The IMF and its Evolution

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ABSTRACT

Observations of the stellar initial mass function are reviewed. The IMF is flat, or possibly declining, below several tenths of a Solar mass, and declining above this mass in a power law with a slope of about $-1.35$ on a log-log plot. The flattening at low mass is evidence for a characteristic mass in star formation, which, according to recent theory, may be either the minimum stellar mass for the onset of deuterium burning, or the thermal Jeans mass. The first of these masses should not vary with environment as much as the second, so any observed variations in the mass of the flattened part are important for understanding star formation. Starburst galaxies may be an example where the characteristic mass is larger than it is locally, but this old observation has been challenged lately. A steeper high-mass slope in the extreme field studied by Massey et al. may be the result of cloud destruction and the termination of star formation by ionization, with a normal IMF in each separate cluster. The lack of a density dependence in cluster IMFs suggests that star and protostar interactions play little role in star formation or the IMF. This is unlike the case for binaries and disks, which do show an environmental influence, and all are consistent with the observed stellar density in clusters, which is high enough to promote interactions between binaries and disks, but not individual stars. These considerations, along with indirect observations of the IMF in the early Universe, suggest that the IMF does not vary much in its basic form over position and time, but that shifts in the characteristic mass might occur in regions with extremely high or low star formation activity, or perhaps light-to-mass ratio, with the characteristic mass, star formation efficiency, and gas consumption rate all following the light-to-mass ratio.

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1. Introduction

Observations of clusters and associations suggest an average stellar initial mass function (IMF) that is approximately a power law like the Salpeter (1955) function, with a slope of $x \sim 1.35$ on a log $n - \log M$ plot, and a flattening below $\sim 0.35 M_\odot$. This IMF appears in clusters and whole galaxies, for all galactic populations, and even in the intergalactic medium (Sect. 2.1, 2.3, 2.4). However, there are still fluctuations in the slope of the power-law by $\pm 0.5$ from cluster to cluster (Scalo 1998), and there are other curious variations too, like a steeper slope in the field (Sect. 2.2), the mass of the most massive star increasing with cloud mass (Sect. 3), the formation of massive stars relatively late and near the centers of clusters (Sect. 3), and the greater proportion of massive stars in starburst galaxies (Sect. 2.5). Considering the robust nature of the IMF, any theory for its origin should be able to reproduce both the average shape and the variations around it with a minimum of free parameters and a minimal dependence on the physical properties of the star-forming clouds.
Another important mass function for star formation is the distribution of cloud and clump masses. This differs from the stellar function in both slope ($x \sim 0.5 - 0.8$ for clouds) and range ($M_{\text{cloud}} \sim 10^{-4} - 10^7 M_\odot$; Heithausen et al. 1998; Dickey & Garwood 1989), leading one to wonder why stars form with a steeper mass distribution than their clumps. There must be a preferential selection of lower clump masses for stars, and a cutoff at some minimum star mass.

There are tantalizing indications that we may be able to understand the IMF without fully understanding the origin of either the cloud structure or the processes involved with individual stars. Given the observed structures of clouds, we can imagine how star formation processes might select pieces of this structure in a certain order and end up with the observed IMF and all of its variations. If such clump selection is the correct explanation for the IMF, then it presumably works because most of the star mass is determined by the gas mass immediately available to it during the protostar phase, and because the IMF is an average over many different processes, with each losing its unique contribution when the mass distribution is averaged over a cluster.

Numerical simulations of such sampling demonstrate this point by reproducing essentially all of the observations of the IMF and its systematic and stochastic variations without any free parameters or physical input other than a single characteristic mass for the minimum clump that can make a star. These models obtain (Elmegreen 1997, 1999a): (1) the correct power-law slope and turnover shape of the IMF, with the correct turnover mass, (2) the tendency for the most massive star in a cluster to increase with cloud mass, (3) the shift in the peak or turnover mass for starburst regions without a change in the power-law slope, (4) the delayed formation of massive stars in a cluster, (5) the fluctuations in the slope of the power-law part from cluster to cluster (which result from sampling statistics), and (6) the tendency for the most massive stars in a cluster to concentrate toward the cluster center. The only input to the model is the hierarchical (and fractal) distribution of cloud structure, and the only assumption is that pieces of this hierarchy make stars at a rate that scales with the square root of the local density, which is the rate at which essentially all of the physical processes involved with the onset of star formation operate, including self-gravity, magnetic diffusion, clump collisions, and turbulence dissipation, given the molecular cloud scaling laws.

The hypotheses that IMF theories may be simplified by the gross averaging of star formation processes during the build up of a cluster, and by the intimate connection between its power-law slope and cloud structure, also help to explain why its power-law slope is so similar from region to region, even in different environments and at different times. The point is that the cloud and star formation details may not matter much for the IMF, and that power-law cloud structures are more-or-less universal, perhaps as a result of pervasive turbulence.

