Die Geburt der Sterne

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Star Formation

Fig. 3.— At Spherical slice of the gas density inside the Jeans volume at \( t = 1 \) pc, with \( 128 \) cells per Jeans length \( \lambda_\text{v} \) and \( 5 \text{ km s}^{-1} \) velocity gradient. Magnetic field lines showing a highly tangled and twisted magnetic field structure typical of the small-scale dynamo; yellow: \( 0.5 \text{ m G} \), red: \( 1 \text{ m G} \). Four randomly chosen individual field lines. The green one in particular is extremely tangled close to the center of the Jeans volume. Contours of the vorticity modulus \( |\nabla \times \mathbf{v}| \) show elongated filamentary structure typical of subsonic turbulence. From Frisch 1995.

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agenda

• star formation theory
  - phenomenology
  - historic remarks
  - our current understanding and its limitations

• application
  - the stellar mass function at birth (IMF)
phenomenology
- star formation sets in very early after the big bang
- stars always form in galaxies and protogalaxies
- we cannot see the first generation of stars, but maybe the second one
- correlation between stellar birth and large-scale dynamics
- spiral arms
- tidal perturbation from neighboring galaxy
- HI gas more extended
- H2 and SF well correlated

Galaxies from THINGS and HERACLES survey
(images from Frank Bigiel, ZAH/ITA)
distribution of molecular gas in the Milky Way as traced by CO emission
- stars form in molecular clouds
- stars form in clusters
- stars form on ~ dynamical time
- (protostellar) feedback is very important
• strong feedback: UV radiation from Θ1C Orionis affects star formation on all cluster scales
Eventually, clusters like the ONC (1 Myr) will evolve into clusters like the Pleiades (100 Myr).
theoretical approach
- **density**
  - density of ISM: few particles per cm$^3$
  - density of molecular cloud: few 100 particles per cm$^3$
  - density of Sun: 1.4 g/cm$^3$

- **spatial scale**
  - size of molecular cloud: few 10s of pc
  - size of young cluster: ~ 1 pc
  - size of Sun: 1.4 $\times$ $10^{10}$ cm
• density
  - density of ISM: few particles per cm$^3$
  - density of molecular cloud: few 100 particles per cm$^3$
  - density of Sun: 1.4 g/cm$^3$

• spatial scale
  - size of molecular cloud: few 10s of pc
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  - size of Sun: 1.4 x 10$^{10}$ cm
• density
  - density of ISM: few particles per cm
  - density of molecular cloud: few 100 particles per cm
  - density of Sun: 1.4 g
• spatial scale
  - size of molecular cloud: few 10s of pc
  - size of young cluster: ~ 1
  - size of Sun: 1.4 x 10

decrease in spatial scale / increase in density
- **contracting force**
  - only force that can do this compression is GRAVITY

- **opposing forces**
  - there are several processes that can oppose gravity
  - GAS PRESSURE
  - TURBULENCE
  - MAGNETIC FIELDS
  - RADIATION PRESSURE
- contracting force
  - only force that can do this compression is **GRAVITY**

- opposing forces
  - there are several processes that can oppose gravity
  - **GAS PRESSURE**
  - **TURBULENCE**
  - **MAGNETIC FIELDS**
  - **RADIATION PRESSURE**

Modern star formation theory is based on the complex interplay between all these processes.
Jeans (1902): Interplay between self-gravity and thermal pressure

- stability of homogeneous spherical density enhancements against gravitational collapse
- dispersion relation:

\[
\omega^2 = c_s^2 k^2 - 4\pi G \rho_0
\]

- instability when \( \omega^2 < 0 \)
- minimal mass:

\[
M_J = \frac{1}{6} \pi^{-5/2} G^{-3/2} \rho_0^{-1/2} c_s^3 \propto \rho_0^{-1/2} T^{3/2}
\]
von Weizsäcker (1943, 1951) and Chandrasekhar (1951): concept of MICROTURBULENCE

