

The Stellar Initial Mass Function in 2007: A Year for Discovering Variations

Bruce G. Elmegreen¹,

ABSTRACT

The characteristic mass (M_c) and slope (Γ) of the IMF are reviewed for clusters, field regions, galaxies, and regions formed during cosmological times. Local star formation has a somewhat uniform M_c and Γ . Statistical variations in Γ are summarized, as are the limitations imposed by these variations. Cosmological star formation appears to have both a higher M_c and a slightly shallower slope at intermediate to high stellar mass. The center of the Milky Way may have a shallow slope too. Field regions have slightly steeper slopes than clusters, but this could be the result of enhanced drift of low mass stars out of clusters and associations. Dwarf galaxies also have steeper slopes. Results from the observation of pre-stellar clumps are reviewed too. Pre-stellar clumps appear to have about the same mass function as stars and are therefore thought to be the main precursors to stars. If this is the case, then the IMF is generally determined by gas-phase processes. Brown dwarf formation also shares many characteristics of star formation, suggesting that they form by similar mechanisms.

Subject headings: stars: mass function; stars: formation

1. Introduction

The year 2007 has been a watershed for the discovery IMF variations. This review includes observations of the characteristic mass, IMF slope, field star IMF, and pre-stellar clump mass functions, along with variations in these quantities that have come to light recently. Theoretical considerations are discussed briefly.

¹IBM T.J. Watson Research Center, Yorktown Hts., NY 10598; bge@us.ibm.com

2. A Characteristic Mass

The IMF is usually fitted to a power law at high mass ($M > 1 M_{\odot}$) and a rollover at low mass. The peak in the IMF on a log-log plot is at $\sim 0.3 M_{\odot}$. We think of this as a characteristic mass for star formation, M_c . Globular clusters and the Milky Way bulge have about the same mass range in the IMF plateau as local clusters. Figure 1 summarizes a sample of observations. A wide variety of clusters with diverse environments, locations, metallicities and ages have the same M_c . For example, cluster Blanco 1 is like the Pleiades in age and mass (Moraux et al. 2007), but Blanco 1 is less dense, lies 240 pc off the plane, and has subsolar abundances for [Ni/Fe],[Si/Fe], [Mg/Fe],[Ca/Fe]; still, the IMFs of these two clusters are almost identical. Digel 2N and S are two far-outer Galaxy clusters ($R_{gal} \sim 19$ kpc), yet they have normal IMFs, like Orion’s, for an age of 0.5-1Myr (Yasui et al. 2007).

The same characteristic mass occurs for low density and low pressure regions like Taurus, and high density and high pressure regions like globular clusters (which formed as super star clusters). This implies that the conditions which determine the characteristic mass vary together. If $M_c \sim M_J$ for thermal Jeans mass M_J (Clark & Bonnell 2005; Bate & Bonnell 2005), then $M_J \sim T^{3/2}G^{-3/2}\rho^{-1/2}$ is constant and so $T \propto \rho^{1/3}$ for temperature T and density ρ . If M_J is determined at the point of grain-gas thermal coupling during collapse (Whitworth et al 1998; Larson 2005; Jappsen et al. 2005), which corresponds to some high density, $\rho_{coupling}$, then T has to scale with $\rho_{coupling}^{1/3}$. Heating from outside and inside a star-forming cloud must balance cooling so that $M_J \sim \text{constant}$. How is this possible? Perhaps M_J is not involved after all. M_c could also depend on core sub-fragmentation and protostellar feedback. A recent discussion of what could conspire to make M_J constant is given in Elmegreen, Klessen & Wilson (2008).

There are recent reports of observations that find different values for M_c :

- Getman et al. (2007) studied triggered star formation in the cometary globule IC 1396N and suggested that the average x-ray luminosity is high for the small number of stars present (e.g., compared to Ophiuchus and Serpens). They concluded that the region could have had preferential triggering of massive stars, which means high M_c .
- On a galactic scale, van Dokkum (2008) studied the mass-to-light ratio and U-V for early type galaxies at redshifts in the range $0.02 < z < 0.83$. He noted that the time dependent variation of M/L depends on the IMF slope near $M = 1 M_{\odot}$, which is the mass range that currently dominates the flux. The best fit model had an approximately flat IMF at $1 M_{\odot}$, which means the plateau has to include this mass. He also noted that this result is consistent with Balmer absorption strengths in early type galaxies as a function of redshift.

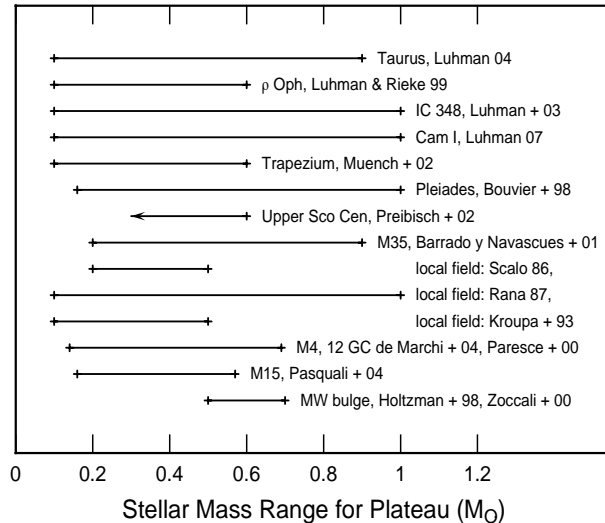


Fig. 1.— Mass ranges for the plateau in the IMF for several clusters and regions of star formation. The plateau is the mass range for the peak of the IMF plotted on a log-log scale, below the power law part characterizing intermediate to high mass stars, and above the apparent drop off in the brown dwarf range.