In the next section we review the observations of the IMF and some of the implications of these observations in an attempt to sort out what is physically significant and compelling for a theory. Other reviews can be found in the conference proceedings The Stellar Initial Mass Function, edited by Gilmore, Parry & Ryan for ASP Press in 1998. A review that compares various theories with the constraints from observations is in Elmegreen (1999b).

2. Observations of the IMF and Implications for the Theory
2.1. The Salpeter Slope in Clusters and Galaxies

The IMF at intermediate to high mass can be written \( n(M)d\log M \propto M^{-x}d\log M \) for slope \( x \) on a log – log plot. For most clusters, \( x \) is in the range 1–1.5. Salpeter (1955) suggested \( x \sim 1.35 \), which is about the average of the values observed today. The most dependable values come from a mixture of photometry and spectroscopy of star clusters. IMFs based on photometry alone are generally steeper than \( x \sim 1.35 \) because of an ambiguity in mass for high mass stars (see discussion in Massey 1998).

Table 1 summarizes the recent observations that obtain \( x \sim 1 – 1.5 \) in various regions. This “Salpeter” slope is found by star counts in local clusters, integrated light from whole galaxies, elemental abundances, and galaxy evolution models. Steeper values of \( x \sim 1.5 – 2 \) are found in samples of local field stars or in the low density parts of some clusters (Table 2). Shallower values are found at low mass, where the IMF flattens to nearly zero slope on a log – log plot (Table 3). Shifts either in the peak or in the slope, favoring higher masses, have been found in starburst galaxies (Table 4).

The observations in these tables suggest that the IMF varies a lot, but in fact most of the functions that deviate from the turned-over Salpeter slope are based on indirect measurements that contain questionable assumptions. For example, the slope determined for the local field tends to get steep only at high mass, and the increased value depends on an assumed recent star formation history and an assumed scale height variation with mass and age. The local field is also more populated by low mass stars than high mass stars because low mass stars live longer and drift further from their sites of star formation than high mass stars.

The low density regions of clusters show a steeper IMF too because of an excess of low mass stars, but this is probably related to the greater concentration of high mass stars in cluster cores, as discussed more in Section 3; the overall cluster can still have a flattened-Salpeter IMF. The Hipparcos results quoted by Brown (1998) were based on photometry, rather than spectra, and are typically steep for photometry. Massey et al. (1995) has shown how such IMF values become shallower, like the Salpeter function, when spectra are considered for the determination of stellar mass.

2.2. A Steep IMF Slope in The Extreme Field

The most extreme deviation for an IMF measurement is in the remote field regions of the LMC and Milky Way (Table 2). These are regions defined by Massey et al. (1995) to be further than 30 pc from the boundaries of catalogued OB associations. Here the slope at high mass has been measured to be around \( x \sim 4 \). Evidently something very unusual is happening. There are several ways to explain this, if it turns out to be true. One way has a normal \( (x \sim 1.35) \) IMF in every individual region of star formation, and a steeper IMF in the composite of many regions. This difference between cluster and integrated IMFs illustrates an important point about cloud destruction, so we discuss it in some detail here (see also Elmegreen 1999a).

In a large region there will in general be many separate clouds that form stars, and these clouds will have some mass function \( n(M_c)dM_c \propto M_c^{-\gamma}dM_c \) for \( \gamma \sim 1.5 – 2 \). If intermediate and high mass stars destroy their clouds because of ionization, and as a result, halt the star formation processes inside them, then more massive clouds will require more massive stars before star formation ends. This leads to a situation where a lot of low mass clouds make primarily low mass stars, with a normal IMF, and where a few high mass clouds make both low mass and high mass stars, also with a normal IMF. But, since there are more low mass clouds, the composite region will have a lot more low mass stars in proportion to high mass stars than is given by each cluster IMF. It follows that even if the IMF inside each region of star formation is the
Table 1: Observations of a Salpeter IMF with $x = 1 - 1.5$

| Star counts in clusters | Elson, Fall, & Freeman 1989; Massey & Thompson 1991; Vallenari et al. 1992; Elson 1992 Massey & Johnson 1993; Phelps & Janes 1993 Hillenbrand et al. 1993; Drissen et al. 1993 Parker & Garmany 1993; Ninkov et al. 1995 Massey et al. 1995; Chiosi et al. 1995; Banks, Dodd, & Sullivan 1995; Will, Bomans, & de Boer 1995; Will et al. 1995, 1997; Deeg & Ninkov 1996; Hunter et al. 1996a,b, 1997 Massey & Hunter 1998 |
| Star counts in the field | Scalo 1986 ($x \sim 1 - 1.5$ for intermediate mass stars) |
| Hα equiv. widths and galaxy photometry | Greggio et al. 1993; Kennicutt, Tamblyn, & Congdon 1994; Marconi et al. 1995 Bresolin & Kennicutt 1997; Holtzman, et al. 1997; Grillmair et al. 1998 |
| Elemental abundances in local ISM | Tsujimoto et al. 1997 |
| Galactic stellar halo | Nissen et al. 1994 |
| QSO damped Lyα | Lu et al. 1996 |
| Lyα forest | Wyse 1998 |
| Intracluster medium | Renzini et al. 1993, Wyse 1997, 1998 (but see Loewenstein & Mushotzky 1996) |
| Elliptical galaxies | Wyse 1998 |
| Galaxy evolution models | Sommer-Larson 1996; Lilly et al. 1996 |

