of an OB association or star cloud. (*c*) Includes only stars located $> 15'$ (300 pc) from the nearest border.

explicitly excluded the WR stars from these diagrams), which these authors define as the point where the surface composition of hydrogen drops below 0.4 for hot stars (log $T_{\rm eff} > 4.0$). We have labeled each evolutionary track by its initial mass; the evolutionary code explicitly takes mass loss into consideration, with the most massive stars losing 20%–30% of their initial mass during the interval shown. The dashed lines are isochrones, at 2 million year intervals. We are extremely grateful to G. Meynet for kindly making these tracks available in electronic form, as well as for providing the FORTRAN program which computes the isochrones.

In Figures 4*b* and 5*b* we show the HRDs for the field star sample defined as described above: stars which lie more than 2′ away from the boundary of an OB association or probable "star cloud."

3.1. *The Most Massive Stars Found in the Field Compared to the Associations*

We can now answer the first of the questions posed above by inspection of Figures 4 and 5: *we do find equally massive stars in the field sample as in the complete sample.* The field contains a healthy number of stars with masses in the 60–85 $M_\odot$ range. This upper mass limit is indistinguishable from what is found from the Galactic and Magellanic Cloud associations studied with the same methods and assumptions (Massey 1993).

We have argued above that our selection of field stars from our larger data set has been fairly conservative—most of the stars in this diagram must have been born in the field, given their present location. We can, however, test our conclusion further by making an even more cautious requirement. There are, after all, "runaway" O stars known, O-type stars whose peculiar space velocity is above 30–40 km s^{-1}. These stars are quite rare: in their study of the radial velocities of over 200 O-type stars, Conti, Leep, & Lorre (1977) identify six stars with peculiar velocities greater than 40 km s^{-1}. None of these had peculiar velocities as great as 100 km s^{-1}, and we will adopt this value as our limit; this is also consistent with the peculiar (total) space velocities given by Blaauw (1993). The ages of the highest mass stars shown in Figures 4, 5*a*, 5*b* are less than 3 Myr, and accordingly we define our population of extreme field stars as those separated from the nearest association boundary by 100 km s^{-1} × 3 Myr, or 300 pc, corresponding to 20′ in the LMC, or 15′ in the SMC. (We note that most runaways are believed to be ejected as its binary companion becomes a supernova; thus, the effective time a star spends as a runaway will actually be only a small fraction of its life.) We show the HRDs of these extreme field stars in Figures 4*c* and 5*a*. Once again, we are forced to the conclusion that the field is capable of producing stars as massive as those found anywhere.[5]

[5] Garmany (1990) has noted that 100% of the known O3–4 stars in the Milky Way are found in clusters and associations, while a few of the earliest type Magellanic stars found here are believed to be "true" field stars. (We note explicitly the O3 stars Sk −67 22, Sk −67 274, Sk −68 137, BI 237, and BI 253, and the O4 stars Sk −67 69, Sk −67 166, and Sk −70 60; finding charts can be found in Sanduleak 1969 and Brunet et al. 1975 for those wishing to check the degree of isolation for themselves.) Does this represent a true difference where the Milky Way? We do not believe so, although additional data on the Galactic field stars would clearly be a welcome addition. We note, for instance, that subsequent to Garmany (1990), Garmany & Vacca (1991) reclassified 32 Galactic OB stars, and noted that several of these are quite isolated, including BD −11°4620, which is an O4 star.

3.2. *The Models and Isochrones*

Before deriving the IMF of the Clouds, we next consider how well the evolutionary models match reality. Is the distribution of stars within the HRD the same as that predicted by the evolutionary models? One very stringent test is to see if the spacing of the isochrones within a given mass bin matches the distribution of stars within that bin; i.e., given our assumption of short-term steady state star formation, we would expect the same number of stars with ages between 2 and 4 million years as between 6 and 8 million years for a given mass—providing the lifetime of a star of that mass was longer than 8 Myr.

We give in Table 8 the number of stars per 2 Myr interval for the field stars in the LMC and SMC, where we have counted between adjacent isochromes in Figures 4*b* and 5*b*. We also list the main-sequence lifetimes based upon these models; in computing the "average" age for a given mass bin, we have weighted the lifetimes of the lower and upper mass bounds of a bin assuming a Salpeter (1955) mass distribution. We enclose in parentheses the number of stars in bins that occur beyond the expected lifetime of a star in that mass range.

In interpreting the numbers in Table 8 there are two effects to keep in mind. As we discussed in the introduction, a *V*-limited catalog will be selectively incomplete for the hotter stars; in this case, this also means the younger stars. The incompleteness will become increasingly severe as we go down in mass. Therefore, for the LMC field star 25–40 $M_\odot$ mass bin we might expect that the reason we see only 27 stars with ages between 2 and 4 Myr but 85 stars with ages between 4 and 6 Myr years is that the younger (hotter) 25–40 $M_\odot$ are simply too faint to have been picked up in the surveys on which our sample is based (see Fig. 1*b*). However, there are fewer stars (50) in the 6 to 8 Myr age bracket because the average lifetime

TABLE 8

DISTRIBUTION OF STARS WITH ISOCHRONES

Masses ($M_\odot$)	Lifetime (Myr)	Isochrones (Myr) 0–2	2–4	4–6	6–8	8–10	10–12	12–14
		LMC Field Stars						
85–120	3.0	0	(3)	(0)	(0)	(0)	(0)	(0)
60–85	3.6	2	(7)	(0)	(0)	(0)	(0)	(0)
40–60	4.1	4	32	(7)	(0)	(0)	(0)	(0)
25–40	7.2	2	27	85	(50)	(0)	(0)	(0)
15–25	12.7	2	0	5	88	108	90	(27)
		SMC Field Stars						
85–120	3.2	0	(1)	(0)	(0)	(0)	(0)	(0)
60–85	3.9	0	2	(0)	(0)	(0)	(0)	(0)
40–60	4.3	0	8	(2)	(0)	(0)	(0)	(0)
25–40	7.2	0	2	31	(14)	(0)	(0)	(0)
15–25	13.0	0	0	3	2	25	33	(13)
		LMC Incompleteness Field						
60–85	3.6	0	(1)	(0)	(0)	(0)	(0)	(0)
40–60	4.1	1	4	(0)	(0)	(0)	(0)	(0)
25–40	7.2	1	2	3	(2)	(0)	(0)	(0)
15–25	12.7	2	2	3	3	6	6	(0)
		SMC Incompleteness Field						
40–60	4.3	0	1	(1)	(0)	(0)	(0)	(0)
25–40	7.2	0	0	5	(1)	(0)	(0)	(0)
15–25	13.0	0	0	0	2	3	5	(2)

of a 25–40 $M_\odot$ star is only 7.2 Myr. Similarly for the 15–25 $M_\odot$ bin we see few stars younger than 6–8 Myr; this is in accord with our expectation that the incompleteness will be considerably worse within this mass bin. We do see roughly the same number of stars in the 6–8, 8–10, and 10–12 Myr bins, but many less in the 12–14 Myr bin; this is consistent with the predicted 12.7 Myr lifetime.

With this in mind we see that the distribution of stars reflected in this table is actually quite consistent with the models, at least to the best of our ability to test these models with the general LMC and SMC field data. In particular, there are never any stars seen in the age bin beyond the star's expected lifetime. This confirms the visual impression one has in studying Figures 4*b* and 5*b*: stars in general extend to the first kink in the evolutionary tracks but not much beyond this, at least for the stars with masses above 15 $M_\odot$. (For stars with lower masses we do see a number of stars beyond the theoretical end of the main sequence; this is in accord with the Fitzpatrick and Garmany 1990 study, discussed below.)

Comparisons between observed HRDs and model predictions in the 1970s and 1980s led to the phrase "main-sequence widening" to describe the fact that the models did not seem to go far enough to the red during core–H burning; a good example of this problem can be found in Figure 4 of Humphreys & McElroy (1984) for the LMC where they have used the Maeder (1981) evolutionary tracks. However, with our more complete data set of the Magellanic Clouds and the newer models of Maeder and his colleagues, this problem appears to have been solved: we run out of stars right as we expect, indicating that the relative ages across the HRD are correct. This is a very encouraging result, given the narrowing of the MS that occurs with lower metallicity models (see discussion in Schaller et al. 1992): even the SMC models show good agreement with the observations.

However, the rarity of the youngest (<2 Myr) massive stars in such diagrams, first noted by Garmany, Conti, & Chiosi (1982) for their mixed population of field and association Galactic stars, is consistent with our LMC and SMC field data as well. Maeder (1993) cites some possible physical explanations for this absence, including the one originally suggested by Garmany et al. (1982) that the youngest O-type stars are still buried in their parent clouds, emerging only after 1–2 Myr. However, we first need to establish if the effect is really present in our data, and not simply due to selection effects. From our Figure 1*a* we judge that a $V \approx 14$ completeness limit should begin to affect the number of stars younger than 2 Myr starting around the 60 $M_\odot$ bin; higher than this in mass we do not have many stars, although the statistics for the 60–85 $M_\odot$ bin are suggestive that there is such a lack.

We can address this better using our incompleteness test fields as these will not suffer from incompleteness at early ages due to selection effects as they go much fainter than the cataloged members. We show in Figure 6 the HRDs for these fields, where we have marked with an asterisk those stars which were known previously to our study. We have used a slightly different criterion here for "field stars": we insist only that a star not be within the boundaries of a probable OB associations in order to qualify, as we are only trying to avoid areas where the catalogs are incomplete. (A more stringent criterion results in fewer stars and poorer statistics but otherwise does not change our result.) We see qualitatively from inspection of these diagrams that most of the newly found stars are indeed younger compared with those previously known within a given mass bin. These additional stars have been added simply by going deeper than the existing catalogs, as expected from the selection effects previously discussed. The scant number of very massive stars—due to the small areas covered by the incompleteness fields—makes a quantitative assessment difficult, however. The problem is exacerbated by the skinniness of the region in log $T_{\rm eff}$ between the start of the tracks (the ZAMS) and the first (2 Myr) isochrone.