- BASIC ASSUMPTION: separation of scales between dynamics and turbulence
  \[ \ell_{\text{turb}} \ll \ell_{\text{dyn}} \]

- then turbulent velocity dispersion contributes to effective soundspeed:
  \[ c_c^2 \mapsto c_c^2 + \sigma_{\text{rms}}^2 \]

- \[ \rightarrow \] Larger effective Jeans masses \[ \rightarrow \] more stability

- BUT:
  1. turbulence depends on \( k \): \[ \sigma_{\text{rms}}^2(k) \]
  2. supersonic turbulence \[ \rightarrow \sigma_{\text{rms}}^2(k) \gg c_s^2 \] usually
problemas de teoría dinámica de los inicios

- Nuevamente, los nubarrones moleculares son altamente Jeans-inestables, pero no forman estrellas a un alto ritmo y con alta eficiencia (conundrum de Zuckerman & Evans 1974) (la eficiencia global de formación de estrellas en nubarrones moleculares es ~5%).
  \[ \rightarrow \text{something prevents large-scale collapse}\]

- A lo largo de los años 1990, los nubarrones moleculares se pensaban como entidades quasiequilibrio.

- Los nubarrones moleculares son magnéticos.
Magnetic star formation

- **Mestel & Spitzer (1956):** Magnetic fields can prevent collapse!!!
  - Critical mass for gravitational collapse in presence of B-field
    \[ M_{cr} = \frac{5^{3/2} B^3}{48\pi^2 \frac{G^{3/2}}{\rho^2}} \]
  - Critical mass-to-flux ratio
    (Mouschovias & Spitzer 1976)
    \[ \left[ \frac{M}{\Phi} \right]_{cr} = \frac{\zeta}{3\pi} \left[ \frac{5}{G} \right]^{1/2} \]
  - Ambipolar diffusion can initiate collapse

Lyman Spitzer, Jr., 1914 - 1997
“standard theory” of star formation

• BASIC ASSUMPTION: Stars form from magnetically highly subcritical cores

• Ambipolar diffusion slowly increases \( \frac{M}{\Phi} : \tau_{AD} \approx 10\tau_{ff} \)

• Once \( \frac{M}{\Phi} > \frac{M}{\Phi}_{\text{crit}} \) : dynamical collapse of SIS
  • Shu (1977) collapse solution
  • \( \frac{dM}{dt} = 0.975 \frac{c_s^3}{G} = \text{const.} \)

• Was (in principle) only intended for isolated, low-mass stars
problems of “standard theory”

- Observed B-fields are weak, at most marginally critical (Crutcher 1999, Bourke et al. 2001)
- Structure of prestellar cores (e.g. Bacman et al. 2000, Alves et al. 2001)
- Strongly time varying dM/dt (e.g. Hendriksen et al. 1997, André et al. 2000)
- More extended infall motions than predicted by the standard model (Williams & Myers 2000, Myers et al. 2000)
- Most stars form as binaries (e.g. Lada 2006)
- As many prestellar cores as protostellar cores in SF regions (e.g. André et al 2002)
- Molecular cloud clumps are chemically young (Bergin & Langer 1997, Pratap et al 1997, Aikawa et al 2001)
- Stellar age distribution small ($\tau_{ff} \ll \tau_{AD}$) (Ballesteros-Paredes et al. 1999, Elmegreen 2000, Hartmann 2001)
- Strong theoretical criticism of the SIS as starting condition for gravitational collapse (e.g. Whitworth et al 1996, Nakano 1998, as summarized in Klessen & Mac Low 2004)
- Standard AD-dominated theory is incompatible with observations (Crutcher et al. 2009, 2010ab, Bertram et al. 2011)

(see e.g. Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125-194)
gravoturbulent star formation

- **BASIC ASSUMPTION:**
  - star formation is controlled by interplay between supersonic turbulence and self-gravity

- turbulence plays a *dual role*:
  - on *large scales* it provides support
  - on *small scales* it can trigger collapse

- some predictions:
  - dynamical star formation timescale $\tau_{ff}$
  - high binary fraction
  - complex spatial structure of embedded star clusters
  - and many more . . .

Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125-194
McKee & Ostriker, 2007, ARAA, 45, 565
turbulent cascade in the ISM

- scale-free behavior of turbulence in the range $\frac{L}{\eta_k} \approx \text{Re}^{3/4}$
- slope between $-5/3 \ldots -2$
- energy "flows" from large to small scales, where it turns into heat

energy source & scale
*NOT known*
(supernovae, winds, spiral density waves?)

dissipation scale not known
(ambipolar diffusion, molecular diffusion?)
turbulent cascade in the ISM

- Molecular clouds:
  \( \sigma_{\text{rms}} \approx \text{several km/s} \)
  \( M_{\text{rms}} > 10 \)
  \( L > 10 \text{ pc} \)

- Dense protostellar cores:
  \( \sigma_{\text{rms}} \ll 1 \text{ km/s} \)
  \( M_{\text{rms}} \leq 1 \)
  \( L \approx 0.1 \text{ pc} \)

Energy source & scale **NOT known**
(supernovae, winds, spiral density waves?)

Dissipation scale not known
(ambipolar diffusion, molecular diffusion?)
turbulence creates a hierarchy of clumps
as turbulence decays locally, contraction sets in
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while region contracts, individual clumps collapse to form stars
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individual clumps collapse to form stars
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in dense clusters, clumps may merge while collapsing --> then contain multiple protostars

$$\alpha = \frac{E_{\text{kin}}}{|E_{\text{pot}}|} < 1$$
in *dense clusters*, clumps may merge while collapsing
--> then contain multiple protostars
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--> then contain multiple protostars
in *dense clusters*, competitive mass growth becomes important
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in dense clusters, $N$-body effects influence mass growth
low-mass objects may become ejected --> accretion stops
feedback terminates star formation
result: *star cluster*, possibly with H\(\text{II}\) region
current status

• stars form from the complex interplay of self-gravity and a large number of competing processes (such as turbulence, B-field, feedback, thermal pressure)

• the relative importance of these processes depends on the environment
  - prestellar cores --> thermal pressure is important
  - molecular clouds --> turbulence dominates

  } (Larson’s relation: $\sigma \propto L^{1/2}$)

  - massive star forming regions (NGC602): radiative feedback is important
  - small clusters (Taurus): evolution maybe dominated by external turbulence

• star formation is regulated by various feedback processes

• star formation is closely linked to global galactic dynamics (KS relation)

Star formation is intrinsically a multi-scale and multi-physics problem, where it is difficult to single out individual processes. Simple theoretical approaches usually fail.
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selected open questions

- what processes determine the initial mass function (IMF) of stars?
- what are the initial conditions for star cluster formation?
  how does cloud structure translate into cluster structure?
- how do molecular clouds form and evolve?
- what drives turbulence?
- what triggers / regulates star formation on galactic scales?
- how does star formation depend on metallicity?
  how do the first stars form?
- star formation in extreme environments (galactic center, starburst, etc.),
  how does it differ from a more “normal” mode?
selected open questions

- what processes determine the initial mass function (IMF) of stars?
- what are the initial conditions for star cluster formation? how does cloud structure translate into cluster structure?
- how do molecular clouds form and evolve?
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- how does star formation depend on metallicity? how do the first stars form?
- star formation in extreme environments (galactic center, starburst, etc.), how does it differ from a more “normal” mode?
stellar mass function
stars seem to follow a universal mass function at birth --> IMF

(Kroupa 2002)

Orion, NGC 3603, 30 Doradus
(Zinnecker & Yorke 2007)
stellar masses

- distribution of stellar masses depends on
  - turbulent initial conditions
    --> mass spectrum of prestellar cloud cores
  - collapse and interaction of prestellar cores
    --> accretion and $N$-body effects
  - thermodynamic properties of gas
    --> balance between heating and cooling
    --> EOS (determines which cores go into collapse)
  - (proto) stellar feedback terminates star formation
    ionizing radiation, bipolar outflows, winds, SN