- Davé (2008) found that the star formation rate per unit mass measured for numerous galaxies in a wide mass range ($10^{9.4} - 10^{12} M_{\odot}$) agrees with that from simulations at $z \sim 0$ but exceeds the simulation predictions for $z \sim 2$. This was explained as a result of increasing IMF characteristic mass with z . High M_c produces a population with a lot of luminosity from star formation but not much build up of mass. He considered the ratio of the high-mass star formation rate to the stellar mass for an evolving IMF compared to a standard IMF and suggested that $M_c = 0.5(1+z)^2$, i.e., the characteristic mass increases monotonically with redshift and was ~ 9 times higher at $z = 2$.
- Fardal et al. (2007) considered constraints from cosmic background starlight and the local luminosity density of galaxies. All of the commonly-observed local IMFs failed to reproduce the observations, while a “Paunchy” IMF, shifted toward higher mass, worked well.
- Komiya et al. (2007) found a high characteristic mass in the IMF of extreme metal-poor (EMP) Milky Way stars. The IMF is constrained by the fractions of EMP stars that are C-rich with s-process and those that are C-rich without s-process. C-rich EMP stars are probably surviving low-mass binary members (models of stellar evolution and binary mass transfer are involved). The theoretical ratio of C-rich-no-s-process to C-

rich s-process stars agrees with the observations only if $M_c \sim 5 M_\odot$. The theoretical fraction of EMP stars that are C-rich s-process also agrees if $M_c \sim 5 M_\odot$. These metal poor stars have $[Fe/H] < -2.5$.

To summarize: local IMFs (mostly in clusters) have a nearly universal M_c . This requires some conspiracy of conditions if M_J is involved. Cosmological evidence suggests M_c was higher in the past, by a factor of ~ 10 , shifting from $\sim 0.3 M_\odot$ today to $\sim 5 M_\odot$ at $z \sim 2$ or more. These are not Population III stars, but normal stars that show up in galaxies today, including the Milky Way. This conclusion about increasing M_c comes from considerations of the mass-to-light ratio and U-V for ellipticals, from the star formation rate per unit mass versus redshift, from cosmic background starlight and the local luminosity density, and from extreme metal-poor stars in the Milky Way.

3. IMF Slope

3.1. Observations

Scalo (1998) summarized various IMF slopes and plotted them as a function of the average mass observed. There was a large scatter in the slope for each mass range, but the general trend and average was consistent with a Salpeter IMF (slope $\Gamma = -1.35$ on a log-log plot) at high mass and a flattening of the IMF (slope ~ 0) around and below $0.3 M_\odot$. Many dense clusters have Salpeter IMFs: R136 in the 30 Dor region of the LMC (Massey & Hunter 1998), h and χ Persei (Slesnick, Hillenbrand & Massey 2002), NGC 604 in M33 (Gonzalez Delgado & Perez 2000), NGC 1960 and NGC 2194 (Sanner et al. 2000), NGC 6611 (Belikov et al. 2004), to name a few. To consider some extreme cases, Westerlund 2, with ~ 5000 stars observed, containing $\sim 7000 M_\odot$, 2 Myr old, at 2.8 kpc distance and behind 5.8 mag extinction, and with mass segregation, has an IMF slope on average that is $\Gamma = -1.2 \pm 0.16$, i.e., approximately Salpeter (Ascenso et al. 2007). NGC 346 in the SMC, which has $\sim 1/5$ solar metallicity, has a Salpeter IMF on average with a radial gradient in the slope (Sabbi et al. 2007). The IMF of the Rosette cluster derived from x-ray observations looks like the Orion IMF from its x-ray stars (Wang et al. 2007).

However, the IMF in the giant Milky Way cluster NGC 3603 is relatively flat with a slope of $\Gamma = -0.74$ (Harayama et al. 2007). There are ~ 7500 stars measured, so the uncertainty in the slope from counting statistics is only ± 0.02 . Mass segregation from dynamical evolution is not significant as there is no IMF steepening beyond the half-mass radius. The uncertainty in the slope from binaries is ± 0.04 . A similar result was found by Stolte et al. (2006), who got a slope of $\Gamma = -0.9 \pm 0.15$, shallower than the Salpeter slope of -1.35 . On the other

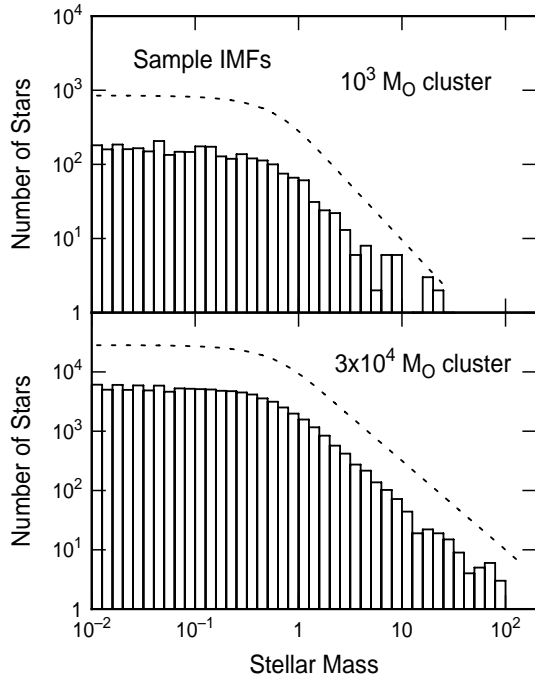


Fig. 2.— Two randomly generated IMFs based on sampling the analytical IMF given by equation 1 with $\Gamma = 1.5$. The top IMF has fewer stars than the bottom IMF, as shown in the label. The break up of the smooth histogram at high mass is to the same degree in each panel because the number of stars close to the maximum likely stellar mass is the same.