Table 2: Observations of the IMF with $x = 1.5 - 2$ or greater

| Local field stars | Miller & Scalo 1979; Garmany, Conti, & Chiosi 1982; Humphreys & McElroy 1984 Blaha & Humphreys 1989; Basu & Rana 1992 Kroupa, Tout, & Gilmore 1993; Scalo 1986 ($x = 1.5 - 2$ for high mass stars) |
| Local OB associations | (review of Hipparcos results: Brown 1998) |
| LMC clusters in regions | J.K. Hill et al. 1994; R.S. Hill et al. 1995 |
| Low young-star density | |
| Unclustered embedded stars in Orion | Ali & DePoy 1995 |
| Extreme field stars in the LMC ($x \sim 4$) | Massey et al. 1995 |
for \( \alpha > 0 \) which is \( M_{\text{stellar mass}} \) comes from the mass-luminosity relation of ionizing radiation, which scales as \( L \propto M^4 \) (Vacca, Garmany, & Shull 1996). A whole cluster’s ionizing luminosity can be evaluated from the expression \( \int_0^{M_L} L(M) n(M) dM \) for maximum mass \( M_L \) and IMF \( n(M) dM = x M_L^c M^{-1-x} dM \). This cluster luminosity scales with \( M_L^c \) too. The constant term in the IMF, \( x M_L^c \), gives one star at a maximum mass \( M_L \) from the expression \( \int_0^{M_L} n(M) dM = 1 \). The luminosity required to destroy a cloud is the binding energy divided by the cloud crossing time, which is \( (G M_c^2 / R_c) \left( G M_c / R_c \right)^{1/2} \propto M_c^{5/4} \), using the Larson (1981) scaling laws for molecular clouds. Setting the luminosity of a cluster, \( \propto M_L^c \), equal to the power required to destroy a cloud, \( \propto M_c^{5/4} \), then gives \( \alpha = 5/16 \) in the expression \( M_L \propto M_c^\alpha \).

With \( \gamma = 2 \) and \( \alpha = 5/16 \), the slope of the composite IMF is \( x_{\text{comp}} = (\gamma - 1) / \alpha = 16/5 \sim 3.2 \). The value observed by Massey et al. (1995) is \( \sim 4 \), which is pretty close to this theoretical result, given the uncertainties in the \( M - L \) relation and other assumptions, and with the observations.

It is important to note that the extreme field IMF found by Massey et al. (1995) is not representative of galaxies in general. Integrated light and elemental abundances give an average IMF for whole galaxies that has the same slope at intermediate and high mass as individual clusters, namely, the Salpeter value of \( x \sim 1.35 \). This simple fact implies that massive stars cannot generally halt star formation in their clouds. If they did, then the composite IMF for a whole galaxy would be significantly steeper than the individual IMF in each cluster. Massive stars may destroy their clouds, in the sense that they push the gas around, but they cannot generally halt star formation in them except possibly in the extreme field. The extreme field could differ from the environment in OB associations because of a much lower pressure in the extreme field. A low pressure could conceivably lead to more efficient cloud ionization and the cessation of star formation in even the dense clumps.

The requirement that the composite IMF be equal to the cluster IMF also means that \( \alpha = 1/x \) in the above analysis (with \( \gamma = 2 \), as required for a hierarchical gas distribution). This is just what is expected for random star formation, where the largest stellar mass increases with cloud mass simply because of random sampling from the IMF. That is, the largest stellar mass satisfies \( \int_0^{M_c} n(M) dM = 1 \), as discussed above, and this gives a constant of proportionality \( n_0 = x M_L^c \) in the expression \( n(M) = n_0 M^{-1-x} \). Thus the total number of stars scales with \( M_L^c \). If the efficiency is about constant with cloud mass (and the smallest mass star is much less massive than \( M_L \)), then this total number scales about with the cloud mass, giving \( M_L^c \propto M_c \), or \( \alpha = 1/x \).

There are other explanations for the steep IMF in the extreme field. Star forming regions are typically much larger than 30 pc, often extending in a coherent fashion up to several hundred parsecs (Efremov 1995), so the 30 pc limit in the definition of the extreme field may allow some normal cluster, association, or star-complex members to be included. In that case, the steep slope in the outer regions of a cluster may occur for the same reasons as the shallow slope in the inner region, i.e., segregation of the most massive stars towards the center.
In summary, the general form of the IMF is probably invariant among clouds of different masses, giving a maximum stellar mass that increases with cloud mass as the power $1/x = 1/1.35$ as a result of random sampling (i.e., more massive clouds sample further out into the high mass tail of the IMF). This explains the similarity between the composite IMF of whole galaxies and the IMFs of individual clusters. However, in the extreme field, where conditions like ambient pressure are very different than in OB associations, star-forming clouds could be more quickly and easily destroyed by ionization from stars, and in this case, the maximum stellar mass could increase much more slowly with cloud mass, as the power $1/4$ instead of $1/1.35$. As a result, the composite IMF can be much steeper than the individual IMFs in each cluster. Alternatively, the extreme field IMF could be sampling only the low mass members of an extended cluster whose other members are more centrally located.