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(Kroupa 2002)
nearby molecular clouds

Pipe
Ophiuchus
Taurus
Perseus

study more closely

scales to same scale

Orion

10 pc
(from A. Goodman)
example: model of Orion cloud

„model“ of Orion cloud:
15,000,000 SPH particles,
$10^4 \, M_{\text{sun}}$ in 10 pc, mass resolution
$0,02 \, M_{\text{sun}}$, forms $\sim 2,500$
„stars“ (sink particles)

isothermal EOS, top bound, bottom unbound

has clustered as well as distributed „star“ formation

efficiency varies from 1% to 20%

develops full IMF
(distribution of sink particle masses)

(Bonnell, Smith, Clark, & Bate 2010, MNRAS, 410, 2339)
example: model of Orion cloud

- 15.000.000 SPH particles
- $10^4 \, M_{\odot}$ in 10 pc, mass resolution 0.02 $M_{\odot}$, forms ~2.500 "stars" (sink particles)

- MASSIVE STARS: form early in high-density gas clumps (cluster center), high accretion rates, maintained for a long time
- LOW-MASS STARS: form later as gas falls into potential well, high relative velocities, little subsequent accretion
dynamics of nascent star cluster

in dense clusters protostellar interaction may be come important!

Trajectories of protostars in a nascent dense cluster created by gravoturbulent fragmentation
Mass accretion rates vary with time and are strongly influenced by the cluster environment.

distribution of stellar masses depends on
- turbulent initial conditions
  --> mass spectrum of prestellar cloud cores
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- thermodynamic properties of gas
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  --> EOS (determines which cores go into collapse)
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(Kroupa 2002)
stellar mass function

• distribution of stellar masses depends on
  - turbulent initial conditions
    --> mass spectrum of prestellar cloud cores
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  - thermodynamic properties of gas
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  - (proto) stellar feedback terminates star formation
    ionizing radiation, bipolar outflows, winds, SN, etc.

application to early star formation

(Kroupa 2002)
thermodynamics & fragmentation

degree of fragmentation depends on EOS!

polytropic EOS: $p \propto \rho^\gamma$

$\gamma < 1$: dense cluster of low-mass stars
$\gamma > 1$: isolated high-mass stars

(see Li et al. 2003; also Kawachi & Hanawa 1998, Larson 2003)
dependency on EOS

for $\gamma < 1$ fragmentation is enhanced $\Rightarrow$ cluster of low-mass stars
for $\gamma > 1$ it is suppressed $\Rightarrow$ isolated massive stars

(from Li, Klessen, & Mac Low 2003, ApJ, 592, 975)
how does that work?

(1) \[ p \propto \rho^\gamma \Rightarrow \rho \propto p^{1/\gamma} \]

(2) \[ M_{\text{jeans}} \propto \gamma^{3/2} \rho^{(3\gamma-4)/2} \]

- \( \gamma < 1 \):
  - large density excursion for given pressure
  - \( \langle M_{\text{jeans}} \rangle \) becomes small
  - number of fluctuations with \( M > M_{\text{jeans}} \) is large

- \( \gamma > 1 \):
  - small density excursion for given pressure
  - \( \langle M_{\text{jeans}} \rangle \) is large
  - only few and massive clumps exceed \( M_{\text{jeans}} \).
EOS as function of metallicity

For those above $10^{-10}$ number density, which is calculated by one-zone models. The dashed lines indicate the constant Jeans masses. For those above $10^{-3}$ number density, which is calculated by one-zone models. In Figure 2 of O05, where similar plots for the evolution of specific heat at the center, one-zone models are presented. In Figure 3.4, we focus on the temperature evolution at each metallicity. The contribution to the pressure evolution is quite similar to that calculated by the one-zone models. Below $10^{-3}$ number density, the evolution of the hydrogen gas is assumed to be fully atomic (molecular) in drawing those lines.