hand, Preibisch et al. (2002) found a step IMF in the Upper Sco-Cen star-forming region, which is not a bound cluster. They derived slopes of $\Gamma = -1.8$ for 0.6 to $2 M_{\odot}$ stars and -1.6 for 2 to $20 M_{\odot}$ stars. W51 has a spatially varying IMF with a mean slope of -1.8 in four subgroups but statistically significant excesses in the numbers of the most massive stars in 2 subgroups (Okumura et al. 2000).

3.2. What should the statistical rms in the slope be?

The expected rms can be found by randomly sampling the IMF in a model cluster. We assume model IMFs, $n(M)$, in equal intervals of mass:

$$n(M)dM = M^{-(\Gamma+1)}(1 - \exp^{-(M/M_t)^\Gamma})dM \quad (1)$$

where $\Gamma = 1.5$ and 1.35 (Salpeter function), which is a power law having a turnover at $M_t = 0.5 M_{\odot}$. We assume a maximum stellar mass $M_{max} = 150 M_{\odot}$.

Figure 2 shows two randomly generated IMFs, one for a cluster with $10^3 M_{\odot}$ and another

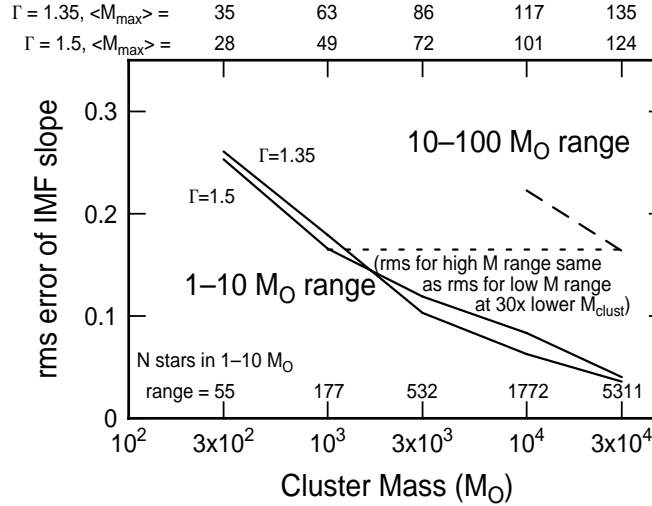


Fig. 3.— The rms deviation in the slope of the IMF is shown versus the mass of the cluster (bottom axis) and the maximum likely mass of the stars (top axis). The two solid lines are for two values of Γ , the intrinsic slope (Eq. 1). The rms is evaluated in two mass ranges, from 1 to 10 M_{\odot} and from 10 to 100 M_{\odot} (for the high mass range, $\Gamma = 1.5$). Only high mass clusters have a well-sampled IMF in the upper stellar mass range. The rms in the high-mass slope of a $3 \times 10^4 M_{\odot}$ cluster is the same as the rms in the intermediate-mass slope of a $10^3 M_{\odot}$ cluster. The number of stars in the 1 – 10 M_{\odot} mass range is indicated along the bottom axis for the $\Gamma = 1.5$ case.

for a cluster with $3 \times 10^4 M_{\odot}$. Clearly the larger cluster has a smoother IMF at low mass because statistical variations are smaller. If we sample the IMFs for 100 clusters of the same mass, we can find the slope for each cluster in a certain mass range, and then find the rms deviations of these slopes around the average value. Figure 3 shows the rms deviations of the slopes versus the cluster mass. The pair of lines ($\Gamma = 1.35$ and 1.5) is for the stellar mass interval between 1 and 10 M_{\odot} and the single dashed line ($\Gamma = 1.5$) is for the 10 – 100 M_{\odot} range. The maximum likely stellar mass in the cluster is shown on the top axis for each assumed IMF slope Γ .

For $10^3 M_{\odot}$ clusters, corresponding to a maximum stellar mass of $\sim 50 M_{\odot}$, the $1 - \sigma$ uncertainty in the IMF slope in the 1 – 10 M_{\odot} range is ~ 0.17 . In the 10 – 100 M_{\odot} range, the IMF uncertainty drops below 0.2 only for clusters more massive than $10^4 M_{\odot}$, corresponding to a maximum stellar mass of $> 100 M_{\odot}$. The rms for the high stellar mass range in a massive cluster is the same as the rms for the low stellar mass range in a cluster that has 30 times lower mass.

