Star Formation

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thanks to...

... people in the group in Heidelberg:

Christian Baczynski, Erik Bertram, Frank Bigiel, Rachel Chicharro, Roxana Chira, Paul Clark, Gustavo Dopcke, Jayanta Dutta, Volker Gaibler, Simon Glover, Lukas Konstandin, Faviola Molina, Mei Sasaki, Jennifer Schober, Rahul Shetty, Rowan Smith, László Szűcs, Svitlana Zhukovska

... former group members:

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

... many collaborators abroad!
I try to cover the field as broadly as possible, however, there will clearly be a bias towards my personal interests and many examples will be from my own work.
Schedule

- Formation of molecular clouds
- Origin and statistical characteristics of ISM turbulence and introduction to star (cluster) formation
- Stellar initial mass function
Literature

Books


Bodenheimer, P. 2012, “Principles of Star Formation” (Springer Verlag)

Draine, B. 2011, “Physics of the Interstellar and Intergalactic Medium” (Princeton Series in Astrophysics)
Books


Bodenheimer, P. 2012, “Principles of Star Formation” (Springer Verlag)

Draine, B. 2011, “Physics of the Interstellar and Intergalactic Medium” (Princeton Series in Astrophysics)
Review Articles

Mac Low, M.-M., Klessen, R.S., 2004, "The control of star formation by supersonic turbulence", Rev. Mod. Phys., 76, 125 - 194


inventory of Galactic disc component

- **stellar disc**
  - thin disc (80% of mass): stars of all ages 0-12Gyr
  - thick disc (5% of mass): older stars with lower metallicity

- **interstellar medium (ISM)**
  - gas (15% of mass): hot, warm, and cool component (atomic and molecular)
  - dust (<1% of gas mass): well mixed with the cool gas
  - cosmic rays: relativistic particles
  - magnetic fields: frozen to the gas (field lines are co-moving with the gas); energy density comparable to the kinetic energy of gas
### Interstellar Matter: ISM

Abundances, scaled to 1,000,000 H atoms

<table>
<thead>
<tr>
<th>element</th>
<th>atomic number</th>
<th>abundance</th>
</tr>
</thead>
<tbody>
<tr>
<td>hydrogen</td>
<td>H</td>
<td>1,000,000</td>
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<tr>
<td>deuterium</td>
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<td>16</td>
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<tr>
<td>helium</td>
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<tr>
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<tr>
<td>silicium</td>
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<td>iron</td>
<td>Fe</td>
<td>34</td>
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<tr>
<td>nickel</td>
<td>Ni</td>
<td>2</td>
</tr>
</tbody>
</table>

Hydrogen is by far the most abundant element (more than 90% in number).
Because hydrogen is the dominating element, the classification scheme is based on its chemical state:

- **ionized atomic hydrogen** $\text{H}^+$
- **neutral atomic hydrogen** $\text{H}$
- **molecular hydrogen** $\text{H}_2$

Different regions consist of almost 100% of the appropriate phase, the transition regions between HII, H and $\text{H}_2$ are very thin.

Star formation always takes place in dense and cold molecular clouds.
Phases of the ISM

A\textsubscript{V} denotes the extinction, the attenuation of radiation due to absorption (mostly on dust grains)
Life-cycle of ISM

- Giant Molecular Cloud Complexes
  - Gravitational Collapse M>MJ
  - Disruption during star formation
  - Merging

- Molecular Clouds
  - Merging & cooling

- Stars
  - Supernovae
  - Planetary Nebulae
  - Breakup of old shells around plan. neb. & SNR

- Warm HI
  - Heating and/or cooling by conduction
  - Cooling after compression

- Cold HI
  - Merging & cooling

- HII
  - Cooling
  - HII regions
  - Coronal Gas
    - Heating by supernovae
    - Cooling

The cycle continues with cold HI, warm HI, and stars as key components.
Galactic Longitude

Molecular Ring

\( M_{\text{tot}} (H_2) \sim 2 \times 10^9 \, M_\odot \)

\( M_{\text{tot}} (\text{H}^1) \sim 6 \times 10^9 \, M_\odot \)

Data from Thomas Dame, CfA Harvard
Predicted $N(H_2)$ ("IRAS - H I")

Observed Velocity-Integrated CO

Data from Thomas Dame, CfA Harvard
Orion Nebula Cluster (ESO, VLT, M. McCaughrean)

data from T. Dame (CfA Harvard)
• stars form in molecular clouds
• stars form in clusters
• stars form on ~ dynamical time
• (protostellar) feedback is very important
Ionizing radiation from central star Θ 1C Orionis

- strong feedback: UV radiation from Θ 1C Orionis affects star formation on all cluster scales

Trapezium stars in the center of the ONC (HST, Johnstone et al. 1998)
eventually, clusters like the ONC (1 Myr) will evolve into clusters like the Pleiades (100 Myr)
nearby molecular clouds

scales to same scale

- Orion
- Ophiuchus
- Perseus
- Pipe
- Taurus

study more closely

(from A. Goodman)
what drives ISM turbulence?

- seems to be driven on large scales, little difference between star-forming and non-SF clouds
  ---> rules out internal sources

- proposals in the literature
  - supernovae
  - expanding HII regions / stellar winds / outflows
  - spiral density waves
  - magneto-rotational instability
  - more recent idea: accretion onto disk
what drives ISM turbulence?

some energetic arguments...

energy decay by turbulent dissipation:

\[ \dot{e} \approx -\frac{1}{2} \rho v_{\text{rms}}^3 / L_d \]

\[ = -(3 \times 10^{-27} \text{ erg cm}^{-3} \text{ s}^{-1}) \left( \frac{n}{1 \text{ cm}^{-3}} \right) \]

\[ \times \left( \frac{v_{\text{rms}}}{10 \text{ km s}^{-1}} \right)^3 \left( \frac{L_d}{100 \text{ pc}} \right)^{-1}, \]

decay timescale:

\[ \tau_d = e / \dot{e} = L_d / v_{\text{rms}} \]

\[ = (9.8 \text{ Myr}) \left( \frac{L_d}{100 \text{ pc}} \right) \left( \frac{v_{\text{rms}}}{10 \text{ km s}^{-1}} \right)^{-1}, \]

(from Mac Low & Klessen, 2004)
what drives ISM turbulence?

magneto-rotational instability:

$$\dot{e} = \left( 3 \times 10^{-29} \text{ erg cm}^{-3} \text{ s}^{-1} \right) \left( \frac{B}{3 \mu \text{G}} \right)^2 \left( \frac{\Omega}{(220 \text{ Myr})^{-1}} \right).$$

gravitational instability (spiral waves)

$$\dot{e} \approx G \left( \frac{\Sigma_g}{H} \right)^2 \lambda^2 \Omega$$

$$\approx \left( 4 \times 10^{-29} \text{ erg cm}^{-3} \text{ s}^{-1} \right)$$

$$\times \left( \frac{\Sigma_g}{10M_\odot \text{ pc}^{-2}} \right)^2 \left( \frac{H}{100 \text{ pc}} \right)^{-2}$$

$$\times \left( \frac{\lambda}{100 \text{ pc}} \right)^2 \left( \frac{\Omega}{(220 \text{ Myr})^{-1}} \right),$$

(from Mac Low & Klessen, 2004)

(from Piotek & Ostriker 2005)

(from Walter et al. 2008)
what drives ISM turbulence?

protostellar outflows

\[ \dot{e} = \frac{1}{2} f_w \eta_w \frac{\Sigma_* H}{v_w^2} \]

\[ \approx \left(2 \times 10^{-28} \text{ erg cm}^{-3} \text{ s}^{-1}\right) \left(\frac{H}{200 \text{ pc}}\right)^{-1} \left(\frac{f_w}{0.4}\right) \]

\[ \times \left(\frac{v_w}{200 \text{ km s}^{-1}}\right) \left(\frac{v_{\text{rms}}}{10 \text{ km s}^{-1}}\right) \]

\[ \times \left(\frac{\Sigma_*}{4.5 \times 10^{-9} M_\odot \text{ pc}^{-2} \text{ yr}^{-1}}\right), \]

(expanding HII regions)

\[ \dot{e} = \frac{\langle \delta p \rangle N(>1) v_i}{V_{t_f}} \]

\[ = \left(3 \times 10^{-30} \text{ erg s}^{-3}\right)^{-1} \]

\[ \times \left(\frac{N_H}{1.5 \times 10^{22} \text{ cm}^{-2}}\right)^{-3/14} \left(\frac{M_{cl}}{10^6 M_\odot}\right)^{1/14} \]

\[ \times \left(\frac{\langle M_* \rangle}{440 M_\odot}\right) \left(\frac{N(>1)}{650}\right) \left(\frac{v_i}{10 \text{ km s}^{-1}}\right) \]

\[ \times \left(\frac{H_c}{100 \text{ pc}}\right)^{-1} \left(\frac{R_{sf}}{15 \text{ kpc}}\right)^{-2} \left(\frac{t_i}{18.5 \text{ Myr}}\right)^{-1}, \]


(note: different numbers by Matzner 2002)

(from Mac Low & Klessen, 2004)
what drives ISM turbulence?

supernovae

\[ \dot{e} = \frac{\sigma_{SN} \eta_{SN} E_{SN}}{\pi R_{sf}^2 H_c} \]
\[= (3 \times 10^{-26} \text{ erg s}^{-1} \text{ cm}^{-3}) \left( \frac{\eta_{SN}}{0.1} \right) \left( \frac{\sigma_{SN}}{1 \text{ SNu}} \right) \]
\[\times \left( \frac{H_c}{100 \text{ pc}} \right)^{-1} \left( \frac{R_{sf}}{15 \text{ kpc}} \right)^{-2} \left( \frac{E_{SN}}{10^{51} \text{ erg}} \right). \]

in star-forming parts of the disk, clearly SN provide enough energy to compensate for the decay of ISM turbulence.

*BUT*: what is outside the disk?

(from Mac Low & Klessen, 2004)
accretion driven turbulence

yet another thought:

- astrophysical objects *form by accretion* of ambient material
- the *kinetic energy* associated with this process is a key agent *driving internal turbulence*.
- this works on *ALL* scales:
  - galaxies
  - molecular clouds
  - protostellar accretion disks

Klessen & Hennebelle (2010, A&A)
turbulence decays on a crossing time
\[ \tau_d \approx \frac{L_d}{\sigma} \]

energy decay rate
\[ \dot{E}_{\text{decay}} \approx \frac{E}{\tau_d} = -\frac{1}{2} \frac{M \sigma^3}{L_d} \]

kinetic energy of infalling material
\[ \dot{E}_{\text{in}} = \frac{1}{2} M_{\text{in}} v_{\text{in}}^2 \]

can both values match, modulo some efficiency?
\[ \epsilon = \left| \frac{\dot{E}_{\text{decay}}}{\dot{E}_{\text{in}}} \right| \]

some estimates from convergent flow studies

Klessen & Hennebelle (2010)
application to galaxies

underlying assumption
  - galaxy is in steady state
    ---> accretion rate equals star formation rate
  - what is the required efficiency for the method to work?

study Milky Way and 11 THINGS
  - excellent observational data in HI:
    velocity dispersion, column density, rotation curve
11 THINGS galaxies
some further thoughts

- Method works for Milky Way type galaxies:
  - Required efficiencies are \(~1\%\) only!
- Relevant for outer disks (extended HI disks)
  - There are not other sources of turbulence (certainly not stellar sources, maybe MRI)
- Works well for molecular clouds
  - Example clouds in the LMC (Fukui et al.)
- Potentially interesting for TTS
  - Model reproduces $dM/dt - M$ relation (e.g. Natta et al. 2006, Muzerolle et al. 2005, Muhanty et al. 2005, Calvet et al. 2004, etc.)
Molecular Cloud Formation
molecular cloud formation

- star formation on galactic scales
  → requires understanding of formation of molecular clouds

- questions
  - *where* and *when* do molecular clouds form?
  - *what* are their properties?
  - *how* do stars form in their interior?
  - global correlations? → *Schmidt law*
Molecular clouds form at stagnation points of large-scale convergent flows, mostly triggered by global (or external) perturbations.