For the SMC we do not see any stars younger than 4 Myr, either newly discovered or previously known. However, the area covered is small, and the data in this diagram may simply suggest that our grism search for hot stars was not as complete as we would have liked; certainly there is little comparison between Figures 6*a* and 6*b*. We note that in our H-R diagram of the SMC cluster NGC 346 (Massey 1989b, and rederived below) we do find plenty of stars between 2 and 4 Myr in age, but none younger than 2 Myr; of course, this latter fact may simply mean that the cluster itself is 2 Myr old.

While we cannot determine if the scarcity of the youngest massive stars is due to a real effect or selection, we can at least test to see if the youngest stars which are present are more heavily reddened than their older counterparts. Wood & Churchwell (1989) have suggested that the first 10%–20% of an O star's main-sequence lifetime is spent embedded in its parent cloud. Thus a massive star arriving on the ZAMS would still be surrounded by circumstellar material. However, we would then expect that the youngest massive stars which *are* in our sample should have higher extinction. We use the data set previously discussed for determining the average reddening and distance moduli of the Clouds (§ 2.4) but restrict ourselves only to the most massive (>40 $M_\odot$). If we further divide this set into stars that are younger or older than 2 Myr, we find identical reddenings: $E(B-V) = 0.17 \pm 0.02$ (10 stars) for the "young and massive" group, and $E(B-V) = 0.17 \pm 0.01$ (67 stars) for the older stars. We believe this renders unlikely the suggestion that the youngest massive stars do not break out of their cocoons until well advanced in core-H burning.

Although we have concluded that the models do a good job of matching the distribution of stars at higher masses, we do see evidence of two potential problems at lower luminosities. The first of these is revealed in Figure 6*a* where it is apparent that the ZAMS lies considerably to the left of the data at lower luminosities ($M_{\rm bol} \approx -4$). The blue edge of the distribution of the majority of the stars is displaced from the theoretical ZAMS by about 0.075 dex in temperature by $M_{\rm bol} = -4$. While our data for the SMC does not go deep enough to detect if there is a problem here, we will see in our reanalysis of the NGC 346 below that the problem is roughly twice as severe (0.15 dex) for the SMC.

Could this gap between the theoretical ZAMS and the distribution of points be due to the fact that we have adopted the Galactic calibration for the conversion of color to effective temperature? For stars in the log $T_{\rm eff} \approx 4.3$ regime, a gap this large would require that the calibration be different for the Magellanic Cloud stars by 0.09 in Q for the LMC and by 0.19 for the SMC; these are considerably greater than the hint of a color difference noted in § 2.5. These differences would require only that a star of spectral type B2 would have colors characteristic of Galactic B1–B1.5 V stars for the LMC and of B0.5–B1 stars for the SMC. Nevertheless, for early and mid-B stars metal lines are so few and weak it would be some surprise that a change in metallicity would have even this much effect on the color of a star. The Kurucz (1979) ATLAS models for log

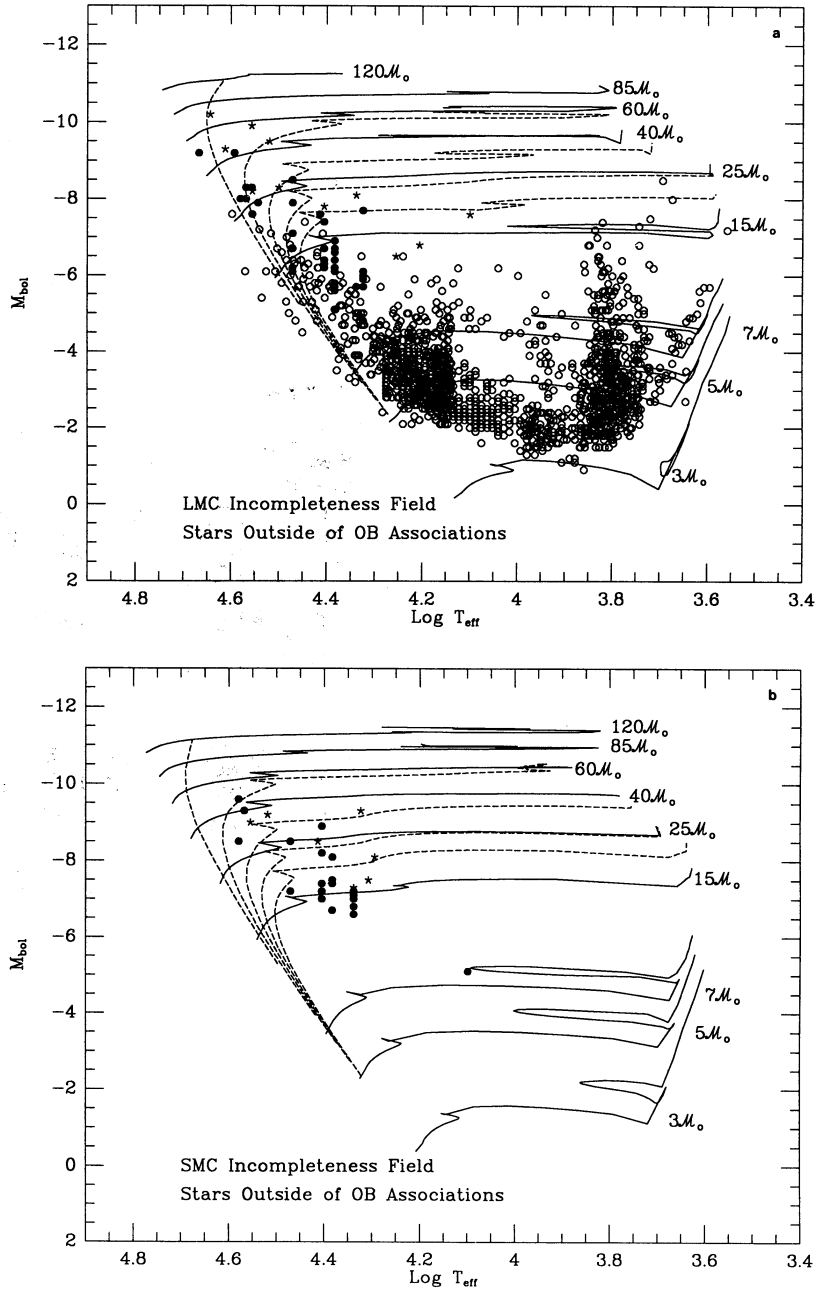

FIG. 6.—The HRDs for the (*a*) LMC incompleteness fields, and (*b*) SMC incompleteness field. Symbols are the same as in Figs. 4 and 5, with the addition that the asterisks denote the stars that were previously cataloged.

$T_{\rm eff} = 4.30$ and $\log g = 4.0$ (essentially a B2 V star) predict a change of 0.01 in $(B-V)_0$ in going from solar to 0.1 solar metallicity, and a corresponding change of 0.03 in $(U-B)_0$: both in the sense of the colors becoming redder toward lower metallicity. The resulting change in Q is only 0.02, in the correct direction but only 10% of what is needed to explain the discrepancy for the SMC. (We are indebted to E. Fitzpatrick for passing on the results of the Kurucz ATLAS code.) In § 2.5 we suggested that the SMC stars might have a color to spectral type calibration that was 0.05 mag redder in Q, somewhat greater than Kurucz's models would predict, but still far short of the 0.19 difference needed here.

The second problem alluded to above is the apparent overabundance of stars with masses of 15 $M_\odot$ and below at modest temperatures ($\log T_{\rm eff} = 3.9$–4.1) evident in Figures 4 and 5. Since the catalogs do not go faint enough to detect any main-sequence stars at this mass, we cannot demonstrate that this is a true overabundance, but we note that these stars are well past the first kink in the tracks. The distribution terminates abruptly around $\log T_{\rm eff} = 3.95$. This diagonal clump is the "blue supergiant ledge" discovered by Fitzpatrick & Garmany (1990), in their analysis of the HRD for the LMC, which was based upon the same data set, minus the new spectral information. (Their study concentrated on the post-main-sequence evolution of 10–40 $M_\odot$ "intermediate mass" stars, as they realized the completeness problem for earlier types.)

They interpret this ledge (the existence of which is quantitatively demonstrated in their Table 3) as evidence for the presence of He-burning, second generation B supergiants, on their way to becoming SN 1987A's. We cannot comment on the cause of this ledge directly, although we note that the objective-prism surveys upon which LMC member stars were found were specifically biased toward OB stars, and that this density enhancement roughly corresponds to the division between the late B and early A stars; the fact that this line is diagonal in the HRD may further suggest an observational bias, as the line is basically a line of constant V (see Fig. 1*a*). We are nevertheless surprised that the density enhancement is so strong, and we agree that this ledge probably has an astrophysical rather than observational explanation.

We conclude from this: (1) The models do an excellent job reproducing the width of the main sequence above a mass of 15 $M_\odot$. (2) Our data cannot address whether or not there is a lack of massive stars younger than 2 Myr, but our reddening data do not support the contention that the youngest massive stars are missing because they are still embedded. (3) There appears to be a problem in the location of the ZAMS for the less luminous and massive stars. The mismatch is evident in our study of the LMC incompleteness field, amounting to ≈ 0.075 in log $T_{\rm eff}$ by $M_{\rm bol} = -4$ (about 7 $M_\odot$). We shall see in § 3.3.3 that this mismatch is also apparent in the HRDs of the OB associations as well, and is about twice as great for the SMC association NGC 346. This is considerably greater than can be attributed to the changes in the $(B-V)_0$ to effective temperature calibration at different metallicities. We suggest that the intrinsic stellar properties are more similar in the Milky Way, LMC, and SMC that the evolutionary models suggest.

3.3. *The Observed Initial Mass Function*

Earlier studies by Lequeux (1979a, b), Lequeux et al. (1980), Vangioni-Flam et al. (1980), and Humphreys & McElroy (1984), among others, have attempted to derive IMFs for the massive star content of the Milky Way and Magellanic Clouds, based upon mixed association and field populations. Typically these studies used luminosity functions and/or spectroscopic data whose completeness limits were not well understood. Nevertheless, values of $\Gamma \approx -2$ were (fortuitously?) derived, similar to the values found by more detailed examination of Magellanic Cloud associations (Massey 1993). In this section we will derive the IMF slope for stars in the field of the LMC and SMC, and redetermine the slopes for the Magellanic Cloud OB associations. We will then compare this to both the field and association values determined for the Milky Way.