Figure 4 shows histograms of the IMF slopes for two cluster masses. For a $10^3 M_{\odot}$

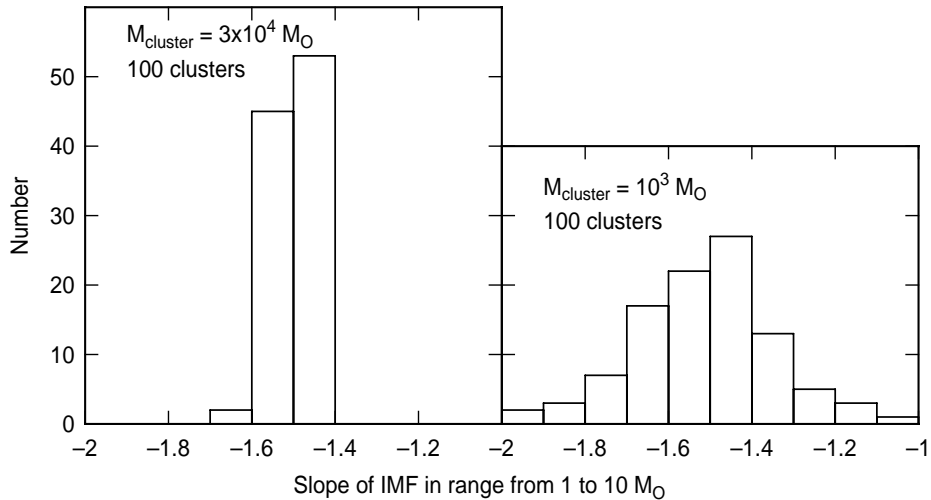


Fig. 4.— Histograms of IMF slopes for $3 \times 10^4 M_{\odot}$ and $10^3 M_{\odot}$ clusters. An IMF with $\Gamma = -1.5$ was used, which reflects the mean value in these histograms. The rms plotted in the previous figure is the breadth of the histogram here for the same cluster mass.

cluster, which is a typically massive cluster in the Milky Way, like NGC 3603, the IMF slope varies from $\Gamma = -1.8$ to -1.1 . Sample IMFs are shown for these two extremes in Figure 5. They look like reputable IMFs without a large amount of scatter from mass interval to mass interval, but their slopes are clearly different. This difference is entirely random. This example illustrates how the *intrinsic* IMF slope cannot be determined to within several tenths for even the most massive clusters in the Milky Way.

The implications of this exercise are worth noting. The upper mass decade for every cluster has an IMF slope uncertainty of ~ 0.2 or more because there is always the same small number of stars in this upper decade (< 500). When the cluster mass is so large ($\sim 10^5 M_{\odot}$) that the maximum possible stellar mass is reached, the number of stars in the upper mass decade goes up and the uncertainty goes down. Second, curvature in the high mass part of the IMF will be difficult to observe or rule out for individual clusters. Similarly, real upper mass IMF variations from cluster to cluster are undetectable.

No cluster IMF has ever been observed throughout the whole stellar mass range. The upper mass IMF needs a very massive cluster, like 30 Dor, massive clusters are rare, and the nearest is too far away to see the low mass stars. On the other hand, the nearest clusters, which are required to study brown dwarfs and low mass stars, are the most common clusters and they are all low mass clusters, having few high mass stars. Until we can observe the lowest mass stars in the highest mass clusters, an IMF makes sense only for an ensemble of clusters or stars.

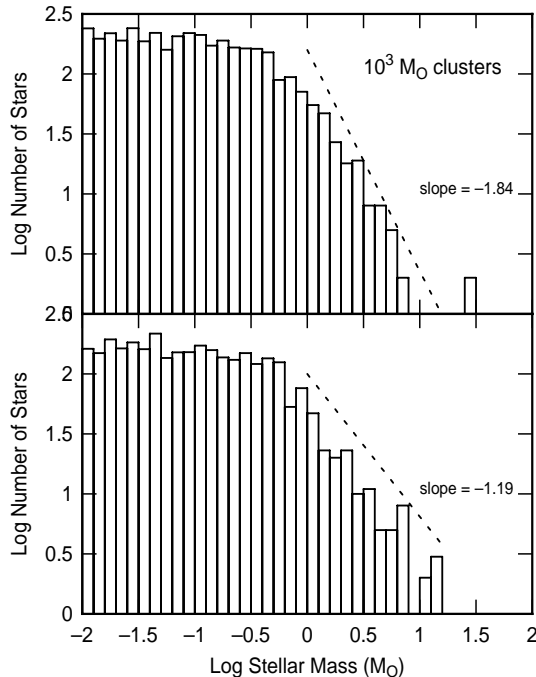


Fig. 5.— Two IMFs in extreme cases of the same randomly sampled model.

4. Is the Field a good place to measure the IMF?

The field is peculiar because there is no star formation there now. A decreasing star formation rate steepens the IMF if the decrease is not accounted for (Elmegreen & Scalo 2006). The field also receives the evaporated stars (low mass stars) from clusters, and the longest-lived stars (low mass stars) from associations, making the field IMF steeper than in clusters even if the total galaxy IMF is the same as in clusters.

Field IMFs can include so many stars that rms errors in the slope are negligible. For example, the field IMF for the LMC measured by Parker et al. (1998) contained 37,300 stars and had a mean slope of $\Gamma = -1.80 \pm 0.09$ for $M > 2 M_{\odot}$. Massey et al. (1995, 2002) found a steep IMF in the remote fields of the LMC ($\Gamma = -4$) and SMC ($\Gamma = -3.6$), more than 30 pc from a Lucke & Hodge (1970) or Hodge (1986) association. Their IMF was complete down to $25 M_{\odot}$ and they assumed a constant star formation rate over the last 10 My; 450 stars were in the in LMC sample, for which we would predict an rms uncertainty of ± 0.15 , which is relatively small. Another field region near LMC4 has a slope of $\Gamma = -5$ for $M = 0.9 - 2 M_{\odot}$ and $\Gamma = -2.6$ for $M = 0.9 - 6 M_{\odot}$ (Gouliermis et al. 2005). However, a field region near 30 Dor has the Salpeter IMF: $\Gamma = -1.38 \pm 0.04$ for $M = 7 - 40 M_{\odot}$ (Selman & Melnick 2005).