(Deul & van der Hulst 1987, Blitz et al. 2004)
The correlation with large-scale perturbations in warm atomic ISM are caused by thermal/gravitational instability and/or supersonic turbulence. Some fluctuations are dense enough to form $H_2$ within "reasonable time" → molecular cloud.

External perturbations (i.e. potential changes) increase likelihood.
star formation on \textit{global} scales

mass weighted $\rho$-pdf, each shifted by $\Delta \log N = 1$

(from Klessen, 2001; also Gazol et al. 2005, Krumholz & McKee 2005, Glover & Mac Low 2007ab)

probability distribution function of the density ($\rho$-pdf)

\textit{varying rms Mach numbers:}

$M_1 > M_2 > M_3 > M_4 > 0$
star formation on global scales

H$_2$ formation rate:

$$
\tau_{H_2} \approx \frac{1.5 \text{Gyr}}{n_H / 1 \text{cm}^{-3}}
$$

for $n_H \geq 100 \text{ cm}^{-3}$, H$_2$ forms within 10 Myr, this is about the lifetime of typical MC’s.

in turbulent gas, the H$_2$ fraction can become very high on short timescale

(for models with coupling between cloud dynamics and time-dependent chemistry, see Glover & Mac Low 2007a,b)

mass weighted $\rho$-pdf, each shifted by $\Delta \log N = 1$

(rate from Hollenback, Werner, & Salpeter 1971)
star formation on global scales

BUT: it doesn’t work
(at least not so easy):

Chemistry has a memory effect!

H2 forms more quickly in high-density regions as it gets destroyed in low-density parts.

mass weighted $\rho$-pdf, each shifted by $\Delta \log N = 1$

(rate from Hollenback, Werner, & Salpeter 1971)

(for models with coupling between cloud dynamics and time-dependent chemistry, see Glover & Mac Low 2007a,b)
SFR estimates from the PDF

log density PDF:

\[ p_s(s) = \frac{1}{\sqrt{2\pi}\sigma_s^2} \exp\left(-\frac{(s - s_0)^2}{2\sigma_s^2}\right) \]

\[ s \equiv \ln\left(\frac{\rho}{\rho_0}\right) \text{, log density, normalized to the mean} \]

relation between mean density and turbulent Mach number \( M \) and magnetic field strength \( \beta \):

\[ s_0 = -\frac{1}{2} \sigma_s^2 \]

\[ \sigma_s^2 = \ln \left(1 + b^2 M^2 \frac{\beta}{\beta + 1}\right) \]

\[ \sigma_s^2 = \ln \left(1 + b^2 M^2 \frac{2M_A^2}{M^2 + 2M_A^2}\right) \]

---

SFR estimates from the PDF

star formation rate (Msun/yr) in terms of the SF efficiency per free-fall time \( \text{SFR}_{ff} \)

\[
\text{SFR} \equiv \frac{M_c}{t_{ff}(\rho_0)} \text{SFR}_{ff}.
\]

\[
\text{SFR}_{ff} = \frac{\epsilon}{\phi_t} \int_{s_{\text{crit}}}^{\infty} \frac{t_{ff}(\rho_0)}{t_{ff}(\rho)} \frac{\rho}{\rho_0} p(s) \, ds.
\]

\[
t_{ff}(\rho) \equiv \left( \frac{3\pi}{32G\rho} \right)^{1/2}
\]


### SFR estimates from the PDF

comparison and extension of existing models

<table>
<thead>
<tr>
<th>Analytic Model</th>
<th>Freefall-time Factor</th>
<th>Critical Density $\rho_{\text{crit}}/\rho_0 = \exp(s_{\text{crit}})$</th>
<th>SFR$_{\text{ff}}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>KM</td>
<td>1</td>
<td>$(\pi^2/5) \phi_x^2 \times \alpha_{\text{vir}} M^2 (1 + \beta^{-1})^{-1}$</td>
<td>$\frac{\epsilon}{(2\phi_t)} { 1 + \text{erf} \left[ \left( \sigma_s^2 - 2s_{\text{crit}} \right) / \left( 8\sigma_s^2 \right)^{1/2} } }$</td>
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<tr>
<td>PN</td>
<td>$t_{\text{ff}}(\rho_0)/t_{\text{ff}}(\rho_{\text{crit}})$</td>
<td>$(0.067) \theta^{-2} \times \alpha_{\text{vir}} M^2 f(\beta)$</td>
<td>$\frac{\epsilon}{(2\phi_t)} { 1 + \text{erf} \left[ \left( \sigma_s^2 - 2s_{\text{crit}} \right) / \left( 8\sigma_s^2 \right)^{1/2} } \exp \left[ (1/2)s_{\text{crit}} \right]$</td>
</tr>
<tr>
<td>HC</td>
<td>$t_{\text{ff}}(\rho_0)/t_{\text{ff}}(\rho)$</td>
<td>$(\pi^2/5) y_{\text{cut}}^{-2} \times \alpha_{\text{vir}} M^{-2} (1 + \beta^{-1}) + \tilde{\rho}_{\text{crit,turb}}$</td>
<td>$e/(2\phi_t) { 1 + \text{erf} \left[ \left( \sigma_s^2 - s_{\text{crit}} \right) / (2\sigma_s^2)^{1/2} } } \exp \left[ (3/8)\sigma_s^2 \right]$</td>
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<td>multi-ff KM</td>
<td>$t_{\text{ff}}(\rho_0)/t_{\text{ff}}(\rho)$</td>
<td>$(\pi^2/5) \phi_x^2 \times \alpha_{\text{vir}} M^2 (1 + \beta^{-1})^{-1}$</td>
<td>$\frac{\epsilon}{(2\phi_t)} { 1 + \text{erf} \left[ \left( \sigma_s^2 - s_{\text{crit}} \right) / (2\sigma_s^2)^{1/2} } } \exp \left[ (3/8)\sigma_s^2 \right]$</td>
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<td>multi-ff HC</td>
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</tr>
</tbody>
</table>
SFR estimates from the PDF

comparison between analytic models and numerical simulations

Federrath & Klessen 2012
SFR estimates from the PDF

comparison between numerical simulations and observations

Federrath & Klessen 2012
modeling galactic SF

SPH calculations of self-gravitating disks of stars and (isothermal) gas in dark-matter potential, sink particles measure local collapse --> star formation

We find a correlation between star formation rate and gas surface density:

\[ \Sigma_{SFR} \propto \Sigma_{gas}^{1.5} \]
observed Schmidt law

\[ \Sigma_{\text{SFR}} \propto \Sigma_{\text{gas}}^{1.5} \]

\( \alpha = 1.48 \pm 0.08 \)

in both cases:

(from Kennicutt 1998)
molecular cloud formation

(from Dobbs et al. 2008)
molecular cloud formation

(Dobbs & Bonnell 2007)
molecular cloud formation

molecular gas fraction as function of time

molecular gas fraction as function of density

(Dobbs et al. 2008)
molecular cloud formation

molecular gas fraction of fluid element as function of time

molecular gas fraction as function of density

$R = 7.1 \text{kpc}$

(Dobbs et al. 2008)
modeling chemistry
(movie from Christoph Federrath)
6 ray approximation to external radiation field

- AMR MHD (B = 2 muG)
- stochastic forcing (Ornstein-Uhlenbeck)
- self-gravity
- time-dependent chemistry
- cooling & heating processes
  --> thermodynamics done right!

- gives you mathematically well defined boundary conditions
  --> good for statistical studies
chemical model 0

- 32 chemical species
  - 17 in instantaneous equilibrium: $\text{H}^-, \text{H}_2^+, \text{H}_3^+, \text{CH}^+, \text{CH}_2^+, \text{OH}^+, \text{H}_2\text{O}^+, \text{H}_3\text{O}^+, \text{CO}^+, \text{HOC}^+, \text{O}^-, \text{C}^-$ and $\text{O}_2^+$
  - 19 full non-equilibrium evolution
    - $\text{e}^-, \text{H}^+, \text{H}, \text{H}_2, \text{He}, \text{He}^+, \text{C}, \text{C}^+, \text{O}, \text{O}^+, \text{OH}, \text{H}_2\text{O}, \text{CO}$,
    - $\text{C}_2, \text{O}_2, \text{HCO}^+, \text{CH}, \text{CH}_2$ and $\text{CH}_3^+$
- 218 reactions
- various heating and cooling processes

(Glover, Federrath, Mac Low, Klessen, 2010, MNRS, 404, 2)
## Chemical Model 1

### Cooling:

<table>
<thead>
<tr>
<th>Process</th>
<th>Reference(s)</th>
</tr>
</thead>
<tbody>
<tr>
<td><strong>Cooling:</strong></td>
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<tr>
<td></td>
<td>Collision rates (H) – Abrahamsson, Krems &amp; Dalgarno (2007)</td>
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<tr>
<td></td>
<td>Collision rates (H₂) – Schroder et al. (1991)</td>
</tr>
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<td>Collision rates (e⁻) – Johnson et al. (1987)</td>
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<td></td>
<td>Collision rates (H⁺) – Roueff &amp; Le Bourlot (1990)</td>
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<tr>
<td></td>
<td>Collision rates (H₂) – Flower &amp; Launay (1977)</td>
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<tr>
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<td>Collision rates (H, T &gt; 2000 K) – Keenan et al. (1986)</td>
</tr>
<tr>
<td></td>
<td>Collision rates (e⁻) – Wilson &amp; Bell (2002)</td>
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<td>Collision rates (H) – Abrahamsson, Krems &amp; Dalgarno (2007)</td>
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<td>Collision rates (H₂) – see Glover &amp; Jappsen (2007)</td>
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<td></td>
<td>Collision rates (e⁻) – Bell, Berrington &amp; Thomas (1998)</td>
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<td></td>
<td>Collision rates (H⁺) – Pequignot (1990, 1996)</td>
</tr>
<tr>
<td>H₂ rovibrational lines</td>
<td>Le Bourlot, Pinceau des Forêts &amp; Flower (1999)</td>
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<tr>
<td>CO and H₂O rovibrational lines</td>
<td>Neufeld &amp; Kaufman (1993); Neufeld, Lepp &amp; Melnick (1995)</td>
</tr>
<tr>
<td>OH rotational lines</td>
<td>Pavlovskii et al. (2002)</td>
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<tr>
<td>Gas-grain energy transfer</td>
<td>Hollenbach &amp; McKee (1989)</td>
</tr>
<tr>
<td>Recombination on grains</td>
<td>Wolfire et al. (2003)</td>
</tr>
<tr>
<td>Atomic resonance lines</td>
<td>Sutherland &amp; Dopita (1993)</td>
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<td>H collisional ionization</td>
<td>Abel et al. (1997)</td>
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<tr>
<td>H₂ collisional dissociation</td>
<td>See Table B1</td>
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<tr>
<td>Compton cooling</td>
<td>Cen (1992)</td>
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</table>