One might reasonably ask what it means to derive an "initial mass function" for stars that were not born together, and we recognize that some of our colleagues may even question if it makes sense to even speak of an IMF under these circumstances. The quantity we are about to measure is simply the average number of stars born per year per (unit logarithmic mass interval) per area, and how this quantity varies with mass: this is, strictly, the definition of the initial mass function (Tinsley 1980), after all. What we hope this means physically is the ensemble average of all of the multitude of small star formation events that have occurred within the LMC or SMC within the past 10 Myr. In comparing this to the more easily understood IMF of a single association or cluster, we make the tacit assumption that the time-averaged rate of star formation in the field has not changed during the last 10 Myr, and that star formation has been "steady state": i.e., the number of stellar births equals the number of stellar deaths for stars of a particular mass. We submit that as long as this assumption is correct, then the quantity we measure is physically meaningful.

In order to derive this quantity for this field population, we merely count the number of stars between pairs of evolutionary tracks (40 and 60 $M_\odot$, say), and then correct for the size of the mass bin (normalizing to 1 dex in mass), and divide by the average age of stars within each mass bin.

As in previous studies, we will adopt the nomenclature of Scalo (1986), where the $\xi(\log m)$ is the mass function in units of numbers of stars born per unit logarithmic (base ten) mass interval per unit area (kpc^2) per unit time (Myr). The slope of the initial mass function is then

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

For a power-law mass spectrum

$$f(m) = Am^\gamma \, ,$$

where the index $\gamma = \Gamma - 1$. In Tinsley's (1980) notation, $x = -\Gamma$; a Salpeter (1955) mass function has $\Gamma = -1.35$.

3.3.1. *Correcting for Incompleteness*

We have previously shown the HRDs for the LMC and SMC incompleteness fields (Fig. 6). The asterisks denote the stars which already were cataloged; the remainder of the points in this diagram are newly found. Within each mass bin, the ratio of the number of stars newly found, to the total number, will provide an empirical correction factor. We give the results of this calculation in terms of the "completeness" of the existing catalogs in Table 9.

The completeness factors in Table 9 suggest that the SMC catalog is actually more complete than the LMC catalog despite the greater distance of the SMC. However, the lack of any stars younger than 4 Myr in our SMC incompleteness sample, despite the fact that many 2–3 Myr-old stars are seen in the NGC 346 HRD discussed below, causes us to suspect that our grism search for candidate blue stars simply was not as effective as our photometric survey in the LMC. Therefore, we will consider what effect the stronger incompleteness correction based upon the LMC data has on the SMC stars when we derive the IMF in the next section.

TABLE 9

COMPLETENESS AS A FUNCTION OF MASS

Masses ($M_\odot$)	Number of Stars: Total	Number of Stars: Cataloged	Completeness
	LMC		
85–120......	0	0	?
60–85.......	1	1	100%?
40–60.......	5	3	60%
25–40.......	8	2	25%
15–25.......	22	3	14%
	SMC		
85–120......	0	0	?
60–85.......	0	0	?
40–60.......	2	0	?
25–40.......	6	4	67%
15–25.......	12	5	40%

TABLE 10

DERIVATION OF INITIAL MASS FUNCTION FIELD STARS

Masses ($M_{\odot}$)	N	log ξ Uncorrected	log ξ Corrected
LMC Field Stars			
85–120......	2	−0.75	−0.75
60–85.......	9	−0.10	−0.10
40–60.......	43	0.39	0.61
25–40.......	164	0.67	1.27
15–25.......	320	0.69	1.55
SMC Field Stars			
85–120......	1	−0.21	−0.21
60–85.......	2	0.00	0.00
40–60.......	10	0.59	0.82
25–40.......	47	0.97	1.57
15–25.......	76	0.89	1.74

3.3.2. *The IMF of the Field Stars*

We can now derive an IMF for the field stars, using the data shown in Figures 4*b* and 5*b*. We begin by counting stars between mass tracks, and then correct for both the size of the mass bin and for the average lifetime of a star within a bin. (For the lifetimes we use the ages in Table 8). In Table 10 we give the raw number of stars within these bins, as well as the computed values of log ξ. We have adopted 25 kpc^2 and 3.4 kpc^2 as the projected surface areas of the LMC and the SMC based upon the fields originally surveyed (i.e., 40 $\deg^2$ and 3.1 $\deg^2$).

We show in Figure 7 the run of log ξ with mass: the slope of this is the slope of the initial mass function, Γ. In fitting the data we have not used any weighting; we have included only the three highest mass points in the fit to the uncorrected data, but the four highest mass points in the fit to data corrected for incompleteness. It is clear that correcting the data does bring the lower mass points more into line with the higher mass points, as we expect if the mass function is a power-law. Fitting the uncorrected data yields a slightly flatter slope, as expected given our contention that the third mass bin (40–60 $M_{\odot}$) is only 60% complete.

The slopes derived in Table 11 are quite steep ($\Gamma \approx -3$ to -4) compared to what has been derived for OB associations in the Magellanic Clouds and the Galaxy, which typically have $\Gamma = -1.0$ to -1.8 (Massey 1993). In the following section we will rederive the IMFs of the Magellanic Cloud associations using the same models as used on field data to ascertain that our comparison is valid, and in § 3.3.4 compare the field result with that of the OB associations of the Magellanic Clouds and the Galaxy.

3.3.3. *OB Associations Revisited*

In previous papers we have investigated star formation in specific OB associations in the Clouds: LH 117 and LH 118 (Massey et al. 1989a), LH 9 and LH 10 (Parker et al. 1992), and LH 58 (Garmany, Massey, & Parker 1994) in the LMC, and NGC 346 in the SMC (Massey et al. 1989b). In these studies, we found IMF slopes that appeared to be somewhat steeper than those in the Milky Way (Massey 1993), although as much variation was found within the LMC as between the Milky Way and the Clouds.

In those analyses, we used the Galactic metallicity ($z = 0.02$) tracks of Maeder & Meynet (1988) simply because, when we began our studies, nothing of comparable quality was available at more appropriate metallicities; in subsequent studies we were reluctant to switch to the newer models as differences between the models might mimic or mask real differences between regions. However, since we have now used our field star data to test the isochrones of these newer models, and derive IMFs, it is important that we be able to make valid comparison to the OB association data. We have therefore chosen to redetermine the HRDs of our associations employing our current transformations, slightly different distances, and the evolutionary tracks of Schaller et al. (1992) and Schaerer et al. (1993).

We present the revised H-R diagrams in Figures 8–12. For each association we computed the locations in the HRD both for the case of variable reddening, in which we compute $(B-V)_0$ from Q (as given in Table 7), and for the case where we fixed the reddening to the average value found in the association. The latter includes more stars but it does so by keeping stars whose colors are inconsistent with the average reddening. (In the above cited published work we sometimes used one method and sometimes the other.) Note that in many cases the assumption of constant reddening includes stars whose location in the two-color plane places them to the left of the ZAMS; these stars are not, however, unusually heavily reddened objects, and their location is simply due to the asymmetrical results that a Gaussian error distribution has in the colors of stars (Massey & Johnson 1993).

We have referred above to the mismatch between the evolutionary models and the location of stars for "lower" masses ($<15\ M_{\odot}$), and we see this effect in all of the LMC associations here. Although the massive stars cluster between the 2–4 Myr isochrones, by lower luminosities the stars are located at considerably cooler temperatures than the evolutionary tracks would have. We have previously noted that we do not see this problem in our Galactic associations, and believe that this is an artifact of the lower metallicity models. Note that in the SMC association NGC 346 (Fig. 8) that the problem is considerably more pronounced. When the stars in NGC 346 are plotted using the Galactic metallicity models of Maeder & Meylan (1988) this problem is not seen (Massey et al. 1989b). This suggests again that the ZAMS in the lowerer metallicity models are too far to the left in the HRD.[6]

We can use these data to now construct new mass functions for these OB associations. We show these mass functions in Figure 13. We can see that there is little difference in the results depending upon how we treat reddenings. The mass function slopes have been fitted to the points from log $M/M_{\odot} = 1.01$ (corresponding to the 7–15 $M_{\odot}$ bin) and higher; i.e., the points from the 5–7 $M_{\odot}$ bin are shown on the diagram but are not

TABLE 11

MASS FUNCTION SLOPES MC FIELD STARS

	Uncorrected	Corrected
LMC......	$\Gamma = -3.7 \pm 0.4$	$\Gamma = -4.1 \pm 0.2$
SMC......	$\Gamma = -2.6 \pm 0.6$	$\Gamma = -3.7 \pm 0.5$

[6] We also note in passing that Kudritzki et al. (1989) have also analyzed NGC 346, finding a considerably higher mass ($>100\ M_{\odot}$) star than the highest shown in Fig. 12. This discrepancy is due to their adoption of an earlier spectral type (O4 III) for the most luminous star; the photometry and spectroscopy of Massey et al. (1989b) found a later type (O5.5 If), with a correspondingly smaller bolometric correction; in addition, Massey et al.'s imaging found that the object is actually multiple.