Chandar et al. (2005) compared field IMF spectra to evolution models in nearby starburst galaxies, often finding a steeper IMF slope than Salpeter, by ~ 0.5 , or a low maximum stellar mass as an alternative explanation. Similarly Úbeda et al. (2007a,b) found a steep IMF in the dwarf galaxy NGC 4214, $\Gamma = -1.83 \pm 0.07$, by counting main sequence stars in intervals of the HR diagram. Field and cluster IMFs were not much different in that study. They suggested that Γ could be even steeper than this. Also, from IUE spectra in this galaxy, Mas-Hesse & Kunth (1999) suggested $\Gamma = -2.0$.

Hoversten & Glazebrook (2007) fit a star formation model to $H\alpha$ equivalent widths and g-r colors in 140,000 SDSS galaxies, minimizing over variations in Γ , metallicity, ages, and star formation history. They found an IMF slope that got steeper for lower mass galaxies, ranging from the Salpeter value for $M_R = -22$ mag to $\Gamma = -1.6$ for $M_R = -17$ mag.

The Milky Way and M31 bulge [Fe/H] abundances suggest shallow IMFs at intermediate to high mass, $\Gamma \sim -1$ to -1.1 (Ballero et al. 2007ab). The central region of the Galaxy may have a shallow current-day IMF too, $\Gamma = -0.85$ (Paumard et al. 2006), based on 40 OB supergiants, giants, and main-sequence stars in 2 rotating disks within 0.5 pc of the nucleus; the ages of these stars are 6 ± 2 Myr and the total mass is $1.5 \times 10^4 M_\odot$. Also in the center, 329 late-type giants within 1 pc of SgrA* have $\Gamma = -0.85$ for 12 Gyr of SF (Maness et al 2007). In a third study of the Galactic center, Nayakshin & Sunyaev (2005) found few x-ray stars associated with the massive stars orbiting Sgr A* and proposed in situ star formation with a shallow IMF. The IMF in the Arches cluster based on deep AO images is also slightly flatter than Salpeter, $\Gamma = -1$ to -1.1 ; Fokker-Planck models were used to account for mass segregation over the age of the cluster (Kim et al. 2006).

What about “top heavy” IMFs in super star clusters? There are several observations that suggest this. Sternberg (1998) found a high value of L/M in NGC 1705-1 that suggests either $|\Gamma| < 1$ or an inner mass cutoff for stars. Smith & Gallagher (2001) found the same for M82F: a high inner cutoff of 2 to 3 M_\odot if $\Gamma = -1.3$. Alonso-Herrero et al. (2001) found a high L/M in the starburst NGC 1614. McCrady et al. (2003) suggested that MGG-11 in M82 is deficit in low mass stars. Mengel et al. (2002) got the same in NGC 4038/9. But other super star clusters have normal IMFs: NGC 1569-A (Ho & Filippenko 1996; Sternberg 1998), NGC 6946 (Larsen et al. 2001), M82 MGG-9 (McCrady et al. 2003). It is unclear if any of these super star cluster IMFs are correct. Bastian & Goodwin (2006) found that all of the odd IMFs in SSCs are in the youngest clusters, and suggested that these young clusters may not be relaxed enough to give an accurate dynamical mass.

There are other IMF oddities too. Massive elliptical galaxies have slightly flatter-than-Salpeter IMFs in studies by Pipino & Matteucci (2004) and Nagashima et al. (2005b). Clusters of galaxies suggest a history of flat IMFs in elliptical galaxy bursts (Renzini et

al. 1993; Loewenstein & Mushotsky 1996; Chiosi 2000; Moretti, Portinari, & Chiosi 2003; Tornatore et al. 2004; Romeo et al. 2004; Portinari et al. 2004; Nagashima et al. 2005a). Low surface brightness galaxies may have steep IMFs, $\Gamma \sim -2.9$ (Lee et al. 2004; Hoversten & Glazebrook 2007; Úbeda et al. 2007). The red halos in BCD galaxies and stacked halos around disks have $\Gamma = -3.5$ (Zackrisson, et al. 2004). Also, more locally, there are apparently bare or single-forming O stars that represent 4% of O stars (deWit et al. 2005), and there are O-stars on the periphery of clusters, possible triggered or not mass segregated.

In summary, the Salpeter slope ($\Gamma = -1.35$) typically occurs in clusters with marginally insignificant variations. A steeper IMF seems to be the rule for field regions, although there are uncertainties in the star formation history and there are possibly important effects from low-mass stellar drift out of nearby clusters and associations. Slightly steeper IMFs are found for low mass or low surface brightness galaxies, $\Gamma \sim -1.5$ to -2.9 . Slightly flatter IMFs are found for extremely active star-forming regions in the Milky Way and M31 bulges, the Milky Way nucleus, and bursting ellipticals, all of which have $\Gamma \sim -1$.