### Heating:

<table>
<thead>
<tr>
<th>Process</th>
<th>Reference(s)</th>
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<tbody>
<tr>
<td><strong>Heating:</strong></td>
<td></td>
</tr>
<tr>
<td>Photoelectric effect</td>
<td>Bakes &amp; Tielens (1994); Wolfire et al. (2003)</td>
</tr>
<tr>
<td>H₂ photodissociation</td>
<td>Black &amp; Dalgarno (1977)</td>
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<tr>
<td>UV pumping of H₂</td>
<td>Burton, Hollenbach &amp; Tielens (1990)</td>
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<td>H₂ formation on dust grains</td>
<td>Hollenbach &amp; McKee (1989)</td>
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<tr>
<td>Cosmic ray ionization</td>
<td>Goldsmith &amp; Langer (1978)</td>
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<td>No.</td>
<td>Reaction</td>
</tr>
<tr>
<td>-----</td>
<td>----------</td>
</tr>
<tr>
<td>1</td>
<td>$H + e^- \rightarrow H^+$</td>
</tr>
<tr>
<td>2</td>
<td>$H^+ \rightarrow H + H_2 + e^-$</td>
</tr>
<tr>
<td>3</td>
<td>$H + H^+ \rightarrow H_2 + \gamma$</td>
</tr>
<tr>
<td>4</td>
<td>$H + H^+ \rightarrow H_2 + H^+$</td>
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<td>$H^+ \rightarrow H + H$</td>
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<td>6</td>
<td>$H_2 + e^- \rightarrow H_2 + e^-$</td>
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<td>$H_2 + e^- \rightarrow H_2 + H$</td>
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<td>$H + e^- \rightarrow H + e^-$</td>
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**Table B1. List of collisional gas-phase reactions included in the chemical model (Glover, Federrath, Mac Low, Klessen, 2010, MNRS, 404, 2)**
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<thead>
<tr>
<th>No.</th>
<th>Reaction</th>
<th>Chemical Model 2</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>H$^+$ + H $\rightarrow$ H + H$^+$</td>
<td>$k_{11} = 2.653 \times 10^{-9} T^{0.7656}$</td>
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<tr>
<td>2</td>
<td>H$^+$ + H$^+$ $\rightarrow$ H$^+$ + H$^+$</td>
<td>$k_{12} = 9.6 \times 10^{-13}$</td>
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<td>3</td>
<td>H + H$^+$ $\rightarrow$ H + H + e$^-$</td>
<td>$k_{13} = 6.9 \times 10^{-3} T^{0.26}$</td>
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<tr>
<td>4</td>
<td>He + e$^-$ $\rightarrow$ He$^+$ + e$^-$</td>
<td>$k_{14} = 9.6 \times 10^{-7} T^{0.0} e^{-0.26}$</td>
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<td>5</td>
<td>He + e$^-$ $\rightarrow$ He$^+$ + e$^-$</td>
<td>$k_{15} = 9.6 \times 10^{-7} T^{0.0} e^{-0.26}$</td>
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<tr>
<td>6</td>
<td>He$^+$ + e$^-$ $\rightarrow$ He + e$^+$</td>
<td>$k_{16} = 1.25 \times 10^{-12} T^{-0.26}$</td>
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<td>7</td>
<td>C + e$^-$ $\rightarrow$ C$^+$ + e$^-$</td>
<td>$k_{17} = 1.24 \times 10^{-17} T^{-0.26} e^{-0.26}$</td>
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<td>8</td>
<td>O + e$^-$ $\rightarrow$ O$^+$ + e$^-$</td>
<td>$k_{18} = 6.85 \times 10^{-3} (0.0156 + u^{-0.36} T^{-2})^{-1}$</td>
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<td>9</td>
<td>O + H$^+$ $\rightarrow$ O$^+$ + H$^+$</td>
<td>$k_{19} = 2.59 \times 10^{-1} (0.0156 + u^{-0.36} T^{-2})^{-1}$</td>
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<td>10</td>
<td>H + O$^+$ $\rightarrow$ H$^+$ + O$^+$</td>
<td>$k_{20} = 1.24 \times 10^{-17} T^{-0.26} e^{-0.26}$</td>
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<td>11</td>
<td>H$^+$ + O$^+$ $\rightarrow$ H + O$^+$</td>
<td>$k_{21} = 9.6 \times 10^{-3} T^{0.26} e^{-0.26}$</td>
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<td>12</td>
<td>C + H$^+$ $\rightarrow$ C$^+$ + H$^+$</td>
<td>$k_{22} = 3.59 \times 10^{-8} (0.0156 + u^{-0.36})^{-1} e^{-0.26}$</td>
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<td>C$^+$ + H$^+$ $\rightarrow$ C$^+$ + H$^+$</td>
<td>$k_{23} = 1.24 \times 10^{-17} T^{-0.26} e^{-0.26}$</td>
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<td>C$^+$ + O$^+$ $\rightarrow$ C$^+$ + O$^+$</td>
<td>$k_{24} = 6.85 \times 10^{-3} (0.0156 + u^{-0.36} T^{-2})^{-1}$</td>
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<td>15</td>
<td>H + O$^+$ $\rightarrow$ H$^+$ + O$^+$</td>
<td>$k_{25} = 3.59 \times 10^{-8} (0.0156 + u^{-0.36})^{-1} e^{-0.26}$</td>
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<td>H$^+$ + O$^+$ $\rightarrow$ H + O$^+$</td>
<td>$k_{26} = 3.59 \times 10^{-8} (0.0156 + u^{-0.36})^{-1} e^{-0.26}$</td>
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<td>H$^+$ + O$^+$ $\rightarrow$ H + O$^+$</td>
<td>$k_{27} = 3.59 \times 10^{-8} (0.0156 + u^{-0.36})^{-1} e^{-0.26}$</td>
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<td>H$^+$ + O$^+$ $\rightarrow$ H + O$^+$</td>
<td>$k_{28} = 3.59 \times 10^{-8} (0.0156 + u^{-0.36})^{-1} e^{-0.26}$</td>
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<td>H$^+$ + O$^+$ $\rightarrow$ H + O$^+$</td>
<td>$k_{29} = 3.59 \times 10^{-8} (0.0156 + u^{-0.36})^{-1} e^{-0.26}$</td>
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<td>20</td>
<td>H$^+$ + O$^+$ $\rightarrow$ H + O$^+$</td>
<td>$k_{30} = 3.59 \times 10^{-8} (0.0156 + u^{-0.36})^{-1} e^{-0.26}$</td>
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<td>Reaction</td>
<td>Rate Constant</td>
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<td>----------------------------------------------</td>
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<td>$1. H + H$</td>
<td>$k_{1} = \frac{4 \times 10^{-13} \exp \left(-\frac{1300}{T}\right)}{\text{cm}^3 \text{ molecule}^{-2} \text{ s}^{-1}}$</td>
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<td>$2. H + \text{He}$</td>
<td>$k_{2} = 3.1 \times 10^{-13} \exp \left(-\frac{1300}{T}\right)$</td>
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<td>$3. H + \text{H}_2$</td>
<td>$k_{3} = 6.9 \times 10^{-13} \exp \left(-\frac{1300}{T}\right)$</td>
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<td>$4. \text{He} + \text{He}$</td>
<td>$k_{4} = 2.0 \times 10^{-12} \exp \left(-\frac{1300}{T}\right)$</td>
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<td>$5. \text{He} + \text{H}_2$</td>
<td>$k_{5} = 1.0 \times 10^{-10}$</td>
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<td>$6. \text{He} + \text{OH}$</td>
<td>$k_{6} = 3.0 \times 10^{-11} \exp \left(-\frac{1700}{T}\right)$</td>
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<td>$7. \text{He} + \text{O}_2$</td>
<td>$k_{7} = 2.1 \times 10^{-10} \exp \left(-\frac{9000}{T}\right)$</td>
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<td>$8. \text{He} + \text{CH}$</td>
<td>$k_{8} = 1.1 \times 10^{-12} \exp \left(-\frac{1700}{T}\right)$</td>
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<td>$9. \text{He} + \text{C}$</td>
<td>$k_{9} = 2.4 \times 10^{-11} \exp \left(-\frac{9000}{T}\right)$</td>
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<td>$10. \text{He} + \text{H}_2\text{O}$</td>
<td>$k_{10} = 7.7 \times 10^{-13} \exp \left(-\frac{1700}{T}\right)$</td>
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<td>$11. \text{He} + \text{CO}$</td>
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<td>$12. \text{He} + \text{H}_2\text{O}_2$</td>
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<td>$13. \text{He} + \text{CO}_2$</td>
<td>$k_{13} = 4.8 \times 10^{-12} \exp \left(-\frac{9000}{T}\right)$</td>
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References:
(Glover, Federbshth, Mac Low, Klessen, 2010, MNRS, 404, 4, 2)
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<td>$H^+ + H \rightarrow H_2 + H^+$</td>
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<td>$H_2 + H \rightarrow H_2 + H$</td>
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<td>$CH + CH \rightarrow CH + CH$</td>
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<tr>
<td>100</td>
<td>$O + H_2 \rightarrow H_2 + O$</td>
<td></td>
<td></td>
</tr>
</tbody>
</table>

Note: $k_{ij}$ values are given in units of $10^{-10}$ cm$^3$ molecule$^{-1}$ s$^{-1}$. 

Table B1. Chemical model 2

<table>
<thead>
<tr>
<th>No.</th>
<th>Reactions</th>
<th>$k_{ij}$</th>
<th>$T$ (K)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>$H + H \rightarrow H + H$</td>
<td>$2.4 \times 10^{-10}$</td>
<td>300</td>
</tr>
<tr>
<td>2</td>
<td>$H + H \rightarrow H + H$</td>
<td>$3.2 \times 10^{-10}$</td>
<td>500</td>
</tr>
<tr>
<td>3</td>
<td>$H + H \rightarrow H + H$</td>
<td>$4.0 \times 10^{-10}$</td>
<td>700</td>
</tr>
<tr>
<td>4</td>
<td>$H + H \rightarrow H + H$</td>
<td>$5.0 \times 10^{-10}$</td>
<td>900</td>
</tr>
</tbody>
</table>

(Glover, Federrath, Mac Low, Klessen, 2010, MNRS, 404, 2)
Table B2. List of photochemical reactions included in our chemical model