2.3. An IMF that is Independent of Cluster Density

One of the most startling aspects of the observed IMF is that it is virtually invariant from cluster to cluster, aside from likely statistical fluctuations (Elmegreen 1999a), and this relative invariance spans a range of a factor of 200 in cluster density (Hunter et al. 1997; Massey & Hunter 1998) and a factor of 10 in metal abundance (Freedman 1985; Massey, Johnson & DeGioia-Eastwood 1995).

The density independence means that the IMF is probably not the result of protostar, star, or clump interactions. If it were, then dense regions, which should have more of these interactions, would differ from low density regions, where there are few or no interactions. The IMF is also not likely to result from accretion of cloud material during stellar orbital motion. Stars in denser regions orbit in a shorter time and have more gas to accrete. Neither is the IMF or any part of it from the coalescence of low mass stars or protostars.

This lack of a density dependence for individual stars in the IMF contrasts with the situation for binary stars and disks. The binary fraction is smaller in denser regions, and protostellar disks are smaller too (see review in Elmegreen et al. 1999). The protostellar binary fraction is lower in both the Trapezium cluster (Petr et al. 1998) and the Pleiades cluster (Bouvier et al. 1997) than it is in the Tau-Aur region, by a factor of $\sim 3$. Also, the peak in the separation distribution for binaries is smaller (90 AU) in the part of the Sco-Cen association that contains early type stars than it is (215 AU) in the part of the Sco-Cen association that contains no early type stars (Brandner & K"ohler 1998).

The cluster environment also apparently affects disks. Mundy et al. (1995) suggested that massive disks are relatively rare in the Trapezium cluster, and N"urnberger et al. (1997) found that protostellar disk mass decreases with stellar age in the Lupus young cluster, but not in the Tau-Aur region, which is less dense. When massive stars are present, as in the Trapezium cluster, uv radiation can photoionize the neighboring disks (Johnstone et al. 1998).

These observations make sense in terms of the relative interaction rates for stars, binaries, and disks (Elmegreen et al. 1999). The size of a typical embedded cluster is $\sim 0.1$ pc, and the number of stars is several hundred. This makes the stellar density on the order of $10^3 - 10^4$ stars pc$^{-3}$. For example, in the Trapezium cluster, the stellar density is $\sim 5000$ stars pc$^{-3}$ (Prosser et al. 1994) or higher (McCaughrean & Stauffer 1994), and in Mon R2 it is $\sim 9000$ stars pc$^{-3}$ (Carpenter et al. 1997). A stellar density of $10^3$ M$_\odot$ pc$^{-3}$ corresponds to an H$_2$ density of $\sim 10^4$ cm$^{-3}$. Molecular cores with densities of $10^5$ cm$^{-3}$ or higher (e.g., Lada 1992) can easily make clusters this dense, because star formation efficiencies are typically 10%-40% (e.g., see Greene & Young 1992; Megeath et al. 1996; Tapia et al. 1996).
The density of \( n_{\text{star}} = 10^3 \) stars pc\(^{-3} \) in a cloud core of size \( R_{\text{core}} \sim 0.2 \) pc implies that objects with this density will collide with each other in one crossing time if their cross section is \( \sigma \sim (n_{\text{star}} R_{\text{core}})^{-1} \sim 0.005 \) pc\(^2 \), which corresponds to a physical size of \( \sim 6500 \) AU. This is the size of protostellar disks and long-period binary stars. Thus disks and binaries should be affected by interactions in the cluster environment, but not individual stars or the IMF.

### 2.4. The Flattening at Low Mass: a Characteristic Mass for Stars

The IMF flattens on a log – log plot at stellar masses of around and below 0.3 M\(_\odot\). Table 3 summarizes the observations. The mass at which this flattening occurs is observed to vary a bit from region to region, particularly in clusters (i.e., the mass at the peak in NGC 6231 is 2.5 M\(_\odot\), much higher than normal; Sung, Bessell, & Lee 1998), but such variations could be the result of mass segregation in the sense that high mass stars are often concentrated towards cluster cores (see Sect. 3). There is even evidence for a turnover in the IMF at masses less than 0.3 M\(_\odot\) for several regions, but this is uncertain because the stars at the low mass end are usually close to the limit of detection.