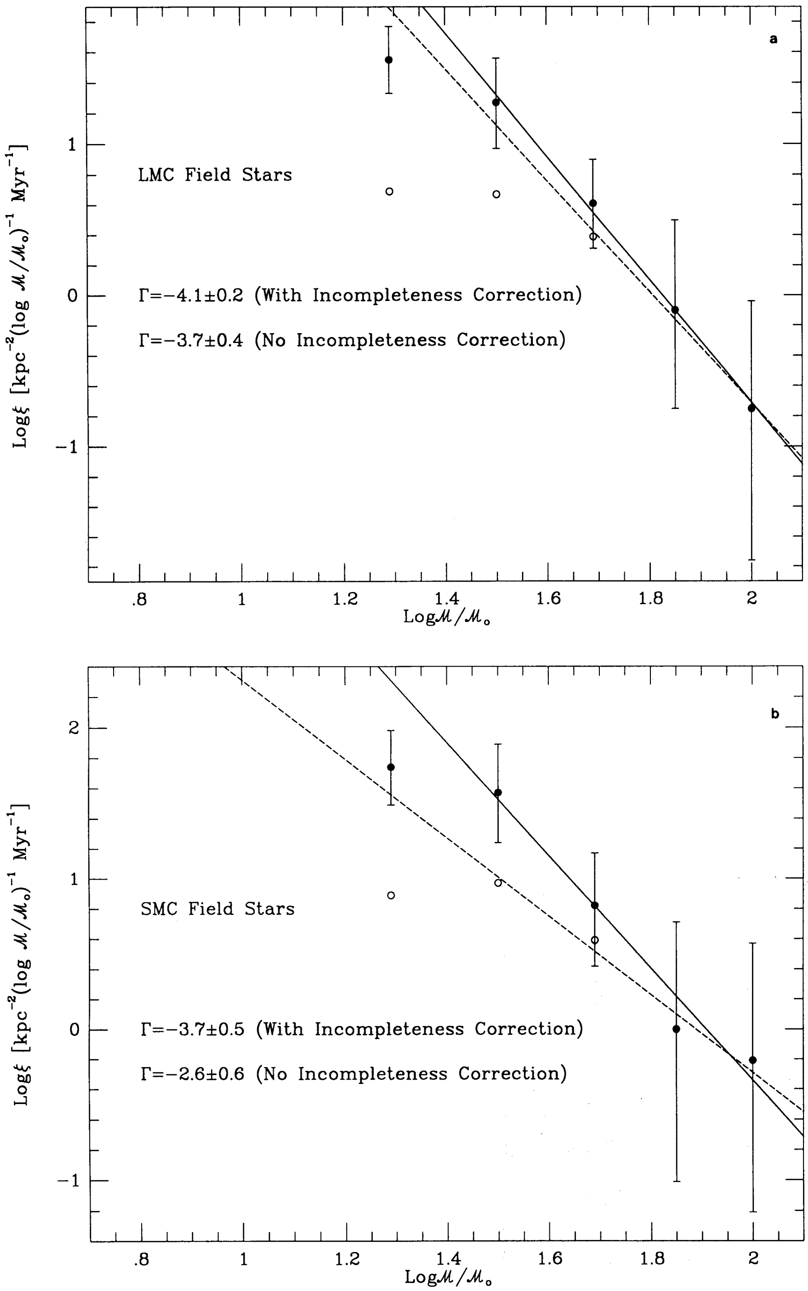

FIG. 7.—The initial mass function for the field stars of (*a*) the LMC and (*b*) the SMC. Solid lines denote the fits to the four most massive points for the data with the incompleteness correction included; dashed lines are the fits to the three highest mass points for the uncorrected data.

included in the fit because of incompleteness. The fits have been made weighting inversely by the errors, assumed to be $1/N^{1/2}$.

These data have not been corrected for the ages of individual stars within an association; instead, we make the assumption that most stars within each OB association are coeval. This assumption has been discussed extensively in the papers cited above; we can see from Figures 8–12 that in general this is a good approximation for the more massive stars, although there are invariably a few evolved supergiants of lower ($\approx 15\ M_\odot$) mass that must be considerably older.

We give the IMF slopes for the Magellanic Cloud OB associations in Table 12. These slopes are marginally less steep than the published values, with the most significant change for the SMC associations, as one might expect. Typically the IMF slopes are found to have $\Gamma = -1.4$, essentially that of Salpeter (1955).

In a recent series of papers, Hill, Madore, & Freedman (1994a, b, c) attempt to determine the IMF of stars in OB associations of the LMC using CCD photometry. They find a considerably steeper slope, $\Gamma = -2.0 \pm 0.5$ (Hill et al. 1994c). Of these, we overlap with only one OB association, LH 58, for

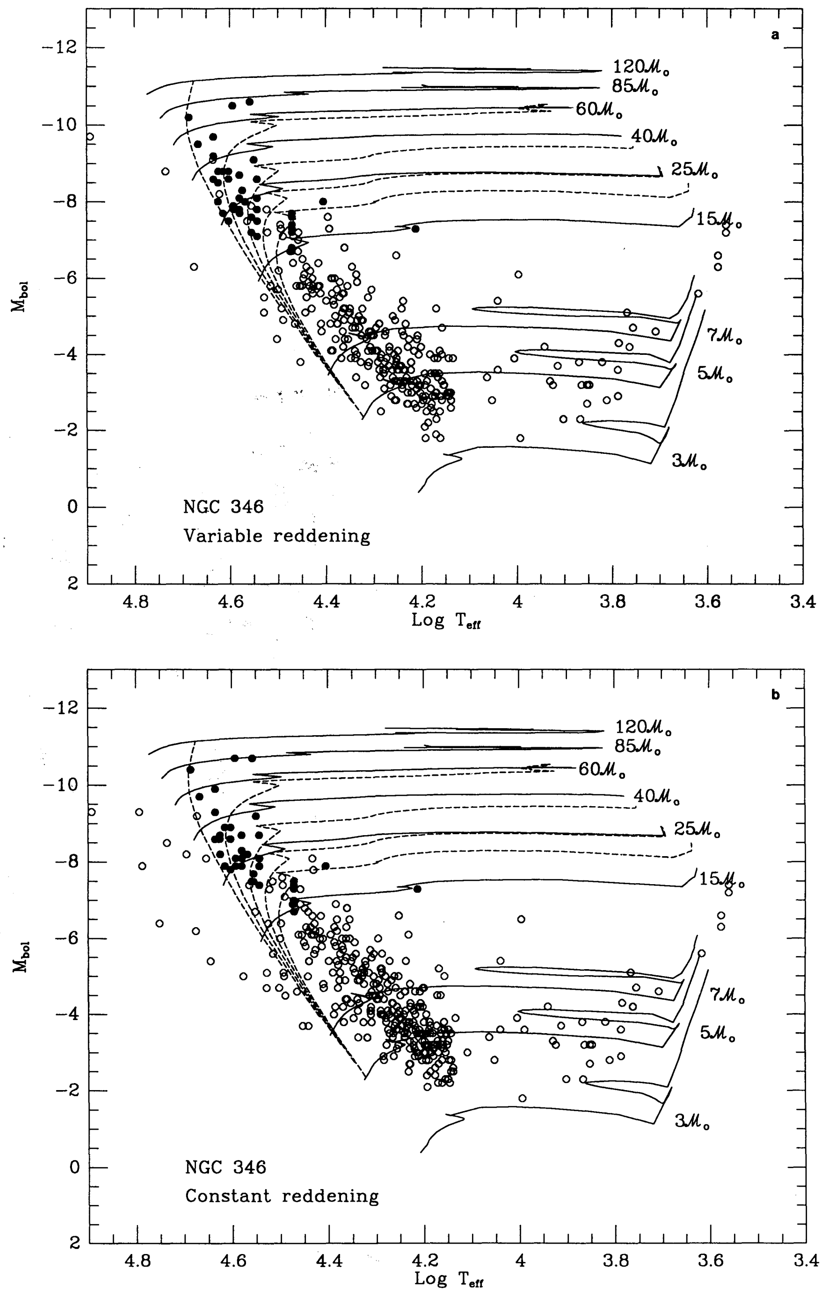

FIG. 8.—The HRD for NGC 346 in the SMC. Data are from Massey et al. (1989b). The filled circles are stars with spectroscopy, while the open circles have been placed in the diagram on the basis of photometry only. The evolutionary tracks (*solid lines*) are the $z = 0.001$ models of Schaller et al. (1992); dashed lines are isochrones at 2 Myr intervals. (*a*) The "variable reddening" case was considered correcting for reddening as in Table 7. (*b*) The average $E(B-V)$ of Massey et al. (1989b) was adopted.

which they find $\Gamma = -2.5 \pm 0.3$, compared to our value of $\Gamma = -1.4 \pm 0.2$—a fairly significant difference! Where does this come from?

We believe that it is simply in the inability of photometry to treat the hottest (and most massive) stars. The Hill et al. studies lacked any spectroscopy, but relied purely upon their photometry for placing stars in the HRD. In order to test this, we redetermined the IMF for LH 58 using our data but ignoring our spectral types—instead, the stars were placed in the HRD using only the photometry. For good measure, we adopted the transformation equations of Hill et al. (1994c, Tables 4 and 5). We show this H-R diagram in Figure 12*c*. By comparison with Figure 12*a*–12*b*, we see just as we might expect: there are far fewer stars populating the upper main sequence. We formally derive an IMF slope of $\Gamma = -2.0 \pm 0.2$, compared to our value (based upon photometry *and* spectroscopy) of $\Gamma = -1.4 \pm 0.2$. This is in accord with our previous experiment on NGC 346: when Massey (1989b) recomputed the IMF