<table>
<thead>
<tr>
<th>No.</th>
<th>Reaction</th>
<th>Optically thin rate (s^{-1})</th>
<th>γ</th>
<th>Ref.</th>
</tr>
</thead>
<tbody>
<tr>
<td>166</td>
<td>H^{-} + γ → H + e^{-}</td>
<td>\text{See S.2.1}</td>
<td>0.5</td>
<td>1</td>
</tr>
<tr>
<td>167</td>
<td>H^{+} + γ → H + H^{+}</td>
<td>R_{167} = 7.4 \times 10^{-7}</td>
<td>1.9</td>
<td>2</td>
</tr>
<tr>
<td>168</td>
<td>H_{2} + γ → H + H^{+}</td>
<td>R_{168} = 5.6 \times 10^{-11}</td>
<td>2.3</td>
<td>3</td>
</tr>
<tr>
<td>169</td>
<td>H^{+} + γ → H_{2} + H</td>
<td>R_{169} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>4</td>
</tr>
<tr>
<td>170</td>
<td>H^{+} + γ → H + H_{2}</td>
<td>R_{170} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>5</td>
</tr>
<tr>
<td>171</td>
<td>CH + γ → C + H</td>
<td>R_{171} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>6</td>
</tr>
<tr>
<td>172</td>
<td>CH + γ → H + CH</td>
<td>R_{172} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>7</td>
</tr>
<tr>
<td>173</td>
<td>CH + γ → CH + H</td>
<td>R_{173} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>8</td>
</tr>
<tr>
<td>174</td>
<td>CH + γ → CH_{2} + e^{-}</td>
<td>R_{174} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>9</td>
</tr>
<tr>
<td>175</td>
<td>CH + γ → C + H_{2}</td>
<td>R_{175} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>10</td>
</tr>
<tr>
<td>176</td>
<td>CH + γ → C + H_{2}</td>
<td>R_{176} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>11</td>
</tr>
<tr>
<td>177</td>
<td>CH + γ → C + H_{2}</td>
<td>R_{177} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>12</td>
</tr>
<tr>
<td>178</td>
<td>CH_{2} + γ → CH_{2} + H</td>
<td>R_{178} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>13</td>
</tr>
<tr>
<td>179</td>
<td>CH_{2} + γ → CH_{2} + H</td>
<td>R_{179} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>14</td>
</tr>
<tr>
<td>180</td>
<td>CH_{2} + γ → CH_{2} + H</td>
<td>R_{180} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>15</td>
</tr>
<tr>
<td>181</td>
<td>C_{2} + γ → C + C</td>
<td>R_{181} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>16</td>
</tr>
<tr>
<td>182</td>
<td>O^{-} + γ → O + e^{-}</td>
<td>R_{182} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>17</td>
</tr>
<tr>
<td>183</td>
<td>OH + γ → H + OH^{-}</td>
<td>R_{183} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>18</td>
</tr>
<tr>
<td>184</td>
<td>OH + γ → OH^{-} + e^{-}</td>
<td>R_{184} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>19</td>
</tr>
<tr>
<td>185</td>
<td>OH + γ → O + H^{+}</td>
<td>R_{185} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>20</td>
</tr>
<tr>
<td>186</td>
<td>H_{2}O + γ → OH + H</td>
<td>R_{186} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>21</td>
</tr>
<tr>
<td>187</td>
<td>H_{2}O + γ → H_{2}O^{-} + e^{-}</td>
<td>R_{187} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>22</td>
</tr>
<tr>
<td>188</td>
<td>H_{2}O + γ → H_{2}O^{-} + e^{-}</td>
<td>R_{188} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>23</td>
</tr>
<tr>
<td>189</td>
<td>H_{2}O + γ → H_{2}O^{-} + e^{-}</td>
<td>R_{189} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>24</td>
</tr>
<tr>
<td>190</td>
<td>H_{2}O + γ → H_{2}O^{-} + e^{-}</td>
<td>R_{190} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>25</td>
</tr>
<tr>
<td>191</td>
<td>H_{2}O + γ → H_{2}O^{-} + e^{-}</td>
<td>R_{191} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>26</td>
</tr>
<tr>
<td>192</td>
<td>H_{2}O + γ → H_{2}O^{-} + e^{-}</td>
<td>R_{192} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>27</td>
</tr>
<tr>
<td>193</td>
<td>H_{2}O + γ → H_{2}O^{-} + e^{-}</td>
<td>R_{193} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>28</td>
</tr>
<tr>
<td>194</td>
<td>H_{2}O + γ → H_{2}O^{-} + e^{-}</td>
<td>R_{194} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>29</td>
</tr>
<tr>
<td>195</td>
<td>H_{2}O + γ → H_{2}O^{-} + e^{-}</td>
<td>R_{195} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>30</td>
</tr>
<tr>
<td>196</td>
<td>O + γ → O^{+} + e^{-}</td>
<td>R_{196} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>31</td>
</tr>
<tr>
<td>197</td>
<td>O + γ → O^{+} + e^{-}</td>
<td>R_{197} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>32</td>
</tr>
<tr>
<td>198</td>
<td>CO + γ → C + O + e^{-}</td>
<td>R_{198} = 5.6 \times 10^{-11}</td>
<td>1.2</td>
<td>33</td>
</tr>
</tbody>
</table>
Table B2. List of photochemical reactions included in our chemical model

<table>
<thead>
<tr>
<th>No.</th>
<th>Reaction</th>
<th>Optically thin rate (s⁻¹)</th>
<th>γ</th>
<th>Ref.</th>
</tr>
</thead>
<tbody>
<tr>
<td>166</td>
<td>H⁻ + γ → H + e⁻</td>
<td>R₁₆₆ = 7.1 × 10⁻⁷</td>
<td>0.5</td>
<td>1</td>
</tr>
<tr>
<td>167</td>
<td>H⁺ + γ → H + H⁺</td>
<td>R₁₆₇ = 1.1 × 10⁻⁹</td>
<td>1.9</td>
<td>2</td>
</tr>
<tr>
<td>168</td>
<td>H₂ + γ → H + H</td>
<td>R₁₆₈ = 5.6 × 10⁻¹¹</td>
<td>See §2.2</td>
<td>3</td>
</tr>
<tr>
<td>169</td>
<td>H + γ → H⁺</td>
<td>R₁₆₉ = 4.9 × 10⁻¹³</td>
<td>1.8</td>
<td>4</td>
</tr>
<tr>
<td>170</td>
<td>H₂⁺ + γ → H⁻ + H⁺</td>
<td>R₁₇₀ = 4.9 × 10⁻¹³</td>
<td>2.3</td>
<td>4</td>
</tr>
</tbody>
</table>

Table B3. List of reactions included in our chemical model that involve cosmic rays or cosmic-ray induced UV emission

<table>
<thead>
<tr>
<th>No.</th>
<th>Reaction</th>
<th>Rate (s⁻¹ cm⁻³)</th>
<th>Ref.</th>
</tr>
</thead>
<tbody>
<tr>
<td>199</td>
<td>H + c.r. → H⁺ + e⁻</td>
<td>R₁₉₉ = 1.0</td>
<td>—</td>
</tr>
<tr>
<td>200</td>
<td>He + c.r. → He⁺ + e⁻</td>
<td>R₂₀₀ = 1.1</td>
<td>1</td>
</tr>
<tr>
<td>201</td>
<td>H₂ + c.r. → H⁺ + H + e⁻</td>
<td>R₂₀₁ = 0.037</td>
<td>1</td>
</tr>
<tr>
<td>202</td>
<td>H₂ + c.r. → H + H</td>
<td>R₂₀₂ = 0.22</td>
<td>1</td>
</tr>
<tr>
<td>203</td>
<td>H₂ + c.r. → H⁺ + H⁻</td>
<td>R₂₀₃ = 6.5 × 10⁻⁴</td>
<td>1</td>
</tr>
<tr>
<td>204</td>
<td>H₂⁺ + c.r. → H⁻ + e⁻</td>
<td>R₂₀₄ = 2.0</td>
<td>1</td>
</tr>
<tr>
<td>205</td>
<td>C + c.r. → C⁺ + e⁻</td>
<td>R₂₀₅ = 3.8</td>
<td>1</td>
</tr>
<tr>
<td>206</td>
<td>O + c.r. → O⁺ + e⁻</td>
<td>R₂₀₆ = 5.7</td>
<td>1</td>
</tr>
<tr>
<td>207</td>
<td>CO + c.r. → CO⁺ + e⁻</td>
<td>R₂₀₇ = 6.5</td>
<td>1</td>
</tr>
<tr>
<td>208</td>
<td>C + γ.c.r. → C⁺ + e⁻</td>
<td>R₂₀₈ = 2800</td>
<td>2</td>
</tr>
<tr>
<td>209</td>
<td>CH + γ.c.r. → C⁺ + H</td>
<td>R₂₀₉ = 4000</td>
<td>3</td>
</tr>
<tr>
<td>210</td>
<td>CH + γ.c.r. → C⁺ + H</td>
<td>R₂₁₀ = 960</td>
<td>3</td>
</tr>
<tr>
<td>211</td>
<td>CH₂ + γ.c.r. → CH⁺ + e⁻</td>
<td>R₂₁₁ = 2700</td>
<td>1</td>
</tr>
<tr>
<td>212</td>
<td>CH₂ + γ.c.r. → CH⁺ + e⁻</td>
<td>R₂₁₂ = 2700</td>
<td>1</td>
</tr>
<tr>
<td>213</td>
<td>C₂ + γ.c.r. → C⁺ + C</td>
<td>R₂₁₃ = 1300</td>
<td>3</td>
</tr>
<tr>
<td>214</td>
<td>OH + γ.c.r. → O + H</td>
<td>R₂₁₄ = 2800</td>
<td>3</td>
</tr>
<tr>
<td>215</td>
<td>H₂O + γ.c.r. → H²O⁺</td>
<td>R₂₁₅ = 5300</td>
<td>3</td>
</tr>
<tr>
<td>216</td>
<td>O₂ + γ.c.r. → O₂⁺ + e⁻</td>
<td>R₂₁₆ = 4100</td>
<td>3</td>
</tr>
<tr>
<td>217</td>
<td>O₂ + γ.c.r. → O₂⁺ + e⁻</td>
<td>R₂₁₇ = 640</td>
<td>3</td>
</tr>
<tr>
<td>218</td>
<td>CO + γ.c.r. → C + O</td>
<td>R₂₁₈ = 0.218T^(1/2)</td>
<td>4</td>
</tr>
</tbody>
</table>

(Glover, Federrath, Mac Low, Klessen, 2010, MNRS, 404, 2)
HI to H2 conversion rate

Figure 4. Time evolution of the mass-weighted H$_2$ abundance in simulations R1, R2 and R3, which have numerical resolutions of 64$^3$ zones (dot-dashed), 128$^3$ zones (dashed) and 256$^3$ zones (solid), respectively.

(Glover, Federrath, Mac Low, Klessen, 2010)
HI to H2 conversion rate

H2 forms rapidly in shocks / transient density fluctuations / H2 gets destroyed slowly in low density regions / result: turbulence greatly enhances H2-formation rate

Figure 4. Time evolution of the mass-weighted H$_2$ abundance in simulations R1, R2 and R3, which have numerical resolutions of 64$^3$ zones (dot-dashed), 128$^3$ zones (dashed) and 256$^3$ zones (solid), respectively.

(Glover, Federrath, Mac Low, Klessen, 2010)
HI to H2 conversion rate

compare to data from Tamburro et al. (2008) study:

t_{\text{form}} \sim \text{few } \times 10^6 \text{ years}

Figure 4. Time evolution of the mass-weighted H$_2$ abundance in simulations R1, R2 and R3, which have numerical resolutions of $64^3$ zones (dot-dashed), $128^3$ zones (dashed) and $256^3$ zones (solid), respectively.
CO, C\(^+\) formation rates

**Figure 5.** Time evolution of the mass-weighted abundances of atomic carbon (black lines), CO (red lines), and C\(^+\) (blue lines) in simulations with numerical resolutions of 64\(^3\) zones (dot-dashed), 128\(^3\) zones (dashed) and 256\(^3\) zones (solid).

(Glover, Federrath, Mac Low, Klessen, 2010)
effects of chemistry 1

(Glover, Federrath, Mac Low, Klessen, 2010)
effects of chemistry 2

(Glover, Federrath, Mac Low, Klessen, 2010)
x-factor

• conversion rate between H$_2$ column density and CO emission (equivalent width W)

$$X = \frac{N_{H_2}}{W} \text{ (cm}^{-2} \text{ K}^{-1} \text{ km}^{-1} \text{ s)}$$

• most mass H$_2$ determinations depend on X!

• in Milky Way $X \sim$ few $\times$ $10^{22}$ cm$^{-2}$ K$^{-1}$ km$^{-1}$ s $\sim$ const.

• why is it constant?