The importance of the IMF flattening is that this is the only characteristic scale known for the star formation process. Molecular clouds and their pieces have a power law mass distribution from sub-stellar masses to the masses of clouds as big as the galactic scale height. There is essentially no characteristic scale for clouds. The mass distributions for open clusters and perhaps even primordial globular clusters are power laws too, with about the same slope as for clouds (Elmegreen & Efremov 1997; see review in Elmegreen et al. 1999). The rest of the IMF is a power law too. But the IMF does have a characteristic scale at the low mass end, where it flattens at about 0.3 M\(_\odot\).

The existence of such a characteristic mass is an important clue to the mechanism of star formation. For example, we know now that the characteristic mass is not the Jeans mass at an optical depth of unity, as formerly suggested, because this mass is too small, \( \sim 10^{-3} \) M\(_\odot\) (e.g., Rees 1976). The two most promising suggestions for the origin of the characteristic mass are: (1) self-limitation of accretion by protostellar winds triggered at the deuterium-burning mass (Nakano, Hasegawa, & Norman 1995; Adams & Fatuzzo 1996), and (2) the inability of a cloud piece smaller than the thermal Jeans mass to become self-gravitating and collapse to a star, given the temperature and pressure of a molecular cloud core (Larson 1992; Elmegreen 1997).

The first of these limits would seem to be relatively independent of environment, while the second should scale with \( T^2/P^{1/2} \) for cloud temperature \( T \) and cloud-core pressure \( P \). Both values are about the same locally, where \( T \sim 10^\text{K} \) and \( P \sim 10^\text{6} \) k\(\text{B} \) cm\(^{-3} \), and since \( T^2 \) and \( P^{1/2} \) tend to vary together with galactocentric radius and star formation activity (Elmegreen 1997, 1999b), the two masses should remain the same in most normal regions.

To check the theoretical predictions, we should look for places where \( T^2/P^{1/2} \) deviates a lot from its local value. If the mass at the peak of the IMF, or where the IMF flattens, varies from region to region along with the quantity \( T^2/P^{1/2} \), then the second model would be preferred; if the peak mass does not, then the first model is better. For example, Larson (1998) suggested that the peak in the IMF was shifted towards higher masses in the early Universe, in order to account for the G dwarf problem, the large heavy element abundance and high temperature in galactic cluster gas, and the high luminosities of distant galaxies. Variations like this would be more easily explained by an IMF model that depends on the thermal Jeans mass.
### Table 3: Observations at Low Mass

| Flat, $x \sim 0 - 0.5$: |  |
|-------------------------|-----------------------------|
| **Clusters**            | Reid 1987; Kroupa, Tout & Gilmore 1990, 1991 Hubbard, Burrows, & Lunine 1990; Zuckerman & Becklin 1992; Kroupa, Gilmore, & Tout 1992 Tinney, Mould, & Reid 1992; Tinney 1993 Laughlin & Bodenheimer 1993; Comeron et al. 1993; Jarrett, Dickman & Herbst 1994; Paresce, de Marchi, & Romaniello 1995; Strom, Kepner, & Strom 1995; Pound & Blitz 1995; Williams, Rieke & Stauffer 1995; Williams, et al. 1995 Kroupa 1995a; Comeron, Rieke, & Rieke 1996 Meusinger et al. 1996; Macintosh, et al. 1996 Festin 1997; Hillenbrand 1997; Luhman & Rieke 1998; Reid 1998 |
| **Binary star mass ratios** | Tout 1991 |
| **Lack of Brown Dwarfs** | Zuckerman & Becklin 1992; Reid & Gazis 1997 Reid 1998 |
| **Globular clusters, halo, and bulge stars** | de Marchi & Paresce 1997; Chabrier & Mera 1997 Holtzman et al. 1998 |
| **Rise at lower mass** | Tinney 1993; Mera et al. 1996; Zapatero Osorio et al. 1997 |
| **Turnover at lower mass** | Reid & Gazis 1997; Hillenbrand 1997; Reid 1998 Sung et al. 1998; Nota et al. 1998; King et al. 1998 Gould, et al. 1997 |
The thermal Jeans mass, which contains the combination of parameters $T^2/P^{1/2}$, is approximately constant in normal regions of star formation. This is because the numerator in this expression is approximately proportional to the cooling rate per unit mass in molecular clouds (which scales about as $T^2 - T^3$ – see Neufeld, Lepp, & Melnick 1995), and the denominator is approximately proportional to the heating rate per unit mass from starlight and cosmic rays in typically active disks. The starlight and cosmic ray intensities scale with the background column density of stars, and the pressure in the midplane of the disk scales with the square of this column density. Thus the square root of pressure goes with the column density of background stars. As long as heating equals cooling and the mass-to-light ratio in a galactic disk is about constant, and as long as the factor by which star-forming clouds have a higher pressure than the ambient pressure is about constant, the thermal Jeans mass is about the same in all dense cloud regions. If the mass-to-light ratio goes down, then the thermal Jeans mass can go up. Perhaps this occurs in starburst regions. Conversely, if the mass-to-light ratio is abnormally high, then the thermal Jeans mass can go down.