5. Theory: Origins of Stars and the IMF

There are three types of theories for the IMF:

- Fragmentation (“top down”), most likely driven by turbulence & self-gravity (“turbulent fragmentation”). Simulations of this process have been studied by several teams (e.g., Tilley & Pudritz 2007; Padoan et al. 2007; Li et al. 2004; Nakamura & Li 2005, 2007; Martel, Evans & Shapiro 2006; see review by MacLow & Klessen 2004). This process is most clearly related to the initial phases of star formation, and may also be most relevant to intermediate stellar masses, possibly including the IMF turnover at $M_c \sim 0.3 M_\odot$. It may be what connects the characteristic mass, M_c to the thermal Jeans mass M_J .
- Accretion and coagulation (bottom up), where clumps accrete from filaments and sheets to grow into stars (including “competitive accretion”; see review by Bonnell et al. 2007). Also included in these models are situations where clumps or protostars coagulate inside dense clusters. This may be most relevant to the late stages in star formation, and to the most massive stars. Coagulation or capture should form some binaries, and protostar interactions should influence the disk fraction.
- “Interruption” to fragmentation and accretion, where either of the previous two processes gets interrupted so the final stellar mass is smaller than it would have been.

Included here are models for core ejection from dense groups, and core ionization (Boss 2001, Reipurth & Clarke 2001, Bate et al. 2002; Whitworth & Zinnecker 2004; Whitworth & Goodwin 2005; Goodwin & Whitworth 2007; Umbreit et al. 2005; and others). This process may be most relevant to brown dwarfs.

6. Brown Dwarfs: Fragmentation or Ejection?

Several observations suggest that brown dwarfs and stars form by the same mechanism. First, brown dwarfs and stars have the same spatial distribution in Taurus (Briceno et al. 2002; Luhman 2004a, 2006), whereas ejection in the “interruption” model should produce more dispersed brown dwarfs than stars (Kroupa & Bouvier 2003). Second, the disk fraction for brown dwarfs is about the same as the disk fraction for stars, and also about the same as the fraction of isolated planetary mass objects ($\sim 30\%$; Luhman et al. 2006; Scholz & Jayawardhana 2007). This is also contrary to the ejection scenario, which might be expected to produce fewer or smaller disks in brown dwarfs. Third, there is a binary brown dwarf in Cha I with a 240 AU separation; this is a fragile system that could probably not have been ejected (Luhman 2004b). Fourth, the accretion rate scales with protostellar mass as $M^{2.1}$, spanning the stellar to brown dwarf boundary without a change (Muzerolle et al. 2005). Fifth, the lack of brown dwarf companions to stars in wide binaries equals the lack of free-floating brown dwarfs relative to stars (Luhman et al. 2005). This implies that free floating brown dwarfs form the same way as binary brown dwarfs.

7. Mass Functions of Pre-Stellar Cores

An important clue to the IMF is that the core mass function is often similar to the stellar IMF, both in the slope at the high mass end and in the turnover around $1 M_{\odot}$. Table 1 lists observations of core mass functions for low masses and similarly steep mass functions for high mass cores. It also lists mass functions for high mass cores that have a shallow slope, like the slope of the GMC mass function.

There are several key properties of the pre-stellar cores that have steep mass functions. First, they move at sub-virial, near sonic, speeds, which implies that they do not accrete much in the Bondi-Hoyle fashion. Measured velocity dispersions are $\sim 0.17 \text{ km s}^{-1}$ (Di Francesco et al. 2004) or $\sim 0.4 \text{ km s}^{-1}$ (André et al. 2007), and so on (Belloche et al. 2001; Walsh et al. 2004, 2007; Jorgensen et al. 2007; Kirk et al. 2007).

Second, they resemble Bonner-Ebert spheres, which means that they are pressure-bound

Table 1. Observations of Proto-Stellar Core Mass Functions

Reference	Region	Method	Number of Cores	Density cm^{-3}	Mass Range M_{\odot}	Slope Γ
Low Mass Cores						
Motte et al. 1998	ρ Oph	1.3mm	58	10^6	0.5 – 3	–1.5
Testi & Sargent 1998	Serpens	1.3mm	26	10^7	0.3 – 30	–1.1
Johnstone et al. 2000	ρ Oph	850mm	55	$10^{6.5}$	0.02 – 6.3	–1 to –1.5
Coppin et al. 2000	Orion A North	450mm/850mm	67	10^5	0.1 – 100	–0.5(shallow)
Kerton et al. 2001	KR140	450/850mm	22	$10^{4.3}$	0.5 – 130	–0.5 (shallow)
Johnstone et al. 2001	Orion B	850mm	75	10^5	0.2 – 12.3	–1.5
Bontemps et al. 2001	ρ Oph	6.7mm/14.3mm	123 Cl.II	...	0.02 – 3	–1.7
Motte et al. 2001	NGC 2069/2071	450mm/850mm	70	10^7	0.3 – 5	–1.1
Sandell et al. 2001	NGC 1333	450mm/850mm	33	10^7	0.03 – 1	–0.4 (near pk only)
Tachihara et al. 2002	Tau, Oph, Lupus L1333, CorAust, Coalsack, Pipe neb.	C^{18}O	174	10^4	1 – 400	–1.5 to –2.6
Onishi et al. 2002	Taurus	N^{13}CO^+	44	10^5	3.5 – 20	–1.5
Tohill et al. 2002	M8	450/850/1.3mm	37	$10^{4.5}$	0.5 – 20	–0.7 (shallow)
Stanke et al. 2006	Oph	1.2mm	111	10^5	1 – 8	–1.6
Enoch et al. 2006	Perseus	1.1mm	122	10^5	1 – 30	–1.6
Johnstone et al. 2006	Orion B S	450mm/850mm	57	10^6	3 – 30	–1.5
Johnstone & Bally 2006	Orion A S	450mm/850mm	71	10^6	0.3 – 22	–1
Chi & Park 06	Polaris	^{13}CO	105	10^3	0.1 – 10	–0.91 (low density)
Young et al. 2006	Oph	1.1mm	44	10^6	0.24 – 3.9	–1.1
Reid & Wilson 2006	M17	450mm/850mm	100	10^5	0.8 – 120	–0.5 to –0.9
Kirk et al. 2006	Perseus	850mm/extinction	58	10^5	0.3 – 5	–2
Alves et al. 2007	Pipe Nebula	extinction	159	10^4	0.5 – 28	–1.4
Ikeda et al. 2007	Orion A	H^{13}CO^+	236	$10^{4.3}$	0.5 – 2	–1.5
Nutter et al. 2007	Orion NS active	450mm/850mm	393	10^8	0.1 – 40	–1.2
Li et al. 2007	Orion S quiescent	sub-mm	51	10^7	0.1 – 46	+0.15
Walsh et al. 2007	NGC 1333	N^2H^+	93	10^6	0.05 – 2.5	–1.4
Massi et al. 2007	Vela Cloud D	1.2mm	29	10^5	0.04 – 88	–0.45 (shallow)
High Mass Cores						
Shirley et al. 2003	many sources	CS	57	10^5	$10^2 – 10^4$	0.91
Reid & Wilson 2005	NGC 7538	450mm/850mm	67	10^5	$10^2 – 10^{3.5}$	–1
Rathborne et al. 2006	IR Dark Clouds	1.2mm	120	$10^{4.5}$	$10 – 10^{3.3}$	–1.1
Cluster Mass Slope						
Moore et al. 2007	W3	850mm	316	10^5	13 – 2500	–0.8
Beltran et al. 2006	Southern sources	IRAS	235	10^6	$10^2 – 10^{3.6}$	–0.9
					10 – 120	–0.5
Munoz et al. 2007	NGC 6334	1.2mm	181		3 – 6000	–0.6