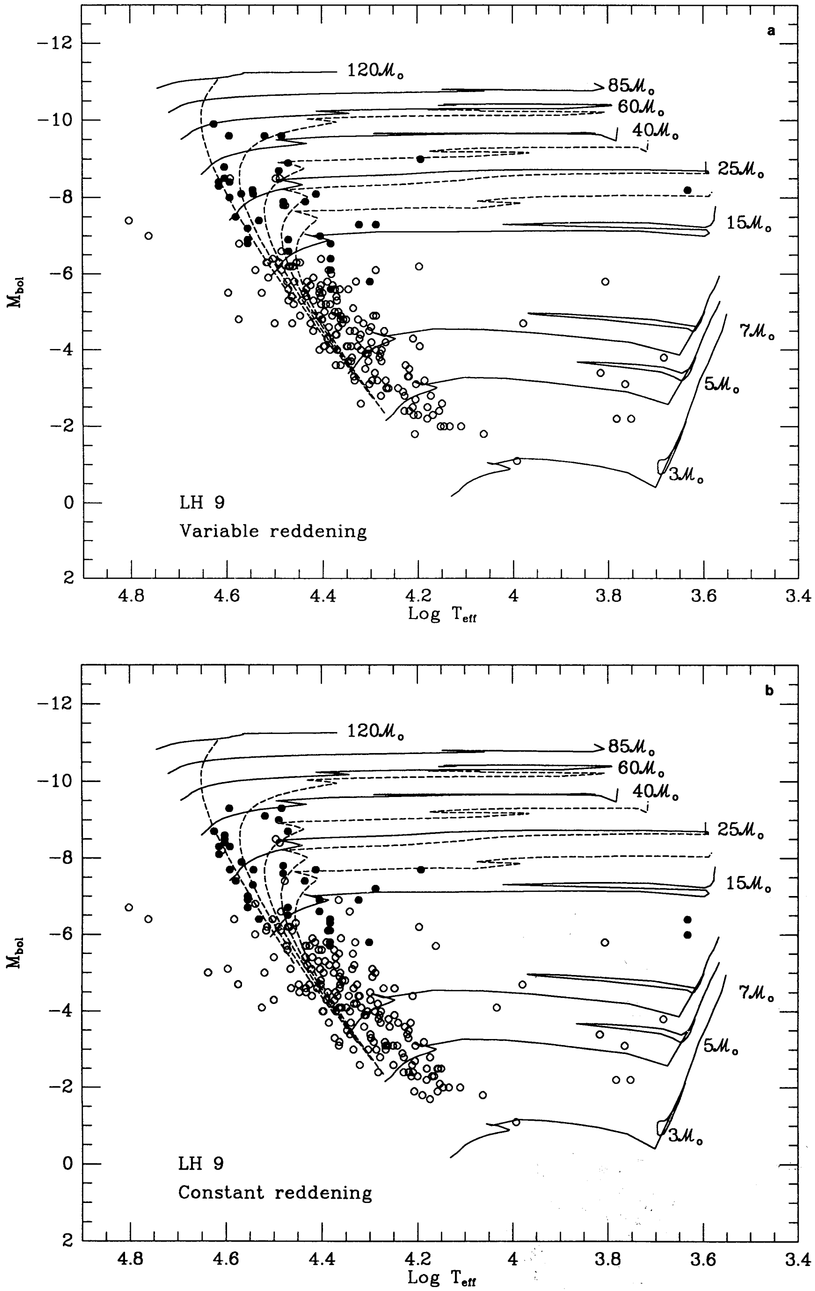

FIG. 9.—Same as Fig. 8, for Lucke-Hodge 9 in the LMC. Data are from Parker et al. (1992), and tracks are the $z = 0.008$ models of Schaerer (1993).

TABLE 12

MASS FUNCTION SLOPES LMC/SMC OB ASSOCIATIONS

	New			
OB Associations	Variable Reddening	Constant Reddening	OLD	REFERENCE
		LMC		
LH 9	$\Gamma = -1.4 \pm 0.2$	$\Gamma = -1.5 \pm 0.2$	$\Gamma = -1.6 \pm 0.1$	Parker et al. 1992
LH 10	$\Gamma = -1.1 \pm 0.1$	$\Gamma = -1.1 \pm 0.1$	$\Gamma = -1.1 \pm 0.1$	Parker et al. 1992
LH 58	$\Gamma = -1.4 \pm 0.2$	$\Gamma = -1.5 \pm 0.2$	$\Gamma = -1.7 \pm 0.3$	Garmany et al. 1994
LH 117/118......	$\Gamma = -1.6 \pm 0.2$	$\Gamma = -1.7 \pm 0.2$	$\Gamma = -1.8 \pm 0.1$	Massey et al. 1989a
		SMC		
NGC 346	$\Gamma = -1.3 \pm 0.1$	$\Gamma = -1.4 \pm 0.1$	$\Gamma = -1.8 \pm 0.2$	Massey et al. 1989b

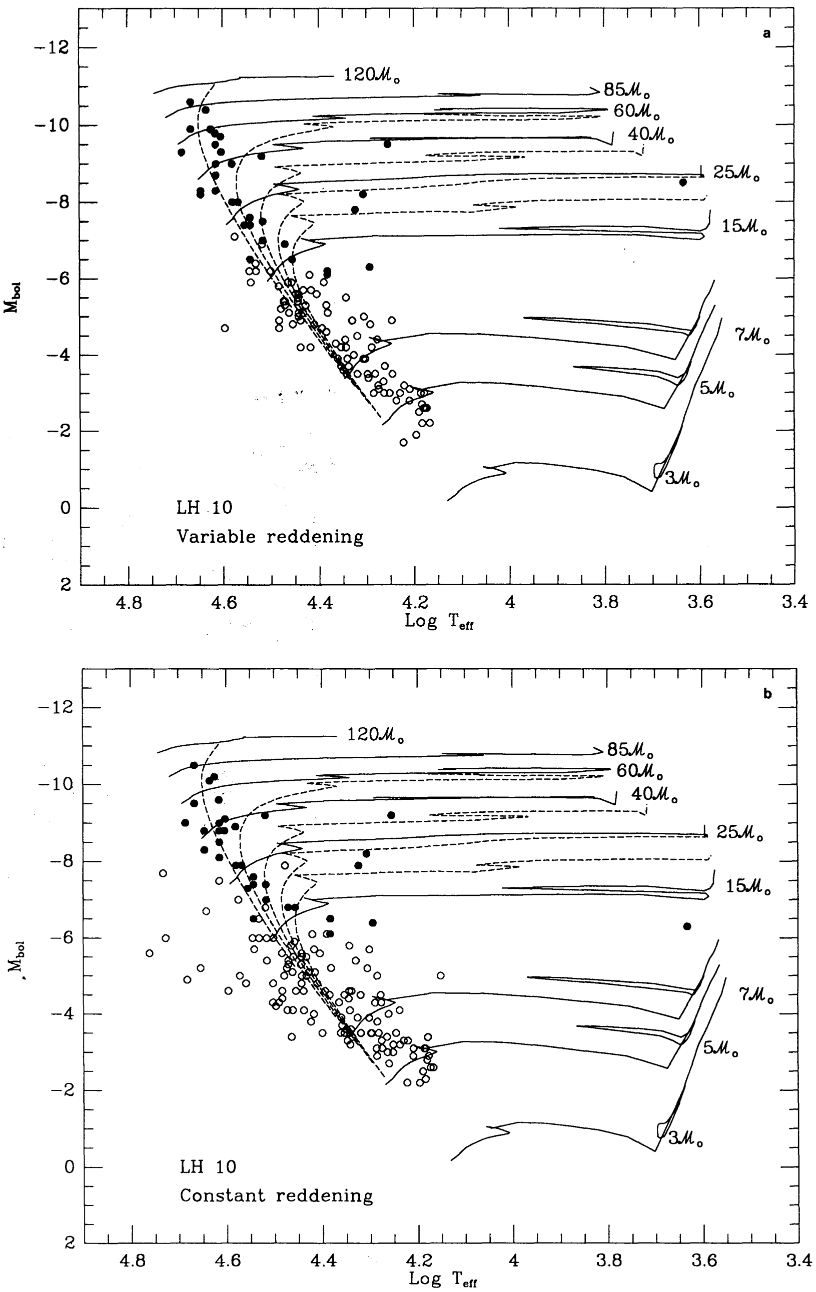

FIG. 10.—Same as Fig. 9, for Lucke-Hodge 10

slope based upon photometry alone, the resulting slope steepened to -2.5 rather than the -1.8 found by the use of spectroscopy to locate the most massive stars.

3.3.4. *The IMFs of the Fields, Associations, and the Galaxy Compared*

The data given above strongly suggest that the IMF of the field stars in the Magellanic Clouds is considerably steeper than that of the OB associations in the Magellanic Clouds. We find that $\Gamma \approx -4$ for the field stars in both the LMC and the SMC, while the OB associations have typically $\Gamma = -1.0$ to $\Gamma = -1.6$ (Table 12).

How do these slopes compare to the field and associations of the Milky Way? We are currently engaged in studying a number of very young clusters and associations within the Milky Way to determine the IMF slope and its variation, as well as ages and age spreads. Three regions have been completed, and the IMF slopes are shown in Table 13. *We conclude from this then that there are no significant systematic differences in the IMF slopes found for the OB stars found in the Milky Way and the LMC or SMC.* This differs from the conclusions of Massey (1993), who suggest that the MCs have somewhat steeper slopes; this latter was based upon the $z = 0.02$ evolutionary models. Use of the lower metallicity models for the Magellanic Clouds has brought the IMF slopes into closer agreement with that of the Milky Way.