• how does it vary with environmental condition?
  - metallicity
  - density, radiation field, etc.
  (“normal” gal. vs star burst)
Figure 4. Images of (a) $N_{\text{CO}}$, (b) $W$, (c) $N_{\text{H}_2}$, and (d) the $X$ factor of model n300-Z03. Each side has a length of 20 pc. In (a) and (b), solid contours indicate $\log(N_{\text{CO}}) = 12, 14$ and $\log(W) = -3, -1$; dashed contours are $\log(N_{\text{CO}}) = 16.5$ and $\log(W) = 1.5$ (see the text and Fig. 2d).

By $\approx 1$ order of magnitude. Since the $X$ factor directly depends on $W$ and only indirectly on $N_{\text{CO}}$, the $X$ factor also only falls into a limited range. Positions with the largest $X$ factors correspond to the lowest $N_{\text{H}_2}$ regions as well as low $N_{\text{CO}}$ and $W$ regions. These are the regions where CO is most affected by photodissociation. Since the amount of photodissociation depends on the 'effective' column density in each location of the 3D simulation volume, regions with similar $N_{\text{H}_2}$ can have very different $N_{\text{CO}}$ values (see also Papers I and II), as evident in Fig. 5(b): at low-to-intermediate $H_2$ densities $10^{21} \lesssim N_{\text{H}_2} \lesssim 10^{22}$ cm$^{-2}$, there is a wide range of $N_{\text{CO}}$ for a given $N_{\text{H}_2}$.

Since the $X$ factor (indirectly) depends on $N_{\text{CO}}$, its densities the $X$ factor also takes on a wide range. For instance, at $N_{\text{H}_2} = 5 \times 10^{21}$ cm$^{-2}$, the $X$ factor varies from $\sim 10^{20}$ to $10^{23}$ cm$^{-2} K^{-1}$ km$^{-1}$ s. Evidently, the $X$ factor can have a wide distribution within an MC, even for regions with identical molecular column densities. This is a consequence of the combination of a large distribution of $N_{\text{CO}}$ for a given $N_{\text{H}_2}$, as well as the lack of a simple correlation between $W$ and $N_{\text{CO}}$ due to the optically thick nature of CO. Fig. 5(d) shows that there can be a wide range in the $X$ factor even in very low-density regions. For this model ($n_0 = 100$ cm$^{-2}$ and $Z = Z_{\odot}$), much of the gas has $N_{\text{H}_2} \lesssim 10^{21}$ cm$^{-2}$. Unlike model n300 (in Fig. 5a), there is a very wide distribution in the $X$ factor in the range $10^{20} \lesssim N_{\text{H}_2} \lesssim 10^{21}$ cm$^{-2}$. This model also differs from model n300-Z03 (in Fig. 5b), showing a much larger distribution in $N_{\text{CO}}$ and the $X$ factor for a given $N_{\text{H}_2}$ at $N_{\text{H}_2} \lesssim 2 \times 10^{21}$ cm$^{-2}$. The $C$⃝2011 The Authors Monthly Notices of the Royal Astronomical Society C⃝2011 RAS (Shetty, Glover, Dullemond, Klessen 2011).
The line shows the only at the highest densities (log(Saturation)) become clearly evident. On the other hand, model n1000 shows a decreasing simply the H factor distributions, as evident from the similarity of the multiple peaks in the distribution in the approximately constant (with a mean value of 2^0.5) from this model. This line is a very good fit at log(Saturation) (shown in Fig. 6c) shows an increase in the factor distributions can be understood from the relationship between the CO and the H factor. Inset figures show the colour scale and previous picture.

Numerous simulations have shown that the (3D) volume density distribution, may be signatures of the internal dynamics of a cloud.

4.1 The column density distributions: can CO observations be more effective. As Fig. 5(c) shows for model n1000, the models with lower metallicity and density (Figs 6b and d) can have rather different distributions than (Fig. 6), however, such a trend is also present at the X factor. We have seen the X factor does not neatly follow the X factor. Inset figures show the colour scale and previous picture.

The correlation between the X factor and the X factor for four models.

The models with lower metallicity and density (Figs 6b and d) can have rather different distributions than (Fig. 6), however, such a trend is also present at the X factor. We have seen the X factor does not neatly follow the X factor. Inset figures show the colour scale and previous picture.

Given that the X factor can have a large distribution and can be well correlated with the X factor. Given that the X factor can have a large distribution and can be well correlated with the X factor. Given that the X factor can have a large distribution and can be well correlated with the X factor. Given that the X factor can have a large distribution and can be well correlated with the X factor. Given that the X factor can have a large distribution and can be well correlated with the X factor.

(Shetty, Glover, Dullemond, Klessen 2011)
Figure 5. X factor for four models. $N_{\text{CO}}$ is plotted as a function of $N_{\text{H}_2}$. The colour of each point indicates the X factor. Inset figures show the colour scale and PDF of the X factor. The corresponding maps of $N_{\text{H}_2}$, $N_{\text{CO}}$ and the X factor from model n300-Z03 are shown in Fig. 4.
observed x-factor

Tacconi et al. (2008)
Figure 1. Compilation of estimated $X$ factors from a range of systems, shown as a function of surface density. Figure reproduced from Tacconi et al. (2008).

Figure 2. Mean $X$ factor in bins of gas surface density $\Sigma_{\text{gas}}$ for 5 models. The $X$ factor is averaged in different $\Sigma_{\text{gas}}$ bins. The value of $X$ is plotted on the midpoint value of $\Sigma_{\text{gas}}$ of each bin. Each model is identified by different colors and symbols (and labeled in the legend). The large symbols shows the global (emission weighted) mean $X$ factor and mean $\Sigma_{\text{gas}}$ from each model. This indicates that the $X$ factor is dependent on three quantities: the column density of $H_2$, the peak CO intensity, and the range in velocities. Due to the coupling between hydrodynamics, thermodynamics, and chemistry, $T_B$ is also dependent on the velocity and density (as well as the kinetic temperature). We aim to understand the relative contribution of each of these three properties of the MC. After assessing the $X$ factor from the original Milky Way MC model, we alter one of these properties, while keeping the others fixed, and recompute the $X$ factor. In this manner, we can identify the most important cloud properties responsible for setting the $X$ factor.

3 MODELING METHOD

3.1 Numerical Magnetohydrodynamics, Chemistry, and Radiative Transfer

To investigate how MC characteristics affect the $X$ factor, we analyze magnetohydrodynamic (MHD) models of molecular clouds that include a treatment of chemistry. We perform radiative transfer calculations on these numerical models, in order to solve for the CO level populations and compute the emergent CO intensity. The ratio of the $H_2$ column density to the emergent CO intensity then gives the $X$ factor (Eqn. [1]).

The MHD grid-based models follow the evolution of an initially fully atomic medium with constant density in a $(20 \text{ pc})^3$ periodic box. Thermodynamics is coupled with...
• extend range of model parameters
  - we are currently running starburst galaxies with higher density, and 1000x increased radiation field and/or 1000x increased cosmic ray intensity

next steps

Genzel et al. (2010)
are molecules needed for star formation?

- it has been proposed that molecule formation (H₂, CO, etc.) is a prerequisite for star formation (e.g. Schaye 2004; Krumholz & McKee 2005; Elmegreen 2007; Krumholz et al. 2009)

- the idea is that CO is a necessary coolant for collapse

- however, also C+ is a very efficient coolant! (Glover & Clark 2011)

- to address this question, we performed dedicated simulations in Heidelberg
are molecules needed to form stars?

NO! CII, CI, provide equal amounts of cooling to CO . . .

image from Simon Glover
are molecules needed for star formation?

- presence of molecular gas has only very minor influence on ability of cloud to form stars
- $C^+$ is equally efficient coolant in atomic phase as CO in molecular
- what is crucial is the ability of cloud to shield itself from interstellar radiation field
- but clouds that are big/dense enough to shield themselves will be molecular!

this suggests that the correlation between $H_2$ and star formation is a coincidence

Glover & Clark (2011)
where is most energy lost?

images from Simon Glover
where is most energy lost?

images from Simon Glover
where is most energy lost?

K km/s

images from Simon Glover
where is most energy lost?

\[ C^+ \]

images from Simon Glover
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 x $10^{10}$ cm
- 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 \textit{GRAVITY}

• opposing forces
  - there are several processes that can oppose gravity
    - \textit{GAS PRESSURE}
    - \textit{TURBULENCE}
    - \textit{MAGNETIC FIELDS}
    - \textit{RADIATION PRESSURE}

Modern star formation theory is based on the complex interplay between all these processes.
**early theoretical models**

- **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

\[ l_{\text{turb}} \ll l_{\text{dyn}} \]

- then turbulent velocity dispersion contributes to effective soundspeed:

\[ c_c^2 \rightarrow c_c^2 + \sigma_{\text{rms}}^2 \]

- \( \rightarrow \) Larger effective Jeans masses \( \rightarrow \) more stability

- BUT: (1) turbulence depends on \( k \):

\[ \sigma_{\text{rms}}^2(k) \]

(2) supersonic turbulence \( \rightarrow \) \( \sigma_{\text{rms}}^2(k) >> c_s^2 \) usually
problems of early dynamical theory

• molecular clouds are *highly Jeans-unstable*, yet, they do *NOT* form stars at high rate and with high efficiency (Zuckerman & Evans 1974 conundrum) (the observed global SFE in molecular clouds is \( \sim 5\% \))

\[ \Rightarrow \text{something prevents large-scale collapse.} \]

• all throughout the early 1990’s, molecular clouds had been thought to be long-lived quasi-equilibrium entities.

• molecular clouds are *magnetized*
Magnetic star formation

- Mestel & Spitzer (1956): Magnetic fields can prevent collapse!!!
  - Critical mass for gravitational collapse in presence of B-field
    \[ M_{cr} = \frac{5^{3/2} B^3}{48\pi^2 G^{3/2} \rho^2} \]
  - 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
“standard theory” of star formation

- BASIC ASSUMPTION: Stars form from magnetically highly subcritical cores

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

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

- Was (in principle) only intended for isolated, low-mass stars

Frank Shu, 1943 -
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 ($\tau_{ff} << \tau_{AD}$) (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)
Observed B-fields are weak

$B$ versus $N(H_2)$ from Zeeman measurements. (from Bourke et al. 2001)

→ cloud cores are magnetically supercritical!!!

$(\Phi/M)_n > 1$ no collapse

$(\Phi/M)_n < 1$ collapse
Fig. 1.— The Arecibo telescope primary beam (small circle centered at 0,0) and the four GBT telescope primary beams (large circles centered 6′ north, south, east, and west of 0,0. The dotted circles show the first sidelobe of the Arecibo telescope beam. All circles are at the half-power points.

Crutcher et al. (2009)
Field reversal in the outer parts. This is incompatible with “standard” ambipolar diffusion theory!

Crutcher et al. (2009)

Fig. 2.— OH 1667 MHz spectra toward the core of L1448CO obtained with the Arecibo telescope (center panel) and toward each of the envelope positions 6′ north, south, east, and west of the core, obtained with the GBT. In the upper left of each panel is the inferred $B_{LOS}$ and its 1σ uncertainty at that position. A negative $B_{LOS}$ means the magnetic field points toward the observer, and vice versa for a positive $B_{LOS}$. 
Table 2. Relative Mass/Flux

<table>
<thead>
<tr>
<th>Cloud</th>
<th>$\mathcal{R}$</th>
<th>$\mathcal{R}'$</th>
<th>Probability $\mathcal{R}$ or $\mathcal{R}' &gt; 1$</th>
</tr>
</thead>
<tbody>
<tr>
<td>L1448CO</td>
<td>$0.02 \pm 0.36$</td>
<td>$0.07 \pm 0.34$</td>
<td>0.005</td>
</tr>
<tr>
<td>B217-2</td>
<td>$0.15 \pm 0.43$</td>
<td>$0.19 \pm 0.41$</td>
<td>0.05</td>
</tr>
<tr>
<td>L1544</td>
<td>$0.42 \pm 0.46$</td>
<td>$0.46 \pm 0.43$</td>
<td>0.11</td>
</tr>
<tr>
<td>B1</td>
<td>$0.41 \pm 0.20$</td>
<td>$0.44 \pm 0.19$</td>
<td>0.010</td>
</tr>
</tbody>
</table>

Fig. 2.— OH 1667 MHz spectra toward the telescope (center panel) and toward each of the positions west of the core, obtained with the GBT. In this figure, each spectrum is shown with its 1σ uncertainty at that position. A negative velocity toward the observer, and vice versa for a positive velocity.
Fig. 1.—Left: Simulated $^{13}$CO (1–0) map of the model in the z-axis direction. The locations of the cloud cores are shown with squares. The circles indicate the locations of telescope beams used in the synthetic observations of three cores. Right: Line-of-sight magnetic field strength as calculated from Zeeman splitting.