An example of the latter situation might arise in the inner regions of M31. There the molecular cloud heating rate is low and the cloud temperature is close to 3K, instead of the usual 10K (Allen et al. 1995; Loinard & Allen 1998). These clouds also exist in the part of the disk where the stellar column density is high in old stars, so the interstellar pressure is not particularly low. As a result, the thermal Jeans mass can be lower in ultracold clouds than in normal clouds, possibly as low as $0.01 \, M_\odot$ instead of $0.3 \, M_\odot$ (Elmegreen 1999c). For this reason, a significant population of Brown Dwarf stars might be present in ultracold molecular clouds. If they are found, then the model based on the thermal Jeans mass would be preferred over the model based on the deuterium burning limit.

The thermal Jeans model is preferred also if a reasonably high fraction, say $> 10\%$, of all the material in a collapsing cloud piece gets into a star. This leaves a lot of mass for wind expulsion and disk erosion, but it also implies that the star mass depends somewhat on the mass of the cloud piece in which it forms. In that case, wind-limitations to the stellar mass would not be very important, causing only a factor of 2–10 variation in the ratio of star mass to cloud mass. Most of the mass variation along the IMF, which spans a factor of $\sim 10^3$ in mass, would then have to come from something else, and the mass of the pre-stellar cloud piece is a likely place.

Another observation that could help distinguish between possible origins for the characteristic stellar mass is the discovery of powerful pre-main sequence winds from extremely low-mass Brown Dwarfs, i.e., stars too small to ignite even deuterium. If pre-main sequence contraction energy alone is enough to start a wind, then deuterium burning would not be relevant to the limitation of stellar mass.

There is some evidence already that the mass function for dense cloud cores containing about a solar mass is similar to the IMF (Motte, Andrè, & Neri 1998; Testi & Sargent 1998). This is the type of observation that could clarify the origin of the characteristic mass for star formation.

### 2.5. Top-Heavy IMFs in Starburst Regions

There has been considerable discussion about a shift in the IMF towards proportionally more high mass stars in starburst regions, although many of the initial reports are now being questioned. The original motivation for this idea was the observation that the luminosity of the starburst was so high, given the total mass from the rotation curve, that there could not be a normal proportion of high and low mass stars but only an excess of high mass stars. Now, more detailed modeling, and in the case of M82, a lower extinction correction (Devereux 1989, Satyapal et al. 1997), makes the stellar luminosity seem about right for the
mass. A summary of these observations is in Table 4. In addition, a top-heavy IMF would produce too much oxygen in proportion to other elements (Wang & Silk 1993), and the aging population of stars would be too red (Charlot et al. 1993).

Considering the basic form of the IMF, which is a power law with a lower cutoff or flattening at some characteristic mass, one can easily envision variations that lead to top-heavy IMFs as a result of an upward shift in the characteristic mass. A predicted downward shift leading to an excess of Brown Dwarfs was mentioned in a previous section. The upward shift would come in the same way, but from an increase rather than a decrease in the value of $T^2/P^{1/2}$. It is more difficult to envision a top heavy IMF that results from a decrease in the slope of the power law part, because the very existence of a power law suggests a scale-free process, which means that it is essentially free of dependence on physical parameters. Power law mass distributions often result from geometric (e.g., fractal) or self-regulatory (e.g., equilibrium coalescence) effects instead.

The IMF model in Elmegreen (1997), in which the power law part comes from a weighted selection of clump pieces in a hierarchically structured cloud and the low mass cutoff comes from the thermal Jeans mass, gets a simple shift in the whole IMF towards higher mass, with a constant slope in the power-law part, as $T^2/P^{-1/2}$ increases. A computer simulation showing this result was in that paper.

An amazing thing about the IMF is that the characteristic mass at the low end, where the flattening occurs, appears to be nearly constant from region to region. As discussed above, this may simply reflect equilibrium thermal conditions with varying $T$ and $P$ but constant $T^2/P^{1/2}$, or it may reflect a constant wind-limited mass at the threshold of deuterium burning. The upward shift for starbursts, if real, provides a good test for the models. It is easier to increase $T^2/P^{1/2}$ in warm regions at slightly elevated pressures than to affect the deuterium burning limit, which would seem to be independent of environment. Thus the exact form of the IMF in starburst conditions is extremely important for the models. In this respect, the reported slight upward shift in the characteristic mass for the 30 Dor cluster in the LMC (Nota et al. 1998) is noteworthy. This is the closest starburst-like region, and therefore the most promising for providing a firm observation of the IMF from direct star counting. Unfortunately, this cluster could suffer from mass segregation effects as in other clusters, in which case the upward shift would appear only in the nuclear region.