The gas is optically thin at $Z=0$ and $Z=0$. The lines are so small that the temperature evolution is almost identical to that of a metal-free gas. The gas is then described as having evolved from metal-free to metal-rich. Metallicity effects contribute comparably to $H_2$ formation (Equation (19)).

With their recombination proceeding, the $H_2$ catalyzed by a small amount of remaining electrons. With the increasing number density, the recombination channel is quenched. Another molecular species in the metal-free gas is $H_2O$. The heating is owing to the compression of the gas. The effective ratio of specific heat remains below $4 \times 10^{-3}$ orders of magnitude. The effective ratio of specific heat is an important index to examine the variation of pressure in response to the density variation.

Let us summarize here the formation processes of $H_2$ in metal-free cases with metals. Below $10^{-10}$ number density, the hydrogen is converted to the molecular form via this continuum absorption. Another molecular species in the metal-free case is $H_2O$. The heating is owing to the compression of the gas. The effective ratio of specific heat remains below $4 \times 10^{-3}$ orders of magnitude.
EOS as function of metallicity

The evolution of H as a function of metallicity is presented in Figure 1. (A color version of this figure is available in the online journal.)

In this section, we review thermal evolution of the cloud core in prestellar collapse for various metallicities. This is calculated by one-zone models with the effective ratio of specific heat at the center, \( \gamma \), being an important index to examine the temperature evolution at each metallicity. The contribution of the curve in Figure 1 is shown for those cases. Note that the temperature evolution is nearly isothermal with temperature difference, \( \Delta T \), equal to the gradient, \( \gamma \), for constant molecular weight. The effective ratio of specific heat gives \( \gamma \) for constant molecular weight.

Let us summarize here the formation processes of H in a prestellar core. With a linear density n \( \propto \lambda \), the temperature T(K) becomes

\[ T = n \times 10,000 \]

The effective ratio of specific heat is an important index to examine the temperature evolution. For example, the clouds easily fragment as long as \( n \gtrsim 10^5 \) cm\(^{-3}\), justifying the one-zone treatment for the metal-free gas is once ionized (Uehara & Inutsuka 2008). In our case, however, it only occurs for \( n \approx 10^4 \) cm\(^{-3}\), while the other lines become optically thick. The effective ratio of specific heat gives \( \gamma \) for constant molecular weight.

The evolution of H is formed by the H\(_2\) catalyzed by a small amount of remaining electrons. With the three-body reaction (Equation (10)), the effective ratio of specific heat gives \( \gamma \) for constant molecular weight.

\[ \gamma = \frac{\partial \ln \rho}{\partial \ln \rho} \]

After this plateau, the H\(_2\) abundance begins to increase again at \( n \approx 10^3 \) cm\(^{-3}\), where the H\(_2\) lines dominate. For the cooling, the H\(_2\) catalyzed by the optical depth, \( \tau \), of the H\(_2\) lines.

\[ \tau = \frac{n \times 10^4}{M_m} \]

With gradual increase of temperature, the balance of chemical processes is maintained, as shown in Figure 2. The amount of formed H\(_2\) is smaller than the amount of formed H by a small factor whereas density increases by many orders of magnitude. The evolution is nearly isothermal with temperature difference, \( \Delta T \), equal to the gradient, \( \gamma \), for constant molecular weight.

\[ \gamma = \frac{\partial \ln \rho}{\partial \ln \rho} \]

The result is justified by the one-zone treatment for a small amount of remaining electrons. With gradual increase of temperature, the balance of chemical processes is maintained, as shown in Figure 2. The amount of formed H\(_2\) is smaller than the amount of formed H by a small factor whereas density increases by many orders of magnitude. The evolution is nearly isothermal with temperature difference, \( \Delta T \), equal to the gradient, \( \gamma \), for constant molecular weight.
Figure 1. Evolution of temperatures in prestellar cloud cores with metallicities. This is calculated by one-zone models for different metallicities. This should be compared with Figure 2 of O05, where similar plots for the metal-free one except for a slight offset at highest densities are so small that the temperature evolution is almost identical to that calculated by the one-zone treatment for the metal-free gas, HD, is known to play an important role in cooling associated with the three-body reaction (Equation (10)). With gradual increase of temperature, the balance of chemical equilibrium between the H$_2$ and ions becomes more important until the H$_2$ dissociates and the amount of formed H by the H$_2$+H$^+$ recombination proceeding, the H$^+$ catalyzed by a small amount of remaining electrons. With increasing temperature, the H$^+$ fraction grows in the range between $10^{-4}$ and $10^{-3}$ of the total gas density.