Bertram et al. (2012)
Figure 4. Distribution of clumps in different LoS directions for (i) PPP and (ii) PP measurements and observed cores by Crutcher et al. (2009). From the top to bottom: different values of $\beta$ ($\beta = 0.01, 0.1, 1, 10$ and $100$). From the left-hand to right-hand side: different time-steps ($t = 2.0, 2.4$ and $2.8T$). The initial magnetic field strength for $\beta$ is marked with a vertical line. Plotted is the absolute value of $R$ against the absolute value of the average of the magnetic field components for a given LoS. In general, we observe a small value of $|R|$ for small magnetic field strengths that might be caused by field reversals. The stronger the magnetic field lines, the higher the value of $|R|$. For PPP and PP configurations, as well as for the three different times, we get statistically the same distribution.
Molecular cloud dynamics

• **Timescale problem:** Turbulence *decays* on timescales *comparable to the free-fall time* $\tau_{ff}$ ($E \propto t^{-\eta}$ with $\eta \approx 1$).

  - Timescale problem

  - Magnetic fields (static or wave-like) *cannot* prevent loss of energy.

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 ... -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?)
molecular clouds
\[ \sigma_{\text{rms}} \approx \text{several km/s} \]
\[ M_{\text{rms}} > 10 \]
\[ L > 10 \text{ pc} \]

能量来源及尺度
NOT known
(supernovae, winds, spiral density waves?)

massive cloud cores
\[ \sigma_{\text{rms}} \approx \text{few km/s} \]
\[ M_{\text{rms}} \approx 5 \]
\[ L \approx 1 \text{ pc} \]

dense protostellar cores
\[ \sigma_{\text{rms}} << 1 \text{ km/s} \]
\[ M_{\text{rms}} \leq 1 \]
\[ L \approx 0.1 \text{ pc} \]

dissipation scale not known
(ambipolar diffusion, molecular diffusion?)

turbulent cascade in the ISM
dynamical SF in a nutshell

- interstellar gas is highly *inhomogeneous*
  - gravitational instability
  - thermal instability
  - *turbulent compression* (in shocks $\delta \rho/\rho \propto M^2$; in atomic gas: $M \approx 1...3$)

- cold *molecular clouds* can form rapidly in high-density regions at *stagnation points* of convergent *large-scale flows*
  - chemical *phase transition*: atomic $\rightarrow$ molecular
  - process is *modulated* by large-scale *dynamics* in the galaxy

- inside *cold clouds*: turbulence is highly supersonic ($M \approx 1...20$)
  $\rightarrow$ *turbulence* creates large density contrast,
  *gravity* selects for collapse

---

[GRAVOTUBULENT FRAGMENTATION]

- *turbulent cascade*: local compression *within* a cloud provokes collapse $\rightarrow$
  formation of individual *stars* and *star clusters*

(e.g. Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125-194)
Density structure of MC’s

Molecular clouds are highly inhomogeneous.

Stars form in the densest and coldest parts of the cloud.

$\rho$-Ophiuchus cloud seen in dust emission.

Let’s focus on a cloud core like this one.

(Motte, André, & Neri 1998)
Evolution of cloud cores

- How does this core evolve? Does it form one single massive star or cluster with mass distribution?

- Turbulent cascade „goes through“ cloud core
  --> NO scale separation possible
  --> NO effective sound speed

- Turbulence is supersonic!
  --> produces strong density contrasts: $\delta \rho / \rho \approx M^2$
  --> with typical $M \approx 10$ --> $\delta \rho / \rho \approx 100$!

- many of the shock-generated fluctuations are Jeans unstable and go into collapse

- --> expectation: core breaks up and forms a cluster of stars
Evolution of cloud cores

indeed $\rho$-Oph B1/2 contains several cores ("starless" cores are denoted by $\star$, cores with embedded protostars by $\star\star$)

(Motte, André, & Neri 1998)
Formation and evolution of cores

- protostellar cloud cores form at **stagnation point in convergent turbulent flows**

  - if $M > M_{\text{crit}} \propto \rho^{-1/2} T^{3/2}$: collapse & star formation
  - pf $M < M_{\text{crit}} \propto \rho^{-1/2} T^{3/2}$: reexpansion after end of external compression

  (e.g. Vazquez-Semadeni et al 2005)

- typical timescale: $t \approx 10^4 \ldots 10^5$ yr
What happens to distribution of cloud cores?

Two extreme cases:

1. Turbulence dominates energy budget:
   \[ \alpha = \frac{E_{\text{kin}}}{|E_{\text{pot}}|} > 1 \]
   --> individual cores do not interact
   --> collapse of individual cores dominates stellar mass growth
   --> loose cluster of low-mass stars

2. Turbulence decays, i.e. gravity dominates:
   \[ \alpha = \frac{E_{\text{kin}}}{|E_{\text{pot}}|} < 1 \]
   --> global contraction
   --> core do interact while collapsing
   --> competition influences mass growth
   --> dense cluster with high-mass stars
turbulence creates a hierarchy of clumps
as turbulence decays locally, contraction sets in
as turbulence decays locally, contraction sets in
while region contracts, individual clumps collapse to form stars
while region contracts, individual clumps collapse to form stars
individual clumps collapse to form stars
individual clumps collapse to form stars
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
in dense clusters, clumps may merge while collapsing
--> then contain multiple protostars
in *dense clusters*, competitive mass growth becomes important
in dense clusters, competitive mass growth becomes important
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
some concerns of simple model

• energy balance
  - in molecular clouds:
    kinetic energy ~ potential energy ~ magnetic energy > thermal energy
  - models based on HD turbulence misses important physics
  - in certain environments (Galactic Center, star bursts), energy density in cosmic rays and radiation is important as well

• time scales
  - star clusters form fast, but more slowly than predicted by HD only (feedback and magnetic fields do help)
  - initial conditions do matter (turbulence does not erase memory of past dynamics)

• star formation efficiency (SFE)
  - SFE in gravoturbulent models is too high (again more physics needed)
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

- 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.
Progress requires a comprehensive theoretical approach.

Star formation is intrinsically a multi-scale and multi-physics problem, where it is difficult to single out individual processes. The color results from assigning different hues (colors) to each monochromatic image. In this case, the assigned colors are:

- CTIO: ([O III] 501nm) blue
- CTIO: (H-alpha+[N II] 658nm) green
- CTIO: ([S II] 672+673nm) red
- HST/ACS: F656N (H-alpha+[N II]) luminosity*
Star formation is intrinsically a multi-scale and multi-physics problem, where it is difficult to single out individual processes. Progress requires a comprehensive theoretical approach.
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?
initial mass function
stellar mass function

stars seem to follow a universal mass function at birth \(\rightarrow\) IMF

(Kroupa 2002)

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

**BUT:** maybe variations with galaxy type (bottom heavy in the centers of large ellipticals)

from JAM (Jeans anisotropic multi Gaussian expansion) modeling

inferred excess of low-mass stars compared to Kroupa IMF

distribution of stellar masses depends on

- turbulent initial conditions
  --> mass spectrum of prestellar cloud cores

- collapse and interaction of prestellar cores
  --> competitive 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

(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)
The IMF distribution of stellar masses depends on:

- **turbulent initial conditions**
  --> mass spectrum of prestellar cloud cores ???

- Collapse and interaction of prestellar cores
  --> competitive accretion and $N$-body effects

- Thermodynamic properties of gas
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(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)
there are different quantitative IMF based on turbulence

- Hopkins (2012)

all relate the mass spectrum to statistical characteristics of the turbulent velocity fields

---

**ANALYTICAL THEORY FOR THE INITIAL MASS FUNCTION: CO CLUMPS AND PRESTELLAR CORES**

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*Received 2008 February 12; accepted 2008 May 4*
there are different quantitative IMF based on turbulence

- Hopkins (2012)

all relate the mass spectrum to statistical characteristics of the turbulent velocity fields
different statistical approaches

- there are different quantitative IMF based on turbulence
  - Hopkins (2012)
- all relate the mass spectrum to statistical characteristics of the turbulent velocity fields
there are different quantitative IMF based on turbulence

- Hopkins (2012)
- all relate the mass spectrum to statistical characteristics of the turbulent velocity fields

there are alternative approaches

- IMF from (proto)stellar feedback (Silk 1995, Adams & Fatuzzo 1996)
- IMF from competitive coagulation (Murray & Lin 1995, Bonnell et al. 2001ab, etc.)
caveat: everybody gets the IMF!

- combine scale free process $\rightarrow$ **POWER LAW BEHAVIOR**
  - gravity in dense clusters (Bonnell & Bate 2006, Klessen 2001)
  - universality: dust-induced EOS kink insensitive to radiation field (Elmegreen et al. 2008)

- with highly stochastic processes $\rightarrow$ central limit theorem $\rightarrow$ **GAUSSIAN DISTRIBUTION**
  - basically mean thermal Jeans length (or feedback)
  - universality: insensitive to metallicity (Clark et al. 2009, submitted)
caveat: everybody gets the IMF!

“everyone” gets the right IMF
→ better look for secondary indicators

- stellar multiplicity
- protostellar spin (including disk)
- spatial distribution + kinematics in young clusters
- magnetic field strength and orientation
caveat: dilatational vs. solenoidal

- density pdf depends on “dimensionality” of driving
  - relation between width of pdf and Mach number
  \[
  \frac{\sigma_\rho}{\rho_0} = bM
  \]
  - with \( b \) depending on \( \zeta \) via
  \[
  b = 1 + \left[ \frac{1}{D} - 1 \right] \zeta = \begin{cases} 
1 - \frac{2}{3} \zeta, & \text{for } D = 3 \\
1 - \frac{1}{2} \zeta, & \text{for } D = 2 \\
1, & \text{for } D = 1
\end{cases}
\]
  - with \( \zeta \) being the ratio of dilatational vs. solenoidal modes:
  \[
  P_{ij}^\zeta = \zeta P_{ij}^\perp + (1 - \zeta) P_{ij}^\parallel = \zeta \delta_{ij} + (1 - 2 \zeta) \frac{k_i k_j}{|k|^2}
  \]

Fig. 3. — Volume-weighted density PDFs \( p(s) \) obtained from 3D, 2D and 1D simulations with compressive forcing and from 3D and 2D simulations using solenoidal forcing. Note that in 1D, only compressive forcing is possible as in the study by Passot & Vázquez-Semadeni (1998). As suggested by eq. (5), compressive forcing yields almost identical density PDFs in 1D, 2D and 3D with \( b \sim 1 \), whereas solenoidal forcing leads to a density PDF with \( b \sim 1/2 \) in 2D and with \( b \sim 1/3 \) in 3D.
caveat: dilatational vs. solenoidal

- density pdf depends on “dimensionality” of driving
  → is that a problem for the Krumholz & McKee model of the SF efficiency?