The discussion about starburst IMFs begs the question of whether there is an upper limit to the mass of a star that can form. No such upper limit has been found yet. That is, the upper limit in any particular region just keeps increasing as the total stellar mass increases, as expected for random star formation (see theory in Elmegreen 1983, 1997, and observations in Massey & Hunter 1998). Yet there would seem to come a time where this stellar mass increase would have to stop. After all, if we scale the $1/x$ power law

| Shifts to high mass | Rieke et al. 1980; Kronberg, Biermann & Schwab 1985; Wright et al. 1988; Telesco 1988; Doane & Matthews 1993; Doyon, Joseph, & Wright 1994 | Rieke et al. 1993; Smith et al. 1995, 1998 | Shier et al. 1996 |
|--------------------|-------------------------------------------------------------------------------------------------|---------------------------------|------------------|
| Normal, $x \sim 1 - 1.5$ | Devereux 1989; Schaerer 1996; Satyapal et al. 1997; Calzetti 1997 |                                                                 |
relation between the maximum star mass and total star mass to all of the young stellar mass in the galaxy, with an age less than the $\sim 2$ million year lifetime of a massive star, then the total young stellar mass is $\sim 10^7 M_\odot$ and the expected maximum stellar mass is

$$M_{\text{max}} \sim 50 \left( \frac{10^7 M_\odot}{10^{4.5} M_\odot} \right)^{1/1.35} M_\odot \sim 3600 M_\odot.$$  \hspace{1cm} (1)

Here we have normalized this power law relation to the maximum mass ($\sim 50 M_\odot$) and total mass ($\sim 10^{4.5} M_\odot$) in the Orion OB association. The result is very inaccurate, of course, but the lack of Galactic stars containing several thousand solar masses suggests that there is an upper mass cutoff.

An alternative explanation for the lack of thousand-$M_\odot$ stars is that each star-forming region is independent, so the total stellar mass used in the above equation is the maximum stellar mass in the largest region of star formation, rather than the maximum for all regions in the Galaxy. In that case, the numerator in the above expression should be $\sim 10^{5.5} M_\odot$ for the largest star complexes forming in $10^7 M_\odot$ spiral arm clouds, and $M_{\text{max}} \sim 200 M_\odot$, which may be possible a few places in the Galaxy. If such stars are found, then there may be no maximum mass based on physical principles, only one based on sampling statistics.

### 3. Peculiarities with Massive Stars: central concentration and late appearances in clusters, and a preference for massive clouds

Most massive stars form in giant molecular clouds in OB associations, and not in small clouds like Taurus, which seem to contain only low mass stars (Larson 1982; Myers & Fuller 1993). Massive stars also form relatively late in the evolution of a star cluster, after many low mass stars have already formed (Herbig 1962; Iben & Talbot 1966; Herbst & Miller 1982; Adams, Strom & Strom 1983).

There have been several attempts to explain the correspondence between extreme star mass and cloud mass as a consequence of different mechanisms for star formation or different physical conditions in large and small clouds (Larson 1992; Khersonsky 1997), however observations like this are expected from random star formation alone (Elmegreen 1983; Walter & Boyd 1991; Massey & Hunter 1998), so the need for any special theory is not compelling.

If stars form randomly in all clouds, with stellar masses in the proportion given by a normal IMF, then statistical effects will make the massive stars, which are relatively rare, more likely to appear after there are $100 - 1000 M_\odot$ of other stars already (Elmegreen 1983; Schroeder & Comins 1988). This means that massive stars tend to show up only in massive clouds, and when they do, they are relatively late compared to the more common low mass stars. Simulations of this effect are in Elmegreen (1999a). Note that the average time of appearance of a star with a particular mass is still independent of that mass in this statistical interpretation, so if there is a systematic bias toward a late appearance of high mass stars, then some physical process for this would be required. Stahler (1995) suggested, however, that even the proposed examples of such bias probably have other interpretations, so the entire effect could be just statistical.

Another peculiar observation of massive stars is that they tend to appear near the centers of star clusters, surrounded by the lower mass stars (see reviews in Elmegreen et al. 1999; Testi, Palla, & Natta 1998). This peculiar distribution for massive stars has been observed using color gradients in 12 clusters (Sagar & Bhatt 1989), and from the steepening of the IMF with radius in several clusters (Pandey, Mahra, & Sagar 1992), including Tr 14 (Vazquez et al. 1996), the Trapezium in Orion (Jones & Walker 1988; Hillenbrand 1997; Hillenbrand & Hartmann 1998), and, in the LMC, NGC 2157 (Fischer et al. 1998), SL...
666, and NGC 2098 (Kontizas et al. 1998).

The usual explanation for this effect is that massive stars sink to the center of a cluster during dynamical relaxation, but several clusters seem too young for this to have happened (Bonnell & Davies 1998), including Orion Trapezium (Hillenbrand & Hartmann 1998). In that case, the high mass stars had to have been born near the cluster centers, perhaps because the most massive clumps were closer to the center at the time the massive stars were born in them. There are other explanations too. The stars near the center could have accreted gas faster and ended up more massive (Larson 1978, 1982; Zinnecker 1982; Bonnell et al. 1997); they or their predecessor clumps could have coalesced more (Larson 1990; Zinnecker et al. 1993; Stahler, Palla, & Ho 1999; Bonnell, Bate, & Zinnecker 1998), or the most massive stars and clumps forming anywhere could have migrated to the center faster because of a greater gas drag (Larson 1990, 1991; Gorti & Bhatt 1995, 1996; Saiyadpour, Deiss, & Kegel 1997). A problem with most of these models is that they are inconsistent with the observation that the IMF is nearly independent of cluster density (Sect. 2.3). Another model without this problem suggests that the central location of the most massive stars is from the central location of the most massive cloud pieces, which is expected for a hierarchical cloud (Elmegreen 1999a).