and self-gravitating, perhaps stable, although sometimes there is evidence for inflow (e.g., André et al. 2007). Note that some theoretical derivations of the characteristic mass are dynamic, which means they occur during the collapse and not in equilibrium. For example, in Larson’s (2005) model, the characteristic mass in the IMF is the thermal Jeans mass at an inflection point in the equation of state. This inflection is important during the collapse as it determines the point where subfragmentation stops (in the soft part of the equation of state) and the existing fragments either keep their mass or grow by accretion (in the hard part of the equation of state).

Third, the mass fractions of the pre-stellar cores inside their clouds are ~ 2 to 10%, like the average star formation efficiency. This implies the observed cores are evolving toward individual stars or binary systems and nearly all the stars that will ever form in the cloud are currently observed in the form of cores.

Fourth, the cores cluster together hierarchically, like the young stars they will eventually form (Johnstone et al. 2000, 2001; Enoch et al. 2006; Young et al. 2006).

All of the pre-stellar cores that are observed with a particular method in a region have about the same internal density, which is the density where that method is most sensitive to finding them. Overall, the pre-stellar cores have a wide range of densities, considering different methods and regions. The densities in the table range from 10^4 to 10^7 cm^{-3} , even though the mass functions are all about the same.

These observations suggest that the stellar IMF is determined in the cloudy phase by gas processes, prior to significant collapse. This differs from the IMF explanation in Bonnell et al. (2007), where the stellar masses are determined by accretion onto existing cores whose initial core mass function is not important and may even be a delta function. Bonnell et al. point out, however, that most pre-stellar cores are not strongly self-gravitating, and they suggest that the cores may not even form stars, or at least not on a one-to-one basis with a preservation of the mass function.

Steep core mass functions are sometimes found for higher mass cores too. The lower part of Table 1 lists the observations of these. Because the cluster mass function has $\Gamma \sim 1$, these cores could form clusters. Other mass functions at high mass are shallower than this, like the GMC mass function. Examples of these shallow functions are at the bottom of the table.

There are many detailed and interesting results about pre-stellar core mass functions that have some bearing on the star formation process. For example, the mass function for class II sources in Ophiuchus is the same as the mass function for pre-stellar cores there (Bontemps et al. 2001). This illustrates well the picture that cores evolve into stars. The

core definition most likely depends on resolution, however. Enoch et al. (2007) found that the Serpens core mass function is flatter (slope of -1.6) than the Perseus and Ophiuchus mass functions (slopes of -2.1), all observed with the same technique, and the distance to Serpens is the smallest. Yet all three surveys have the same core size distribution relative to the beam size. Thus, smaller cores are inferred for the closest regions.

Ikeda et al. (2007) studied Orion A with H^{13}CO^+ at a characteristic density of $10^{4.3} \text{ cm}^{-3}$. They got core mass functions with a turnover at low mass like the IMF, which is similar to what other prestellar core studies found (as listed in Table 1), but they got these turnovers only when they did not consider blending corrections. They got a straight power law mass function with no turnover when they included blending corrections. They concluded from this that the turnover at low mass, which is one of the key features that makes a pre-stellar core mass function look like a stellar IMF, is the result of blending and source confusion. If this is true in general, then we should not conclude without further study that pre-stellar cores evolve into stars on a one-to-one basis.

Another curious result is that the starless cores seen at 1.2mm in Ophiuchus have mass segregation (Stanke et al. 2006): the mass functions look the same in the inner and outer regions, but they extend to higher mass in the inner regions.

Young et al. (2006) and Enoch et al. (2007) found that pre-stellar cores are spatially correlated, which means they are hierarchically clustered in the same way that stars are. This is further evidence they will turn into stars without moving much further, considering their velocity dispersions are low (see above).