- density pdf of compressive driving is NOT log-normal
  → is that a problem for the Padoan & Nordlund, or Hennebelle & Chabrier IMF model?

- most “physical” sources should be compressive (convergent flows from spiral shocks or SN)

Federrath, Klessen, Schmidt (2008b)
caveat: dilatational vs. solenoidal

- density power spectrum differs between dilatational and solenoidal driving!

→ dilatational driving leads to break at sonic scale!

- can we use that to determine driving sources from observations?

compensated density spectrum $kS(k)$ shows clear break at sonic scale. Below that shock compression no longer is important in shaping the power spectrum...

Federrath, Klessen, Schmidt (2008b)
distribution of stellar masses depends on
  
  - turbulent initial conditions
    --> mass spectrum of prestellar cloud cores
  
  - collapse and interaction of prestellar cores
    --> competitive mass growth 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

(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)
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 ~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

Bonnell & Clark 2008
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.

initial conditions for cluster formation
ICs of star cluster formation

• key question:
  - what is the initial density profile of cluster forming cores? how does it compare low-mass cores?

• observers answer:
  - very difficult to determine!
    - most high-mass cores have some SF inside
    - infra-red dark clouds (IRDCs) are difficult to study
  - but, new results with Herschel

IRDC near Aquila rift, studied with the SMA: J. Swift & E. Churchwell
ICs of star cluster formation

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  - how does it compare low-mass cores?

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IRDC observed with Herschel, Peretto et al. (2010)
ICs of star cluster formation

• key question:
  - what is the initial density profile of cluster forming cores? how does it compare low-mass cores?

• theorists answer:
  - top hat (Larson Penston)
  - Bonnor Ebert (like low-mass cores)
  - power law $\rho \propto r^{-1}$ (logotrop)
  - power law $\rho \propto r^{-3/2}$ (Krumholz, McKee, et
  - power law $\rho \propto r^{-2}$ (Shu)
  - and many more
different density profiles

• does the density profile matter?

• in comparison to
  - turbulence ...
  - radiative feedback ...
  - magnetic fields ...
  - thermodynamics ...
different density profiles

- address question in simple numerical experiment
- perform extensive parameter study
  - different profiles (top hat, BE, $r^{-3/2}$, $r^{-3}$)
  - different turbulence fields
    - different realizations
    - different Mach numbers
    - solenoidal turbulence
dilatational turbulence
both modes
- no net rotation, no B-fields
(at the moment)
Girichids et al. (2011abc)
for the $r^{-2}$ profile you need to crank up turbulence a lot to get some fragmentation!
ICs with flat inner density profile form more fragments

Girichids et al. (2011abc)
however, the real situation is very complex: details of the initial turbulent field matter very high Mach numbers are needed to make SIS fragment

Girichids et al. (2011abc)
different density profiles

- different density profiles lead to very different fragmentation behavior
- fragmentation is strongly suppressed for very peaked, power-law profiles

- this is *good*, because it may explain some of the theoretical controversy, we have in the field
- this is *bad*, because all current calculations are “wrong” in the sense that the formation process of the star-forming core is neglected.

Girichids et al. (2011abc)
distribution of stellar masses depends on
  - turbulent initial conditions
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thermodynamic properties of gas
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(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)
dependency on EOS

• 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

dependency on EOS

\( \gamma = 0.2 \)

\( \gamma = 1.0 \)

\( \gamma = 1.2 \)

for \( \gamma < 1 \) fragmentation is enhanced \( \rightarrow \) cluster of low-mass stars

for \( \gamma > 1 \) it is suppressed \( \rightarrow \) formation of 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) \( M_{\text{jeans}} \propto \gamma^{3/2} \rho^{(3\gamma-4)/2} \)

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

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

In this section, we review thermal evolution of the cloud core in the metal-free case. We then describe the effects of metallicity on the core evolution. There are, however, small disagreements, (Omukai et al. 2005, 2010) which are not discussed here. The thermal evolution is nearly isothermal with temperature differences of order 10% until 10^4 M_\odot. Beyond this mass, the dynamical collapse is halted as the pressure becomes greater than the star-forming gas. Another critical value is exceeded when the Jeans mass is 10^2 M_\odot, which is important in the thermal evolution. The effective ratio of specific heat remains below 4, while fragmentation is strongly prohibited for Z = 0. Another critical value is ∼ 10^{-3} M_\odot, which is associated with the three-body reaction (Equation (9)). Until very high density (Figure 3), the heating is due to the three-body reaction H + H → H_2 + e, which dominates. For the cooling, the H_2 line emission contributes strongly, especially toward a higher density. The steep decline of the H_2 lines is due to the H_2 dissociation channel by the density increase. After this plateau, the H_2 cooling rate at 10^6 M_\odot is due to the H_2 recombination proceeding, the H_2 recombination heating channel is quenched. Another molecular species in the metal-free gas, HD, is known to play an important role in cooling and heating. Another color version of this figure is available in the online journal. (Nagakura & Omukai 2003).
EOS as function of metallicity

In Figure 1, we present the temperature evolution at the center of the core as a function of the number density. The constant Jeans masses are indicated by the dashed lines. For those above 10^0 number density, which is calculated by one-zone models. The dashed lines play a crucial role in the thermal evolution. The evolution of H\_\gamma (below 10^0), the effective ratio of specific heat is an important index to examine the contribution to the cooling and heating rates by individual processes are presented. In Figure 2 of O05, where similar plots for the overall evolution is quite similar to that calculated by the one-zone model until 10^0, justifying the one-zone treatment for the prestellar cores as a function of the number density. The contribution of the curve in Figure 1, while fragmentation is strongly prohibited for number density log (n\_H (cm^-3)) ∝ (Z=0, [M/H]=-6, -5, -4, -3, -2, -1, 0) presents the temperature evolution at the center of the core evolution. There are, however, small disagreements, but is above 1 in this period except for brief intervals around 10^0 cm^-3, but is above 1 in this period except for brief.

Let us summarize here the formation processes of H\_\gamma over the gravity, and a hydrostatic object is formed. This value, the dynamical collapse is halted as the pressure dominates below 4. If a metal-free gas is once ionized (Uehara & Inutsuka 2007), the evolution is nearly isothermal with temperature difference ∼ \(10^3\) cm^-3, and the amount of formed H\_\gamma abundance begins to increase again at line-channel: [M/H]=-6, -5, -4, -3, -2, -1, 0, as functions of the number density. The constant Jeans masses are indicated by the dashed lines. For those above 10^0 number density, which is calculated by one-zone models. The dashed lines play a crucial role in the thermal evolution. The evolution of H\_\gamma (below 10^0), the effective ratio of specific heat is an important index to examine the contribution to the cooling and heating rates by individual processes are presented. In Figure 2 of O05, where similar plots for the overall evolution is quite similar to that calculated by the one-zone model until 10^0, justifying the one-zone treatment for the prestellar cores as a function of the number density. The contribution of the curve in Figure 1, while fragmentation is strongly prohibited for number density log (n\_H (cm^-3)) ∝ (Z=0, [M/H]=-6, -5, -4, -3, -2, -1, 0) presents the temperature evolution at the center of the core evolution. There are, however, small disagreements, but is above 1 in this period except for brief intervals around 10^0 cm^-3, but is above 1 in this period except for brief.
present-day star formation

\[ \gamma = 0.7 \]

\[ \gamma = 1.1 \]

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

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

at SFE $\approx 50\%$

need appropriate EOS in order to get low mass IMF right

EOS as function of metallicity

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

Indicate the constant Jeans masses. For those above 10 concentrations is presented in Figure Z/Z

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

Play a crucial role in the thermal evolution. The evolution of H one-zone models are presented. In Figure compared with Figure 2 of O05, where similar plots for the

\[ \gamma \]

of the curve in Figure 1, while fragmentation is strongly prohibited for

\[ \sim \]

20

1000

100

10

1

\[ T(K) \]

number density log \( n_H \) cm\(^{-3} \)

\[ 10^6 M_{\odot} \]

\[ 10^4 M_{\odot} \]

\[ 10^2 M_{\odot} \]

\[ M_{\odot} \]

\[ 10^{-2} M_{\odot} \]

\[ 10^6 M_{\odot} \]

\[ 10^4 M_{\odot} \]

\[ 10^2 M_{\odot} \]

\[ M_{\odot} \]

\[ 10^{-2} M_{\odot} \]

\[ Z=0 \]

[M/H]=-6

-5

-4

-3

-2

-1

0

\[ \tau = 1 \]

\[ 10^{-4} M_{\odot} \]

\[ \text{line cooling} \]

\[ \text{dust cooling} \]

\[ (\text{Omukai et al. 2005, 2010}) \]
transition: Pop III to Pop II.5

two competing models:

- cooling due to atomic fine-structure lines ($Z > 10^{-3.5} \ Z_{\text{sun}}$)
- cooling due to coupling between gas and dust ($Z > 10^{-5...-6} \ Z_{\text{sun}}$)

- which one is explains origin of extremely metal-poor stars

NB: lines would only make very massive stars, with $M > \text{few } \times 10 \ M_{\text{sun}}$. 

(Okukai et al. 2005, 2010)
transition: Pop III to Pop II.5

SDSS J1029151+172927

- is first ultra metal-poor star with $Z \sim 10^{-4.5} Z_{\text{Sun}}$ for all metals seen (Fe, C, N, etc.)
  [see Caffau et al. 2011]
- this is in regime, where metal-lines cannot provide cooling
  [e.g. Schneider et al. 2011, 2012, Klessen et al. 2012]

  - new ESO large program to find more of these stars
    (120h x-shooter, 30h UVES)
    [PI E. Caffau]

(Caffau et al. 2011, 2012)

(Schneider et al. 2011, 2012, Klessen et al. 2012)

<table>
<thead>
<tr>
<th>Element</th>
<th>$\Delta$</th>
<th>$\Delta$</th>
<th>$\Delta$</th>
<th>$\Delta$</th>
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<th>$S_H$</th>
<th>$A(X)_{0}$</th>
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<td>N</td>
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<td>$-5.0$</td>
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<td>$-4.68 \pm 0.11$</td>
<td>$-4.52 \pm 0.11$</td>
<td>$-4.49 \pm 0.12$</td>
<td>5</td>
<td>0.1</td>
<td>7.54</td>
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<td>$-3.93$</td>
<td>$-3.96$</td>
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<td>$-4.83 \pm 0.16$</td>
<td>$-4.76 \pm 0.18$</td>
<td>$-4.84 \pm 0.16$</td>
<td>6</td>
<td>1.0</td>
<td>4.90</td>
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<td>$-5.02 \pm 0.10$</td>
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<td>$-4.89 \pm 0.10$</td>
<td>43</td>
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<td>7.52</td>
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<tr>
<td>Ni i</td>
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<td>$-4.90 \pm 0.11$</td>
<td>$-4.94$</td>
<td>$-5.09$</td>
<td>10</td>
<td>6.23</td>
<td></td>
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<tr>
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<td>$-5.25$</td>
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<td>$-5.25$</td>
<td>1</td>
<td>0.01</td>
<td>2.92</td>
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</tbody>
</table>
transition: Pop III to Pop II.5

approach problem with high-resolution hydrodynamic calculations of central parts of high-redshift halos

- SPH (40 million particles)
- time-dependent chemistry (with dust)
- sink particles to model star formation
- external dark-matter potential

(Omukai et al. 2005, 2010)
transition: Pop III to Pop II.5

approach problem with high-resolution hydrodynamic calculations of central parts of high-redshift halos

- SPH (40 million particles)
- time-dependent chemistry (with dust)
- sink particles to model star formation
- external dark-matter potential
- focus on relevant density regime (i.e. include dust dip and optically thick regime)

(Omukai et al. 2005, 2010)
Fragmentation of star-forming clouds at very low metallicities

We expect that dust cooling becomes important at high densities, preventing the gas temperature from getting higher than 1500 K. For instance, the metal-free case reaches temperatures as high as 3000 K, which is significantly higher than the CMB temperature. In contrast, the gas temperature in the 10^{-4} Z⊙ case is close to the CMB temperature in the low density region.