### 4. Evolution of the IMF

The discussion above suggests that the IMF has been somewhat constant in time and place, except possibly for an upward shift in the mass at the IMF peak for starburst regions (Sect. 2.5). There was also a suggestion that the IMF was shifted towards higher mass in the early Universe (Larson 1998), although very old stars and old intergalactic gas seem to show evidence for a normal IMF (Table 1).

The direct observations of normal star-forming regions point to a *universal* IMF, with deviations perhaps only from statistical fluctuations in small samples and from mass segregation in clusters. The observations of regions with extremely low star-forming activity suggest a shift towards lower masses, either with a steeper IMF (as observed by Massey et al. 1995) or, possibly, a downward shift in the peak (as predicted by Elmegreen 1999c). Observations of regions with extremely high star-forming activity suggest an analogous shift towards higher masses, possibly as a result of an upward shift in the peak.

If the mass at the peak of the IMF can really change with star formation activity, possibly as a result of changes in the ratio $T^2/P^{1/2}$, which is in the thermal Jeans mass, then there are several important implications. First, the ratio $T^2/P^{1/2}$ depends roughly on the light-to-mass ratio in a galaxy disk, because the numerator is proportional to the cooling, and therefore heating rate in molecular clouds, and the denominator is proportional to the local mass column density (Sect. 2.4). This means that if the light-to-mass ratio is high, the peak in the IMF can shift towards higher masses, and vice versa. Now it follows from the Schmidt law, which has a star formation rate proportional to average density to some power greater than unity (e.g., Kennicutt 1998), that the gas consumption rate in a star-forming region increases with higher density, and the luminosity-to-mass ratio for luminous young stars increases too. If the peak in the IMF increases along with the higher L/M ratio, then we get the interesting result that the IMF peak increases with the gas consumption rate (Elmegreen 1999a). We might also have a higher efficiency of star formation in such a region, because of the generally greater self-binding of clouds in high pressure or high velocity-dispersion gas (Elmegreen, Kaufman, & Thomasson 1993). This circumstance could then explain why some starburst regions have all three of these peculiarities at the same time (see review in Telesco 1988).
What happened in the early Universe is more difficult to assess. Although the temperature was higher from the cosmic microwave background, in proportion to $(1+z)$, the average density of the Universe was higher too, in the proportion $(1+z)^3$, and the pressure, which is a product of density and temperature, was higher by $(1+z)^4$. Thus the ratio of $T^2/P^{1/2}$ in the thermal Jeans mass was independent of $z$. However, $T$ and $P$ variations in newly forming galaxies should dominate these average $z$ variations, and the thermal Jeans mass could have gone either way. If the earliest stars formed in cool high-pressure shocks, then perhaps the thermal Jeans mass was lower than it is today, producing Brown Dwarfs. On the other hand, if the temperature was higher because of an inability to cool without metals, then the Jeans mass could have been higher. The observation of nearly normal abundances in Ly α forest lines and the intercluster medium (Table 1) suggest that this characteristic mass probably did not vary much in the early Universe.

5. Summary

The IMF is a power law at intermediate to high mass, with a flattening on a log-log plot at low mass. The mass at which the flattening occurs is the only characteristic mass that has been clearly observed for star formation, and is therefore an important indicator of physical processes that depend on scale. Examples might be the thermal Jeans mass or the minimum mass for deuterium burning and stellar winds, both of which have about the right value. Methods to distinguish between these two possibilities were discussed in Section 2.4. The power-law part of the IMF may not indicate specific physical processes, but be more of a remnant from the observed scale-free geometry of pre-stellar clouds. Random sampling models for such geometries reproduce essentially all of the IMF properties with very little sensitivity to free parameters. In this case, much of the physics of the star formation process may be unrecoverable from the power-law part of the IMF alone.

The lack of any obvious dependence of the IMF on cluster density places strong constraints on the physical processes that might be involved (Sect. 2.3). The steep IMF in the extreme field (Sect. 2.2), as well as other systematic variations in the IMF, such as the concentration of massive stars in cluster cores (Sect. 3) and the shift in the IMF towards higher masses in starburst regions (Sect. 2.5), all suggest specific physical differences in the properties of star-forming regions and perhaps in the mechanisms of star formation too. Differences in the IMF from place to place and time to time may eventually tell us more about star formation than any single IMF, which may have washed out any such details in the averaging process.

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