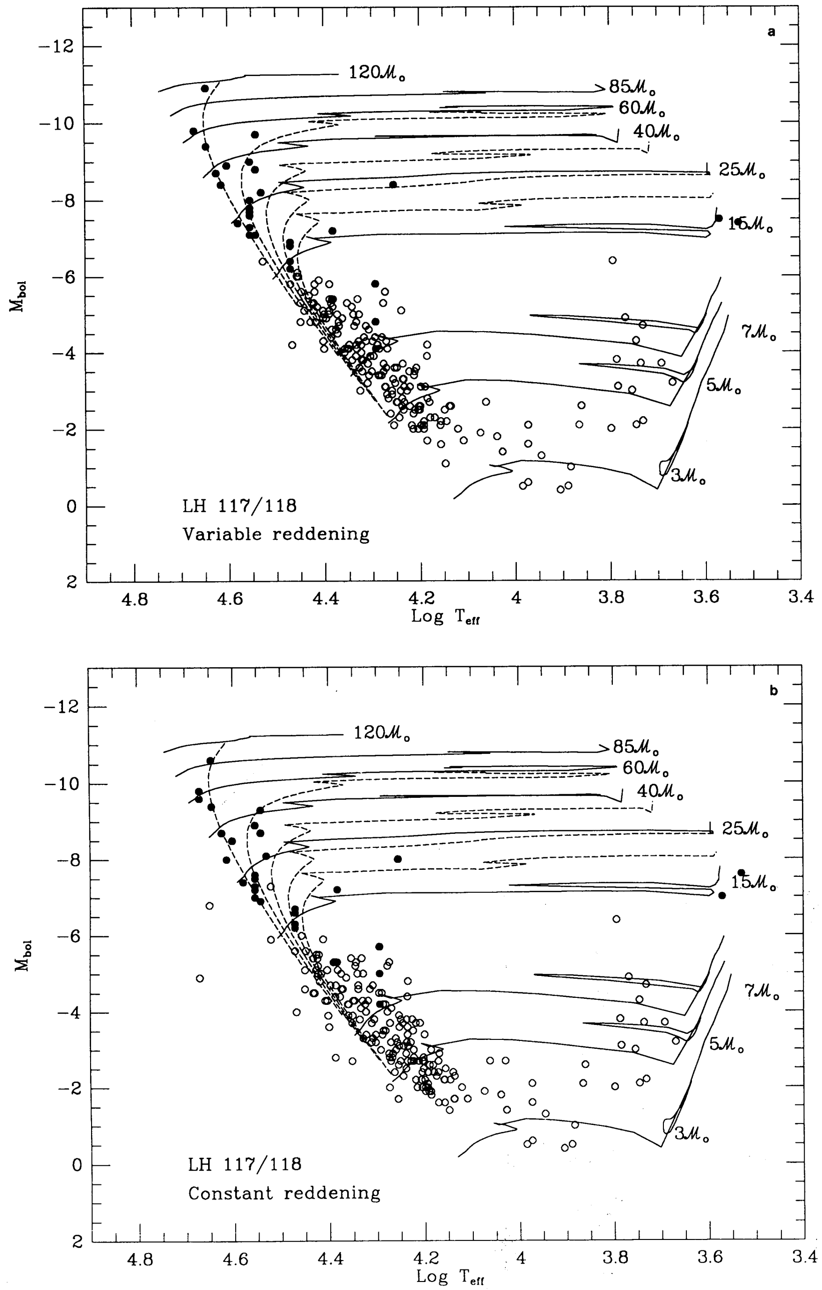

FIG. 11.—Same as Fig. 9, for Lucke-Hodge 117 and 118. Data are from Massey et al. (1989a).

What about the IMF of the Galactic field stars? Garmany, Conti, & Chiosi (1982) have determined the IMF of a volume-limited sample of massive stars within a few kiloparsecs of the Sun, and indeed it was this work that precipitated the present study. However, that analysis used a mixed sample of both field and association stars. We have made a preliminary reanalysis of their data set to see if the field stars in the Galaxy also show this much steeper slope. The data set used by Garmany et al. contain 424 stars within 2.5 kpc, of which 187 (44%) are listed as field stars. The original calculation was done using earlier evolutionary tracks by Bressan, Bertelli & Chiosi (1981). We have reanalyzed these data with the Galactic metallicity evolutionary tracks by Maeder & Meynet (1988) and computed the mass function for stars of mass 25 $M_\odot$ and greater, for the entire sample and for the field stars alone. The slope of the mass function for the entire data set is $\Gamma = -1.45 \pm 0.2$ much like the one derived by Garmany et al. ($\Gamma = -1.6$). However, when we consider the field stars alone we find $\Gamma = -3.2 \pm 1.4$. Although the Galactic field star data

TABLE 13

MASS FUNCTION SLOPES GALACTIC OB ASSOCIATIONS

OB Association	Γ	Reference
Cyg OB2	-1.0 ± 0.1	Massey & Thompson 1991
Tr 14/16	-1.3 ± 0.2	Massey & Johnson 1993
NGC 6611......	-1.1 ± 0.3	Hillenbrand et al. 1993

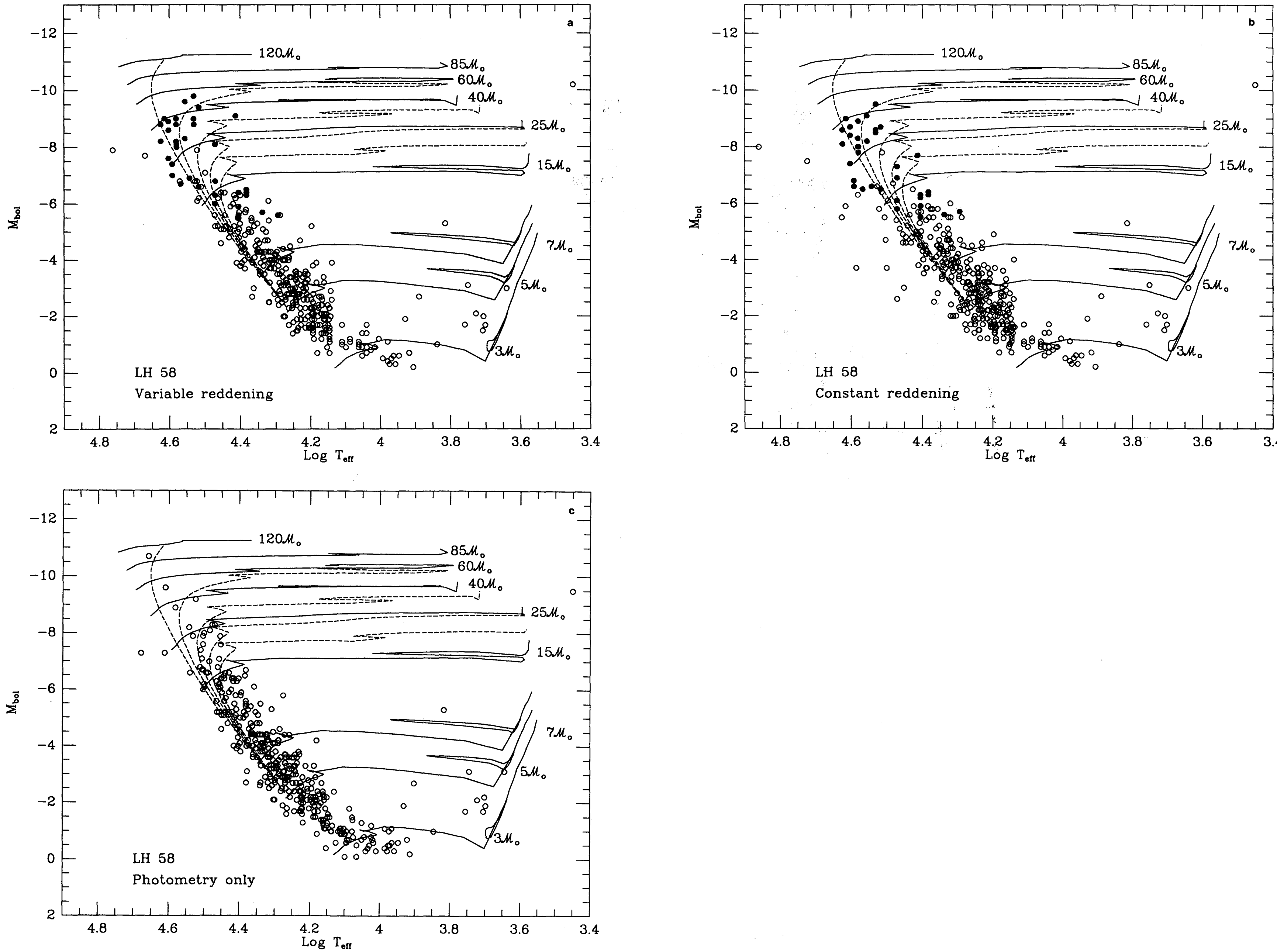

FIG. 12.—Same as Fig. 9, for Lucke-Hodge 58. Data are from Garmany et al. (1994). The additional plot, (*c*), shows the placement of stars in the HRD if the spectroscopy is ignored and the transformations to log T_{eff} and M_{bol} were made purely on the basis of photometry. For good measure, we have adopted the transformation equations of Hill et al. (1994c). Comparison with (*a*) and (*b*) shows the problems of attempting to derive the HRD without the benefit of spectroscopy.

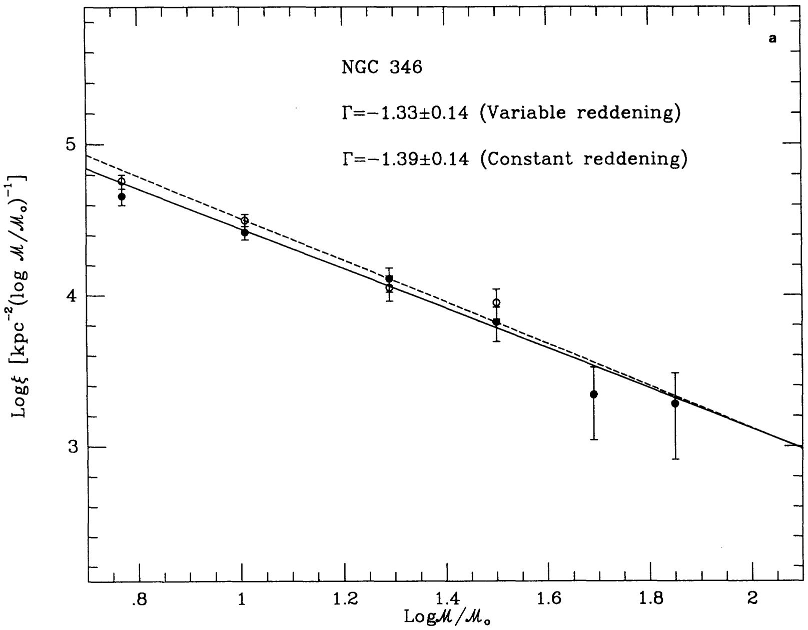

FIG. 13.—The IMFs for the OB associations of Figs. 8–12. The solid curve shows the fit to the variable reddening (*solid points*) calculations, and the dashed curve shows the fit to the constant reddening (*open circles*) calculations. In all cases the lowest mass point was not used in computing the fit.

should be treated cautiously as we are unsure of the selection effects in this sample, the results are in accord with what we see here for the Magellanic Clouds.

3.4. *Stellar Evolution and the Relative Number of Main-Sequence and WR Stars Compared in the LMC and SMC*

We have answered the three main questions posed at the beginning of this paper, but there is a fourth area that we wish to address: how does massive star evolution proceed past the main sequence? In particular, how does the formation of Wolf-Rayet stars depend upon metallicity? Before making our modest contribution in this area, we first need to place it in context.