When we compare our results to the calculations of Omukai et al. (2010), we find good agreement with their 1D hydrodynamical models, although we expected some small differences due to the effects of the turbulence and rotation (see Dopcke et al., 2012, submitted to ApJ, arXiv:1203.6842). The subsequent evolution of the gas is close to isothermal. When cooling and heating balance, the evolution is close to adiabatic. For 10^{-6} Z⊙, dust cooling becomes important for gas temperatures to be seen for different dust opacity models.

The accretion rate varies from 0.02 to 0.17 M⊙ yr^{-1} for the number of sink particles (N) divided by the time to accrete 4.7 M⊙. At 10^{-8} Z⊙, the accretion rate is comparable to the other thermal processes, and necessary to determine the density-temperature diagrams. We did not take this thermal process into account during the calculations, but it is relevant to speculate if it is different accretion can also influence the expected accretion luminosity. We did not take this thermal process into account during the calculations, but it is relevant to speculate if it is different accretion can also influence the expected accretion luminosity.

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Hints for differences in mass spectrum

Disk fragmentation mode

Gravoturbulent fragmentation mode
The thermal evolution of prestellar cloud cores with metallicities is presented in Figure 1. For the metal-free case, the dynamical collapse is halted as the pressure exceeds a certain value, which is calculated by one-zone models. In this period, cooling and heating are always almost balanced, so the evolution is nearly isothermal with temperature differences of orders of magnitude. The effective ratio of specific heat, which is an important index to examine the variation of pressure in response to the density variation, is shown for those cases. Note that the contribution of a metal-free gas dominates. For the cooling, the H$_2$ molecule plays a crucial role in the thermal evolution. The evolution of H$_2$ is catalyzed by a small amount of remaining electrons. With their recombination proceeding, the H$_2$ cooling associated with the three-body reaction (Equation (9)) saturates at a temperature around 10$^3$ K, but is above 1 in this period except for brief intervals around 10$^4$ K, where it falls slightly. That metallicity has a strong effect on the dynamical collapse of self-gravitating clouds to thermal equilibrium between the H$_2$ formation (Equation (10)), the three-body reaction (Equation (11)), and the cooling, via the three-body H$_2$ formation: $2H + H^+ \rightarrow H_2 + e$.

Let us summarize here the formation processes of H$_2$. 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 model (Figure 2). In this section, we review thermal evolution of the cloud core in particular, at high densities and for low-metallicity cases. Below $10^{-3}$, [M/H] = -6, the dynamical collapse is halted as the pressure exceeds 10$^3$ M$_{\odot}$, which is calculated by one-zone models. In Figure 3, we present the temperature evolution at the center of the curve in Figure 1 for those cases. Note that the temperature evolution is almost identical to that calculated by the metal-free case with [M/H] = 0. We then describe the effects of metallicity on the formation processes of H$_2$.
• 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

(Omukai et al. 2005, 2010)
most current numerical simulations of Pop III star formation predict very massive objects (e.g. Abel et al. 2002, Yoshida et al. 2008, Bromm et al. 2009)

similar for theoretical models (e.g. Tan & McKee 2004)
	here are some first hints of fragmentation, however (Turk et al. 2009, Stacy et al. 2010)
detailed look at accretion disk around first star

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


detailed look at accretion disk around first star

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

what is the time evolution of accretion disk around first star to form?

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.
important disk parameters

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, $Q = \frac{c_s}{\kappa} \frac{\pi G \Sigma}{\nu}$, 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 $\nu$ 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 (Clark et al. 2011b, Science, 331, 1040).

Toomre $Q$:

$Q = \frac{c_s \kappa}{\pi G \Sigma}$

instability for $Q < 1$

(Clark et al. 2011b, Science, 331, 1040)
similar study with very different numerical method (AREPO)

one out of five halos

We see “flat” mass spectrum

(expected mass spectrum)

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)
primordial star formation

• just like in present-day SF, we expect
  - turbulence
  - thermodynamics
  - feedback
  - magnetic fields
to influence first star formation.

• masses of first stars still uncertain (surprises from new generation of high-resolution calculations that go beyond first collapse)

• disks unstable: first stars should be binaries or part of small clusters

• effects of feedback less important than in present-day SF
The distribution of stellar masses depends on:

1. **Turbulent initial conditions**: Mass spectrum of prestellar cloud cores
2. **Collapse and interaction of prestellar cores**: Competitive accretion and N-body effects
3. **Thermodynamic properties of gas**: Balance between heating and cooling, EOS (determines which cores go into collapse)
4. **(Proto) stellar feedback terminates star formation**: Ionizing radiation, bipolar outflows, winds, SN
We want to address the following questions:

- how do massive stars (and their associated clusters) form?
- what determines the upper stellar mass limit?
- what is the physics behind observed HII regions?

**Imagery:**
- IMF (Kroupa 2002)
- Rosetta nebula (NGC 2237)
**(proto)stellar feedback processes**
- radiation pressure on dust particles
- ionizing radiation
- stellar winds
- jets and outflows

**ionization**
- few numerical studies so far (e.g. Dale 2007, Gritschneder et al. 2009), detailed collapse calculations with ionizing and non-ionizing feedback still missing
- HII regions around massive stars are directly observable
  --> direct comparison between theory and observations
our (numerical) approach

- focus on collapse of individual high-mass cores...
  - massive core with 1,000 $M_\odot$
  - Bonnor-Ebert type density profile
    (flat inner core with 0.5 pc and rho $\sim r^{-3/2}$ further out)
  - initial $m=2$ perturbation, rotation with $\beta = 0.05$
  - sink particle with radius 600 AU and threshold density of $7 \times 10^{-16}$ g cm$^{-3}$
  - cell size 100 AU
our (numerical) approach

• method:
  - FLASH with ionizing and non-ionizing radiation using raytracing based on hybrid-characteristics
  - protostellar model from Hosokawa & Omukai
  - rate equation for ionization fraction
  - relevant heating and cooling processes
  - some models include magnetic fields
  - first 3D MHD calculations that consistently treat both ionizing and non-ionizing radiation in the context of high-mass star formation

Peters et al. (2010a,b,c)
- disk is gravitationally unstable and fragments
- we suppress secondary sink formation by “Jeans heating”
- H II region is shielded effectively by dense filaments
- ionization feedback does not cut off accretion!

Peters et al. (2010a,b,c)
all protostars accrete from common gas reservoir
accretion flow suppresses expansion of ionized bubble
cluster shows “fragmentation-induced starvation”
halting of accretion flow allows bubble to expand

Peters et al. (2010a,b,c)
mass load onto the disk exceeds inward transport
--> becomes gravitationally unstable (see also Kratter & Matzner 2006, Kratter et al. 2010)

fragments to form multiple stars --> explains why high-mass stars are seen in clusters

younger protostars form at larger radii

“burst” of star formation

mass load onto the disk exceeds inward transport --> becomes gravitationally unstable (see also Kratter & Matzner 2006, Kratter et al. 2010)

fragments to form multiple stars --> explains why high-mass stars are seen in clusters

- compare with control run without radiation feedback
- total accretion rate does not change with accretion heating
- expansion of ionized bubble causes turn-off
- no triggered star formation by expanding bubble

• magnetic fields lead to weaker fragmentation
• central star becomes more massive (magnetic breaking

Fragmentation-induced starvation in a complex cluster

![Diagram](image)

- **Gas density as function of radius at different times**
- **Mass flow towards the center as function of radius at different times**

Girichidis et al. (2011a,b)
Overview of collapse simulations.

<table>
<thead>
<tr>
<th>Name</th>
<th>Resolution</th>
<th>Radiative Feedback</th>
<th>Multiple Sinks</th>
<th>$M_{\text{sinks}}$ (M$_\odot$)</th>
<th>$N_{\text{sinks}}$</th>
<th>$M_{\text{max}}$ (M$_\odot$)</th>
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<td>98 AU</td>
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<td>yes</td>
<td>125.56</td>
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<td>151.43</td>
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<td>14.64</td>
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Peters et al. (2010a,b,c)
relation between maximum stellar mass and total stellar mass

Bonnell et al. (2004): competitive accretion

Peters et al. (2010): fragmentation-induced starvation
- thermal pressure drives bipolar outflow
- filaments can effectively shield ionizing radiation
- when thermal support gets lost, outflow gets quenched again
- no direct relation between mass of star and size of outflow

Peters et al. (2010a,b,c)
- bipolar outflow during accretion phase
- when accretion flow stops, ionized bubble can expand
- expansion is highly anisotropic
- bubbles around most massive stars merge
numerical data can be used to generate continuum maps
calculate free-free absorption coefficient for every cell
integrate radiative transfer equation (neglecting scattering)
convolve resulting image with beam width
VLA parameters:
  - distance 2.65 kpc
  - wavelength 2 cm
  - FWHM 0.14
to the accuracy of 10$^{-3}$ Jy

Wood & Churchwell 1989 classification of UC H II regions

Question: What is the origin of these morphologies?

UC H II lifetime problem: Too many UC H II regions observed!
- synthetic VLA observations at 2 cm of simulation data
- interaction of ionizing radiation with accretion flow creates high variability in time and shape
- flickering resolves the lifetime paradox!

Peters et al. (2010a,b,c)
Morphology of HII region depends on viewing angle

Peters et al. (2010a,b,c)
<table>
<thead>
<tr>
<th>Type</th>
<th>WC89</th>
<th>K94</th>
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<td>55</td>
<td>19</td>
<td>60 ± 5</td>
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<tr>
<td>Cometary</td>
<td>20</td>
<td>16</td>
<td>7</td>
<td>10 ± 5</td>
</tr>
<tr>
<td>Core-halo</td>
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<td>Shell-like</td>
<td>4</td>
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<td>3</td>
<td>5 ± 1</td>
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<tr>
<td>Irregular</td>
<td>17</td>
<td>19</td>
<td>57</td>
<td>21 ± 5</td>
</tr>
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</table>


- statistics over 25 simulation snapshots and 20 viewing angles
- statistics can be used to distinguish between different models
- single sink simulation does not reproduce lifetime problem
- correlation between accretion events and H II region changes
- time variations in size and flux have been observed
- changes of size and flux of $5\text{–}7\%\,\text{yr}^{-1}$ match observations

Some results

- ionization feedback cannot stop accretion
- ionization drives bipolar outflows
- HII regions show high variability in time and shape
- all classified morphologies can be observed in one run
- lifetime of HII regions determined by accretion timescale (and not by expansion time)
- rapid accretion through dense and unstable flows
- fragmentation limits further accretion of massive stars
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. Many different processes need to be considered simultaneously.

- stars form from the complex interplay of self-gravity and a large number of competing processes (such as turbulence, B-field, feedback, thermal pressure)
- thermodynamic properties of the gas (heating vs cooling) play a key role in the star formation process
- detailed studies require the consistent treatment of many different physical and chemical processes (theoretical and computational challenge)
- star formation is regulated by several feedback loops, which are still poorly understood
Protostars and Planets VI in Summer 2013
Thanks!