After this plateau, the H$_2$ equilibrium is formed by the dissociative recombination of H$^+$. For the cooling, the H$^+$ collision-induced line emission contributes comparably to H$_2$-line emission and the amount of formed H$_2$ quickly increases as the temperature decreases. For the higher density, the H$_2$ equilibrium is formed by the H$^+$ dissociative recombination of H$^+$. For the cooling, the H$^+$ collision-induced line emission contributes comparably to H$_2$-line emission and the amount of formed H$_2$ quickly increases as the temperature decreases. For example, the clouds easily fragment as long as the dynamical response of self-gravitating clouds to thermal evolution is quite similar to that calculated by the one-zone treatment for the metal-free one except for a slight offset at highest densities.

The temperature evolution at each metallicity. The contribution of the curve in Figure 1 to the cooling and heating rates by individual processes are so small that the temperature evolution is almost identical to that calculated by the one-zone treatment for the metal-free gas. We then describe the effects of metallicity later in Section 4. PRESTELLAR COLLAPSE.

The temperature evolution at the center of cloud cores during the prestellar collapse for various metallicities. This is calculated by one-zone models for different metallicities. This should be compared with Figure 2 of O05, where similar plots for the metal-free one except for a slight offset at highest densities are so small that the temperature evolution is almost identical to that calculated by the one-zone treatment for the metal-free gas, HD, is known to play an important role in cooling associated with the three-body reaction (Equation (10)). With gradual increase of temperature, the balance of chemical equilibrium between the H$_2$ and ions becomes more important until the H$_2$ dissociates and the amount of formed H by the H$_2$+H$^+$ recombination proceeding, the H$^+$ catalyzed by a small amount of remaining electrons. With increasing temperature, the H$^+$ fraction grows in the range between $10^{-4}$ and $10^{-3}$ of the total gas density.

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EOS as function of metallicity

The dashed lines in Figure 1, while fragmentation is strongly prohibited for cases with metals. Below this value, the dynamical collapse is halted as the pressure is formed by the H\(_2\) channel:

\[ \text{H}_2 \rightarrow \text{H}^+ \text{H} \text{+} \text{H} \ (11) \]

Next, let us see the cooling and heating processes associated with the three-body reaction (Equation (6)), the gas is assumed to be fully atomic (molecular) in drawing those lines. The evolution of H\(_2\) and the amount of formed H\(_2\) saturates at \(\sim 10^6 \text{M}_\odot\) until very high density. For constant molecular weight, the effective ratio of specific heat remains below \(4\) orders of magnitudes. The effective ratio of specific heat is an important index to examine the variation of pressure in response to the density variation, while the dynamical response of self-gravitating clouds to thermal evolution. For example, the clouds easily fragment as long as their recombination proceeding, the H\(_2\) recombination:

\[ \text{H}_2 \rightarrow \text{H}^+ \text{H} \text{+} \text{H} \ (11) \]

With gradual increase of temperature, the balance of chemical processes controlling the temperature evolution at each metallicity. The contribution to the cooling and heating rates by individual processes are quite similar to that calculated by the one-zone models are presented. In Figure 2 of O05, where similar plots for the overall evolution is quite similar to that calculated by the one-zone models. The evolution of H\(_2\) gas is assumed to be fully atomic (molecular) in drawing those lines. The evolution of H\(_2\) number density log (n\(_H\) cm\(^{-3}\)) is presented in Figure Z/Z. OMUKAI, HOSOKAWA, & YOSHIDA V ol. 722 (A color version of this figure is available in the online journal.)
present-day star formation