Chi & Park (2006) studied the polaris flare, which is a weakly self-gravitating (diffuse) cloud visible in ^{12}CO and ^{13}CO at a density of $\sim 10^3 \text{ cm}^{-3}$. The cores in this cloud have a steep mass function too, $\Gamma = -0.91 \pm 0.13$. This suggests that steep, Salpeter-like mass functions are made independently of the star formation processes, long before self-gravity is important in the gas. They seem to be characteristic of turbulence. For example, a three-dimensional region with a Kolmogorov power-law power spectrum and a log-normal density probability distribution function has a mass function for clumps that is shallow at a low threshold density and steep at a high threshold density (Elmegreen 2002; Elmegreen et al. 2006). The high-density mass function has the same slope as the Salpeter IMF and the low-density mass function has the same slope as the GMC mass function. Thus the mass functions for low density clouds and pre-stellar cores may be sampled from the same scale-free density distribution, but sampled at different density thresholds.

8. IMF Origins

My best guess at the present time is that the IMF is a superposition of several processes (Elmegreen 2004). There is an isolated star-formation mode which is dominated by fragmentation near the thermal Jeans mass, M_J . There is a dense cluster mode which adds to this some extra accretion and clump coagulation to boost up the high-mass component relative to that in the isolated mode. And finally, there is the tiny dense-core mode where star mass is heavily influenced by ejections from multiple systems and ablation due to nearby stars. These latter processes modulate the isolated mass function in the brown dwarf and sub- M_J range.

Starbursts and major mergers (forming elliptical galaxies) could have more of the dense cluster mode, which should give them slightly flatter IMFs, as is sometimes observed. Low surface brightness galaxies, dwarf galaxies with low ISM pressures, and quiescent regions like the Taurus clouds have less of this dense cluster mode and slightly steeper IMFs. This difference is not just a temperature, pressure, or M_J effect. It is the result of different combinations of distinct physical processes.

9. Maximum Stellar Mass

The IMF appears to have a maximum stellar mass of $120 - 150 M_\odot$ (Weidner & Kroupa 2004; Oey & Clarke 2005; Koen 2006). This may not be the maximum mass as far as the star formation process is concerned, however. The IMF could in principle go to much higher masses, but by the time the star appears as a pre-main sequence or main-sequence object, it has lost so much mass via intense winds that $\sim 150 M_\odot$ is all we can see. Figure 6 shows the time-dependent masses of stars that have the mass loss rates indicated. Mass loss rates are assumed to depend on the remaining mass. The initial masses are taken to be large, from $100 M_\odot$ to $500 M_\odot$, but after only 10^5 yrs, they have all come down to the $100 - 200 M_\odot$ mass range. This is the time when they are likely to appear. The assumed mass loss rates are not unreasonably high for young O-type stars.

10. Summary

The characteristic mass and slope of the IMF both show variations, mostly in extreme environments. The characteristic mass increases and/or the slope flattens a little ($\Gamma \sim 1$ compared to the Salpeter IMF which has $\Gamma = 1.35$) in several regions that formed at high redshift: the Milky Way and M31 bulges, the Milky Way extremely metal-poor stars,

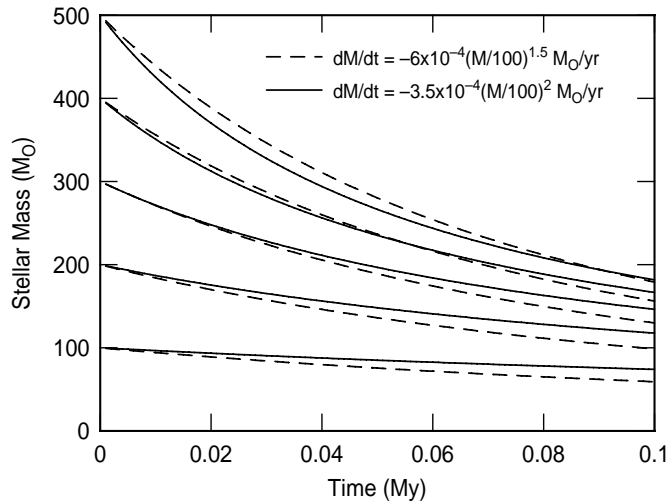


Fig. 6.— A model of stellar mass versus time when a strong wind is present. The processes that determine the IMF may be able to form $500 M_{\odot}$ or larger “protostars”, but these objects will not show up as real stars if the mass loss rate is as high as $\sim 10^{-3} M_{\odot} \text{ yr}^{-1}$.

elliptical galaxies, clusters of galaxies, and other high redshift star formation. Perhaps the common physical process is that these regions formed by major mergers and starbursts in the early universe. The Milky Way nucleus also appears to have a somewhat flattened IMF, and a local bright rim of presumably triggered star formation seems to have a high characteristic mass. On the other hand, the IMFs for modern starburst galaxies or super star clusters do not seem to be much different from the Salpeter function. Perhaps they are not extreme enough.

At the other extreme, dwarf galaxies and low surface brightness galaxies have slightly steeper IMFs, $\Gamma \sim -1.6$. The field regions between clusters and OB associations also have systematically steeper IMFs.

Of the 3 models for the IMF discussed here (top down, bottom up, interruption), the top-down model has the most direct evidence. This evidence comes from the mass distribution functions of pre-stellar cores and stars and from the slow random motions of pre-stellar cores. Pre-stellar core mass functions resemble the IMF, which implies that the IMF may be determined in the gas phase, with possible modifications at high and low mass from stellar-scale processes, such as fragmentation, collisions, and system ejections.

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