The most massive stars arrive on (or near) the ZAMS as O-type stars, and evolve to B-type supergiants during their H-burning lifetimes. Our analysis of the HRDs presented above clearly indicates that the evolutionary models are consistent with the observations in this region, at least. The details of subsequent evolution, however, are murky, due to the extreme role that mass loss plays. It is now believed that above some mass ($\approx 40\ M_\odot$) stars will become Luminous Blue Variables (LBVs) for a short ($\approx 10^4$ yr) period, undergoing episodic mass loss. These very massive stars will then evolve to the left in the HRD, and become spectroscopically identified as Wolf-Rayet stars. Here a star will spend a majority of its He-burning life ($\approx 10^5$ yr). (A recent review may be found in Maeder & Conti 1994). For stars of slightly less mass (10–15 $M_\odot$), evolution proceeds from the main sequence to a red supergiant phase instead, and may then turn blueward spending its last (nuclearly) productive years as a "second generation" blue supergiant before becoming a "puny" supernova à la SN 1987A.

The mass above which a star becomes a WR star rather than a RSG should depend upon mass-loss rate, which in turn should depend upon the initial metallicity of the star, as first noted explicitly by Maeder, Lequeux, & Azzopardi (1980). Recent evolutionary modeling suggests that the mass needed to become a WR star may be 60–85 $M_\odot$ if z is SMC-like, but 25–30 $M_\odot$ if z is like that of the Milky Way (Maeder & Meynet 1994).

Needless to say, observational checks of this are difficult to come by. The effective temperatures and bolometric corrections for WR stars are extremely uncertain (Conti 1988), making their placement in the HRD quite speculative. (See, for example, discussion in Massey & Johnson 1993 concerning the location in the HRD of a WR star in the η Car region.) Humphreys, Nichols, & Massey (1985) studied a number of Galactic associations containing WR stars and attempted to answer this question by determining the location of the highest mass star that was still core H-burning. This study did not pressume to address the issue of metallicity dependence, but did suggest that stars above $\approx 40\ M_\odot$ became WR stars while most RSGs must come from lower mass stars. Their study also demonstrated that RSGs and WR stars were anticorrelated in the OB associations of M33 (metallicity similar to that of the LMC), as would be expected if they come from different mass progenitors.

The evidence for metallicity effects is unfortunately convolved with the possibility of IMF/SFR differences, and, we will note shortly, the difficulty of studying these matters with mixed association and field statistics. We *do* know with some certainty that the number of WR stars per unit area projected on the sky is a factor of 2 to 3 lower in the SMC than in either the LMC or the solar neighborhood. We also know that there is a strong change in the surface density of Wolf-Rayet stars as one moves inward to outward of the solar circle. Garmany, Conti, & Chiosi (1982) attributed this last result to the galactocentric gradient they though they had detected in the slope of the IMF.

However, in both our present study of the field and in previous work on associations we have demonstrated that incompleteness is significant for the H-burning massive star populations even in the Magellanic Clouds. On the one hand,

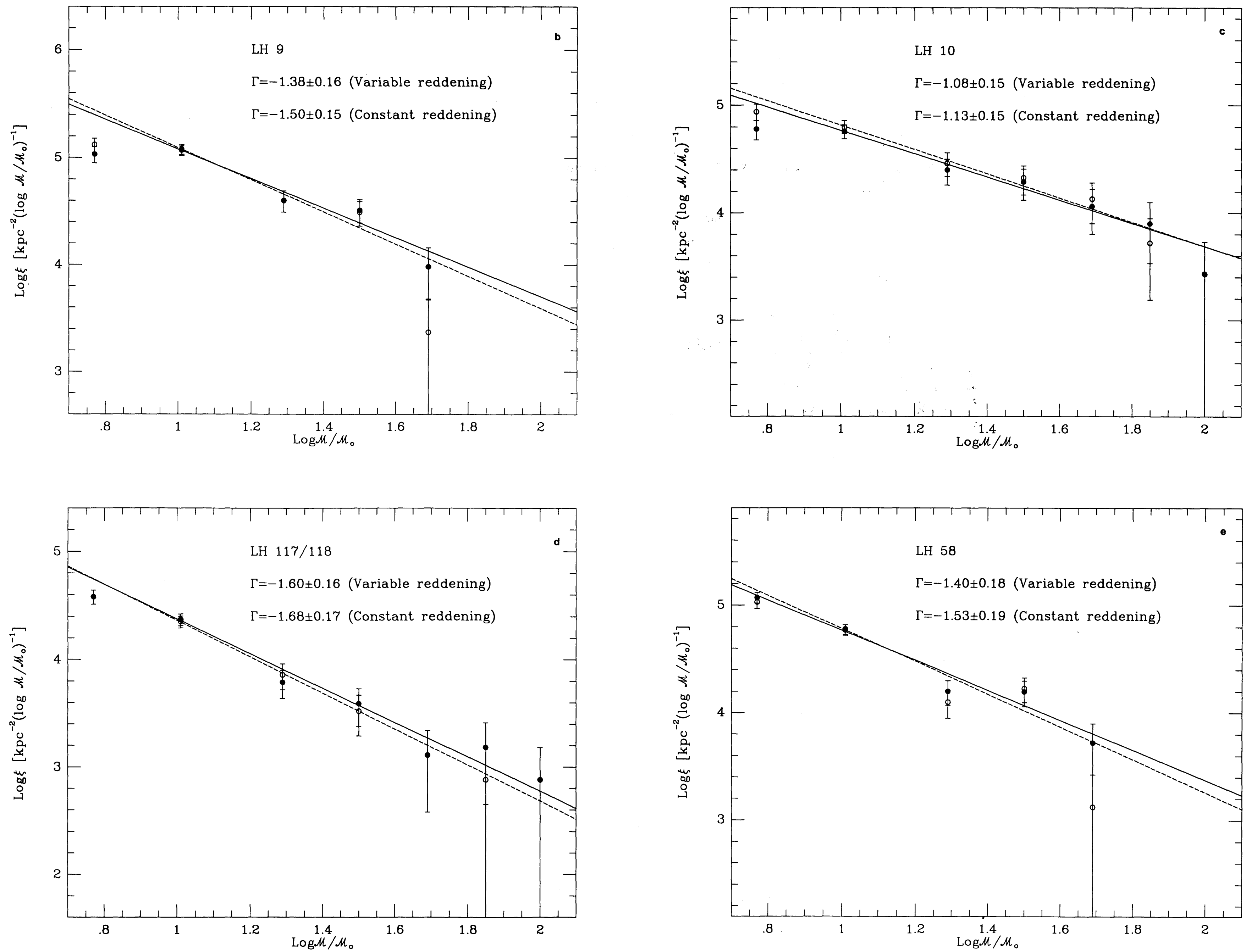

FIG. 13—*Continued*

our knowledge of the Wolf-Rayet population is probably complete in the LMC, SMC, and within a few kilopersec of the Sun, despite the fact that most of these stars are found in OB associations and/or clusters. (An unpublished survey of many SMC associations some years ago by the first author using the CITO 4 m and interference filters failed to reveal even a single new candidate WR star.) However, although most of the WR stars are found in associations and clusters (at least in the Magellanic Clouds), most of the O-type stars known are in the uncrowded field regions. This last fact is implicit in our work on the Magellanic Cloud OB associations: for example, Massey et al. (1989b) identified over 30 O-type stars in NGC 346 in the SMC, *of which only* one *appears in the Azzopardi & Vigneau* (*1982*) *catalog.* The situation is similar for the LMC, where only one of the O-type stars found by Massey et al. (1989a) had been previously identified by Sanduleak (1969). This is hardly surprising given the difficulty that objective-prism surveys have in crowded regions that include nebulosity. However, we find that this is a problem even within the "volume-limited" sample of O-type stars cataloged by Garmany, Conti, & Chiosi (1982): although they state completeness out to 2.5 kpc, two of the three Galactic OB associations we have previously studied fall within this sphere: Cyg OB2, at 1.7 kpc (Massey & Thompson 1991) and NGC 6611, at 2.0 kpc (Hillenbrand et al. 1993). The completeness for O-type stars in these two clusters can be inferred as $\approx 50\%$ from the relevant tables in these papers.

In any event, when making comparisons to the number of WR stars, it is not the number of O-type stars which are of interest but rather the total number of main-sequence stars (of all type) above a certain mass. Conti et al. (1983) used the number of O-type stars as equivalent to the number of main-sequence stars, but their results simply confirmed that the theoretical main-sequence was considerably wider than the region occupied by O-type stars: the ratio of WR stars to O stars found was $\frac{1}{4}$, while one expects that the relative numbers of WR to main-sequence stars will simply be that of the relative He- and H-burning lifetimes, which first principles suggest should be $\frac{1}{10}$.

Using our Magellanic Cloud data we can instead use the number of main-sequence stars (not just the O-type stars) as a function of mass to compare with the WR population in a relatively unbiased method. We assure that the main-sequence and WR samples are similarly complete by selecting only those stars (including WR stars) that meet our criterion of being field stars. We find that there are 38 field WR stars in the LMC, out of a total population ≈ 100 (Breysacher 1981; Conti & Garmany 1983; Azzopardi & Breysacher 1985). (Fourteen of these would also meet our definition of being an "extreme" field star.) For the SMC there are only eight WR stars (Azzopardi & Breysacher 1979), and of these only one meets our definition of being a field star.

If examines such a mixed-age ensemble, one expects that the number of WR stars relative to the number of unevolved stars of similar (initial) mass will simply scale as their relative ages, which is, to a good approximation, just the relative lengths of the He- and H-burning lifetimes. Even with the modern evolutionary models considered here, this ratio is expected to be approximately 0.1. Schaerer et al. (1993) give H-burning lifetimes of 2.8–4.8 Myr for 120–40 $M_\odot$ stars, while the WR lifetimes are 0.3–0.4 Myr for $z = 0.008$. The models of Schaller et al. (1992) predict only the absolutely most massive ($>85\ M_\odot$) stars become WR stars, but even these have ages ≈ 0.3 Myr, while the main-sequence lifetimes are 3.1–4.9 Myr (85–40 $M_\odot$).