\( \tau = 1 \) (Larson 1985, Larson 2005)

\( \gamma = 1.1 \)

\( \gamma = 0.7 \)

(Larson 1985, Larson 2005)
IMF in nearby molecular clouds

with $\rho_{\text{crit}} \approx 2.5 \times 10^5 \text{ cm}^{-3}$

at SFE $\approx 50\%$

$M_{\text{low}} / M_{\text{total}} = 40\%$

need appropriate EOS in order to get low mass IMF right

EOS as function of metallicity

Indicate the constant Jeans masses. For those above $10^{17}$ number density, which is calculated by one-zone models. The dashed lines play a crucial role in the thermal evolution. The evolution of $\frac{\Delta T}{\Delta n}$ is shown for those cases. Note that small disagreements, are so small that the temperature evolution is almost identical to case. In the case of metallicity $[M/H] < -3$, which give $s(t) = H_{\text{eff}} - \frac{\Delta H}{\Delta t} = 0$, justifying the one-zone treatment for overall evolution is quite similar to that calculated by the one-zone models are presented. In Figure 2 of O05, where similar plots for the prestellar cores as a function of the number density. The effective ratio of specific heat is an important index to examine of the curve in Figure 1. In the case of metallicity $[M/H] = 0$, metallicity effects dominate. For the cooling, the $H_2$ molecules contribute.

Figure 1. (A color version of this figure is available in the online journal.)

Figure 2. (A color version of this figure is available in the online journal.)

(Omukai et al. 2005, 2010)
**EOS as function of metalliclicity**

- slope of EOS in the density range $5 \text{ cm}^{-3} \leq n \leq 16 \text{ cm}^{-3}$ is $\gamma \approx 1.06$.
- with non-zero angular momentum, disk forms.
- disk is unstable against fragmentation at high density

(Onukai et al. 2005, 2010)
detailed look at accretion disk around first star

successive zoom-in calculation from cosmological initial conditions (using SPH and new grid-code AREPO)

Redshift: $z = 21$
Boxsize: 150/h kpc (comoving)
Slice Width: 10/h kpc (comoving)


A detailed look at the accretion disk around the first star is provided. The time evolution of the accretion disk is shown through successive zoom-in calculations from cosmological initial conditions using SPH and the new grid-code AREPO. The image illustrates the number density maps for a slice through the high-density region, with different metallicities ($Z$). Figure 3 shows a sequence of zooms in the density structure in the gas immediately before the formation of the first protostar.

Figure 4 presents the enclosed gas mass divided by the Bonnor-Ebert mass versus radius for different metallicities. The values were calculated at the time just before the first sink was formed, with the center taken to be the position of the densest SPH particle. This indicates a more flat mass distribution. The mean fragmentation time and the mean accretion time explain the difference in the sink particle mass distribution in Figure 6.

For $Z \leq 10^{-4} Z_{\odot}$, the fragmentation time is always higher than the accretion time, indicating that the sink particles will accrete faster than they can be generated, resulting in a more flat mass distribution. When the fragmentation time is higher than the accretion time (for $Z = 10^{-4} Z_{\odot}$), the gas rather fragments than moves to the center and is accreted. As a consequence, more mass goes into the low-mass objects, in comparison to the high-mass ones. This behavior agrees well with the predictions from before fragmentation started, shown in Figure 7.

Another property of the star-forming cloud that is observed to vary in the calculations is the mass spatial distribution. The dependence of the enclosed gas and sink mass on the distance from the sink center of mass, for the different $Z$, is shown in Figure 9. The $Z = 0$ case has almost all the sink particle mass in $r < 8$ AU. The gas density for this case is also higher in this region, when compared to the other metallicities, showing that the gas and sink particles mass density follow each other.

In the $Z = 0$ simulation, there is $\sim 80\%$ of the mass in sinks within 8 AU from the center of mass. And for the other cases, this happens for radius $\sim 30$ AU. For radius bigger than 150 AU, the gas becomes the most massive component, for all $Z$.

This more concentrated gas and sink mass towards the center happens probably because for the $Z = 0$ case, the gas had higher temperatures in the central region. And so there was less influence by turbulent and rotational motions, which were successive zoom-in calculation from cosmological initial conditions (using SPH and new grid-code AREPO).

The figure shows the following:

- The accretion disk around the first star is illustrated through successive zoom-in calculations from cosmological initial conditions using SPH and the new grid-code AREPO.
- Figure 3 shows a sequence of zooms in the density structure in the gas immediately before the formation of the first protostar.
- Figure 4 presents the enclosed gas mass divided by the Bonnor-Ebert mass versus radius for different metallicities ($Z$). The values were calculated at the time just before the first sink was formed, with the center taken to be the position of the densest SPH particle. This indicates a more flat mass distribution.
- Figure 9 shows the mass spatial distribution, with the dependence of the enclosed gas and sink mass on the distance from the sink center of mass.

References:
Figure 1: Density evolution in a 120 AU region around the first protostar, showing the build-up of the protostellar disk and its eventual fragmentation. We also see ‘wakes’ in the low-density regions, produced by the previous passage of the spiral arms.
Figure 2: Radial profiles of the disk's physical properties, centered on the first protostellar core to form. The quantities are mass-weighted and taken from a slice through the midplane of the disk. In the lower right-hand plot we show the radial distribution of the disk's Toomre parameter, $Q = \frac{c_s \kappa}{\pi G \Sigma}$, where $c_s$ is the sound speed and $\kappa$ is the epicyclic frequency. Because our disk is Keplerian, we adopted the standard simplification, and replaced $\kappa$ with the orbital frequency.

The molecular fraction is defined as the number density of hydrogen molecules $(n_{H_2})$, divided by the number density of hydrogen nuclei $(n)$, such that fully molecular gas has a value of $0.5$.

\[ Q < 1 \] indicates instability.
similar study with very different numerical method (AREPO)

one out of five halos

Most recent calculations: 
**fully sink-less simulations, following the disk build-up over ~10 years**
(resolving the protostars - first cores - down to $10^5$ km $\sim$ 0.01 $R_\odot$)
expected mass spectrum

- **expected IMF is flat** and covers a wide range of masses

- implications
  - because slope > -2, most mass is in massive objects as predicted by most previous calculations
  - most high-mass Pop III stars should be in *binary systems* --> source of *high-redshift gamma-ray bursts*
  - because of ejection, some *low-mass objects* (< 0.8 $M_\odot$) might have *survived* until today and could potentially be found in the Milky Way

- consistent with abundance patterns found in second generation stars
The metallicities of extremely metal-poor stars in the halo are consistent with the yields of core-collapse supernovae, i.e. progenitor stars with 20 - 40 $M_\odot$

(e.g. Tominaga et al. 2007, Izutani et al. 2009, Joggerst et al. 2009, 2010)
Star formation is intrinsically a multi-scale and multi-physics problem. Many different processes need to be considered simultaneously.
Star formation is intrinsically a multi-scale and multi-physics problem, where it is difficult to single out individual processes.

- stars form from the competing processes (such as pressure)
- thermodynamic properties: the star formation process
- detailed studies require the physical and chemical processes
- star formation is understood
- primordial star formation: star formation
... people in the star formation group at Heidelberg University:

Christian Baczynski, Erik Bertram, Frank Bigiel, Andre Bubel, Diane Cormier, Volker Gaibler, Simon Glover, Dimitrious Gouliermis, Tilman Hartwig, Juan Ibanez, Christoph Klein, Lukas Konstandin, Mei Sasaki, Jennifer Schober, Rahul Shetty, Rowan Smith, László Szűcs

... former group members:

Robi Banerjee, Ingo Berentzen, Paul Clark, Christoph Federrath, Philipp Girichidis, Thomas Greif, Milica Micic, Thomas Peters, Dominik Schleicher, Stefan Schmeja, Sharanya Sur, . . .

... many collaborators abroad!
thanks