TABLE 14

RATIO OF FIELD WR STARS TO FIELD MS STARS

MS Stars	Number	Ratio WR/MS
LMC (Number of WR Stars = 38):		
$>85\ M_\odot$	2	19
$>60\ M_\odot$	11	3.5
$>40\ M_\odot$	83	0.45
$>25\ M_\odot$	740	0.05
SMC (Number of WR Stars = 1):		
$>85\ M_\odot$	1	1
$>60\ M_\odot$	3	0.3
$>40\ M_\odot$	20	0.05
$>25\ M_\odot$	210	0.005

In Table 14 we give the ratio of the number of main-sequence stars above a certain mass limit compared to the total number of WR stars. The mass at which this ratio is roughly 0.1 is the low mass limit for progenitor WR stars.

Using this, we indeed find much as we expect that stars more massive than, say, 30 $M_\odot$ become WR stars in the LMC. The statistics are of course far less certain for the SMC, given that there is only one field WR star, but suggest that the cut-off might be a bit higher for the SMC: perhaps 50. For the $z = 0.001$ models Schaller et al. (1992) find that only stars above 85 $M_\odot$ spend a significant time as a WR star; if that were correct, we would be surprisingly lucky to see the one that we do.

4. SUMMARY

We have examined the massive star content of the Magellanic Clouds, separating, we believe, massive stars that were truly born in the "field," i.e., stars not born as part of large OB complexes. We have obtained spectra of the bluest stars in order to make accurate HRDs, and corrected for incompleteness reliably down to 25 $M_\odot$. We find the following:

1. Stars as massive as those found in OB associations are produced in the field.

2. The initial mass function of the field stars is much steeper ($\Gamma \approx -4$) than that of the associations and clusters we have studied in the Magellanic Clouds and Milky Way: far fewer of the most massive stars are produced compared to the number of lower mass stars. We confirm that this is true for the Milky Way as well.

3. The slopes of the initial mass function of the Magellanic Cloud OB associations are found to be $\Gamma = -1.3 \pm 0.3$, considerably flatter than that of the field. We do not find any significant differences between the Milky Way, LMC, and SMC.

4. The current Geneva evolutionary models computed for metallicities appropriate to the Magellanic Clouds do an excellent job of reproducing the distribution of stars in the HRD for the higher mass ($>15\ M_\odot$) stars. There is an increasing discrepancy between the theoretical ZAMS and the observed blue edge of the main sequence, amounting to 0.075 dex in effective temperature by $M_{\rm bol} = -4$ in the LMC, and 0.15 dex for the SMC. This discrepancy suggests that there may be a problem in the location of the theoretical ZAMS. Alternatively, there may be a larger metallicity dependence in the intrinsic colors of stars of the same effective temperature than the stellar atmosphere models allow (Kurucz 1979), amounting to 0.19 in Q for the SMC around spectral type B2, and increasing toward lower temperature.

5. The number of field Wolf-Rayet stars to field main-sequence stars suggests that stars more massive than 30 $M_\odot$ evolve to WR stars in the LMC. The limit may be a bit higher in the SMC, possibly 50 $M_\odot$, but is unlikely to be as high as $z = 0.001$ models predict (85 $M_\odot$). We suggest that previous studies of this number ratio have been in error by as much as a factor of 2: the number of O stars in OB associations and clusters is underrepresented compared to the number of WR stars.

This long-term project was made possible only through the generous, long-term allocation of telescope time at CTIO. We gratefully acknowledge the always excellent support we have received on Cerro Tololo over the years. Useful conversations with Bruce Bohannan, Peter Conti, Ed Fitzpatrick, Lynne Hillenbrand, Deidre Hunter, and Stephen Strom have all helped shaped the course of this work. In addition, Conti, David De Young, Jay Gallagher, Hillenbrand, Hunter, Eric Wilcots, and Sidney Wolff were kind enough to provide useful comments for improving this manuscript. This research made use of the Simbad database, operated at CDS, Strasbourg, France. P. M. also wishes to thank Barry Lutz for the warm hospitality he enjoyed while on sabbatical at NAU, where much of this work was completed.

APPENDIX

In this appendix we will illustrate the most interesting of the spectra obtained in this program, including the earliest type stars, and the Be stars with Fe II emission.

We show in Figure 14*a* the spectra of stars of type O3 and O4 newly observed in the LMC and SMC (Tables 1–4). Balmer and He II lines dominate. The dividing line between O3 and O4 comes from the presence or absence of He I; in principle, this depends upon the quality of the spectra, although in practice we have found good agreement between photographic, moderate-dispersion spectra and higher signal-to-noise CCD spectra of the same objects.

The star lmc2-675 is composite; we see the strong N V λ4603, 19 absorption signature characteristic of an O3 If star; strong N IV λ4058 emission is also consistent with an O3 If type (Walborn & Fitzpatrick 1990.) However, the presence of He I λ4471 is unmistakable, and there is probable He I λ4387 absorption. We believe this is a spectroscopic binary, and as such is a potential find for followup radial velocity study, which would allow the first direct determination of the mass of an O3 star, were the system to prove to be eclipsing. The only other star in our sample that showed definite signs of being a binary is lmc1-104, an O9.5 V star whose lines appear to be double (Table 3).

We earlier stated that the seemingly phenomenal number of Oe and Be stars seen in our sample is not as surprising as first appears since the presence of Balmer continuum emission will bias the $U-B$ color and would render the star a good candidate for spectroscopy owing to its strongly negative Q value. What we did find remarkable, however, was the large number of these stars that also displayed emission lines of (permitted) Fe II. We illustrate several such spectra in Figure 14*b*. Fe II line identifications were made with reference to the line list of V380 Ori, as given in the paper which defined the characteristics of "Herbig Be" stars (Herbig 1960). However, as Hillenbrand et al. (1993) recently have emphasized, the *spectroscopic* distinction between "Be" and "Herbig Ae/Be" objects is overstated, and the spectral similarities of these stars to Herbig Ae/Be stars should not be used to infer youth. Knowledge of their spectral energy distributions extending into the infrared may be needed to distinguish whether the emission arises in a circumstellar accretion disk or a gaseous circumstellar disk or envelope (Hillenbrand et al. 1992). Historically, emission lines of Fe II have also been used as one of the key signatures of "extreme Be" stars (Schild 1966), although this designation subsequently came to be used for any Be star with higher luminosity than typical (Schild & Romanishin 1976; Garmany & Humphreys 1985). The presence of Fe II emission in these spectra is primarily evidence of an extended envelope (or disk); i.e., simply

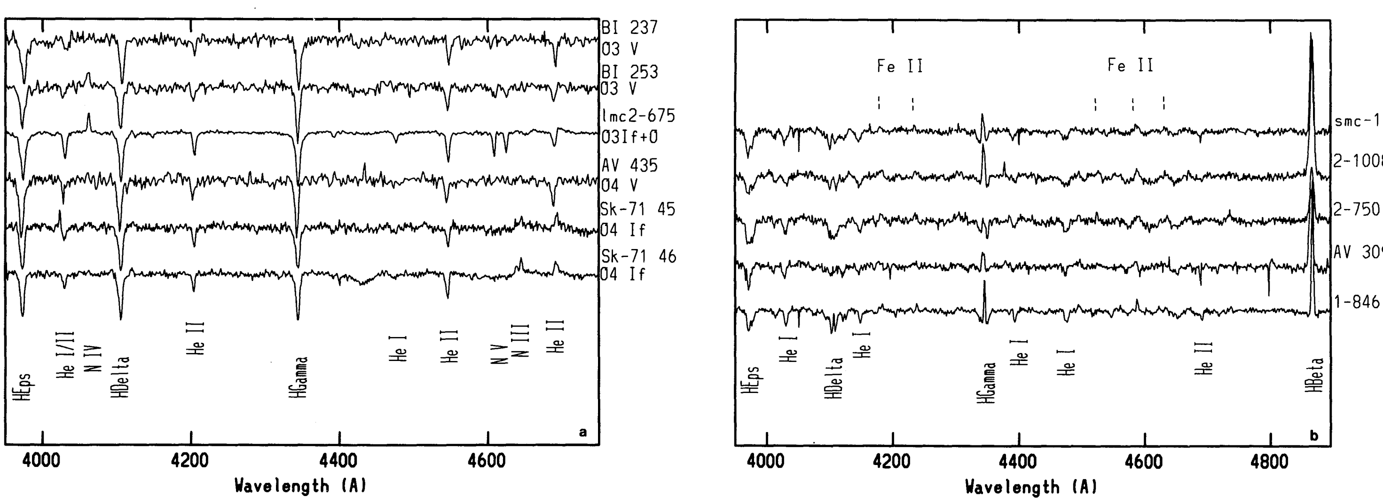

FIG. 14.—(*a*) The spectra of the earliest-type stars observed in this project in the LMC and SMC. (*b*) The spectra of the Be (Fe II) stars. The lines of Fe II marked along the top are λ4173–78; λ4233, λ4522, λ4584, and λ4629.

an extension of the same physical phenomenon that gives rise to the Balmer emission spectrum that causes the "Be" designation in the first place. We further note that our spectra are unlike that of the Zickgraf et al. (1986) "B[e]-supergiants" found in the Magellanic Clouds; although these stars display Fe II emission, the lines are quite narrow, while in ours these lines are quite broad. In addition, the other emission in our spectra are not nearly as strong as that illustrated by Zickgraf et al.

In fact, the closest similarity that we can find to these Be (Fe II) spectra are those of W235 and W503 in NGC 6611, illustrated in Fig. 5*i* of Hillenbrand et al. (1993). Hillenbrand et al. conclude that these stars are spectroscopically similar to the Herbig Be stars BD + 40°4124, but that their infrared signatures clearly link them with "classical Be" stars: their emission lines are likely due to significant mass loss rather than optically thick accretion disks. The strong continuum (as witnessed by the weakness of even He I λ4471 in Fig. 14*b*) is further suggestive that these are simply slightly more "extreme" Be stars than those typically found.

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